Cosmological Distance Measurements with ROTSE Supernovae IIP and Observational Systematics on DESI Emission Line Galaxy Clustering

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COSMOLOGICAL DISTANCE MEASUREMENTS WITH ROTSE SUPERNOVAE IIP
AND OBSERVATIONAL SYSTEMATICS ON DESI EMISSION LINE GALAXY
CLUSTERING

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COSMOLOGICAL DISTANCE MEASUREMENTS WITH ROTSE SUPERNOVAE IIP AND OBSERVATIONAL SYSTEMATICS ON DESI EMISSION LINE GALAXY CLUSTERING

A Dissertation Presented to the Graduate Faculty of the Dedman College
Southern Methodist University
in Partial Fulfillment of the Requirements for the degree of Doctor of Philosophy with a Major in Physics by Govinda Dhungana

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Both Supernovae (SNe) and Baryon Acoustic Oscillations (BAO) surveys emerged as complementary probes of the expansion history of the universe in the last few decades. SNe Ia cosmology has reached the systematic limits in the optical surveys. The most frequently occurring SNe Type IIP are emerging as equally rich distance probes for the next generation larger surveys. In this thesis, I highlight the astrophysical observables of these events in the context of ROTSE III SN survey and using the ROTSE SNe IIP sample, I present calibration in the framework of expanding photosphere method (EPM) to use them as cosmological distance indicators and present the measurement of Hubble expansion in the low $- z$ universe. The upcoming DESI experiment will obtain largest spectroscopic sample to explore the universe for the last 10 billion years through BAO precision measurements. The design sensitivity however requires full control of the systematics. I present an analysis of systematic effects of the spectroscopic performance on the cosmological signal using simulated spectra and galaxy sample for the DESI survey. By modeling the spectroscopic inefficiencies for the simulation samples of emission line galaxies, I estimate the systematic bias and uncertainty on the BAO scale measurement due to various observing conditions.
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This dissertation is dedicated to my beloved parents Nilmani Dhungana and Yashoda Dhungana, my dearest wife Shraddha and my lovely kids Shreyash and Sampada.
Chapter 1

INTRODUCTION

This chapter introduces the basic cosmological principles and highlights some of the key parameters and tools to measure them from astronomical survey data. The subsequent chapters will use these tools to address the objective of this thesis, which is to measure the cosmological distances using supernovae and to measure the observational effects on galaxy clustering. The end of this chapter highlights the structure of this thesis.

1.1. The cosmological framework

This section describes the framework of the Lambda Cold Dark Matter (ΛCDM) model, which is often described as the standard model of Big Bang cosmology. The ΛCDM model has been very successful in accounting for the observed properties of various cosmological data including the structures in the cosmic microwave background temperature and polarization spectra, the accelerated expansion of the universe, the large scale structures of the galaxies and so on. ΛCDM relies on the assumption that the general theory of relativity is the correct theory of gravity and also offers significant space for extensions.

1.1.1. General theory of relativity

The evolution of the universe after the Big Bang is frequently described by Einstein’s field equations, in the framework of the general theory of relativity,

\[ G_{\mu\nu} = 8\pi T_{\mu\nu} - \Lambda g_{\mu\nu} \quad (1.1) \]

where \( G_{\mu\nu} \) is the \( 4 \times 4 \) Einstein tensor describing the curved space time geometry, \( T_{\mu\nu} \) is the \( 4 \times 4 \) stress-energy tensor in the space-time manifold, \( g_{\mu\nu} \) is the metric of the space-time
and $\Lambda$ is considered the vacuum energy density or a cosmological constant. Note that I have used the geometric units $c=1$ and $G=1$ in Eq. 1.1. This will be followed throughout. The implication of Eq. 1.1 is that the matter tells the geometry how to curve and the geodesic equation of a particle in space-time implies that geometry tells matter how to move. The last term of Eq. 1.1 is the natural zero-point of the theory that Einstein introduced to exactly balance out any dynamics of the universe, so as to maintain his static perception of the universe.

With the observations of receding galaxies, Hubble, in 1929, gave the first experimental insight of a dynamic, expanding universe via the equation

$$V = H_0D$$

where $V$ is the radial velocity of a particular galaxy, $D$ is the distance to the galaxy and $H_0$ is the Hubble constant.

With this finding, Einstein had to admit the idea of an evolving universe and called the cosmological constant in his theory his “biggest blunder” of life. From the observed correlation of recession velocity with distance in Eq. 1.2, a fundamental basis of calibrating distances using the brightness of astronomical sources was established. This thesis will revisit more of these sources in depth in later chapters, but first, I will introduce the basic framework of the standard cosmological model.

1.1.2. The cosmological principle

The evolution of the field equations 1.1 governs the dynamics of the universe. On a very small scale, the universe consists of stars, galaxies and gravitationally bound clusters. The distribution of galaxies appears inhomogeneous at scales as large as tens of megaparsecs (Mpc). However, from a statistical viewpoint, the homogeneity of the universe becomes evident on large scales. A concrete example of the homogeneity is evident in the correlation function of galaxy clustering of large astronomical surveys (e.g., [4]). Likewise, statistical
isotropy has been established from observations such as temperature fluctuations in the cosmic microwave background (CMB) (e.g., COBE, WMAP, Planck satellites). Thus at large scales, in a statistical sense, the universe preserves both translational invariance (homogeneity) and rotational invariance (isotropy). To understand the motion of particles in a space-time manifold, it is useful to describe the metric of space-time.

1.1.3. The space-time metric

The most general expression for an interval (length-squared) in 4D space-time manifold is given by

$$ds^2 = g_{\mu\nu}dx^\mu dx^\nu$$

(1.3)

where $\mu$ and $\nu$ run over 1, 2, 3 and 4, and a sum is implied over repeated indices, in generic tensor algebra convention. The elements of the metric tensor $g_{\mu\nu}$ define the space-time. Various combinations of derivatives of the metric coefficients of $g_{\mu\nu}$ give the elements of $G_{\mu\nu}$.

The metric is commonly expressed by resolving the time and space components, considering the orthogonality of temporal axis with spatial axes.

$$ds^2 = dt^2 - a^2(t)\eta_{ij}dx^idx^j$$

(1.4)

The component $\eta_{ij}$ constitutes the usual $3 \times 3$ matrix elements of the spatial coordinate system. The dynamics of the universe is then governed by the scale factor $a(t)$. It is also the “radius” of the universe for the given time. In expanding the metric in Eq. 1.4, using isotropy and homogeneity of the universe, one can obtain the famous Friedmann-Lemaître-Robertson-Walker (FLRW or FRW) metric. This is expressed in spherical polar coordinates as,

$$ds^2 = dt^2 - a^2(t) \left( \frac{dr^2}{1-\kappa r^2} + r^2(d\theta^2 + \sin^2\theta d\phi^2) \right)$$

(1.5)

where the curvature parameter $\kappa$ determines the spatial curvature of the universe. With $\kappa$ being negative, positive, or zero, the universe is open, closed or flat respectively.
1.1.4. Stress energy tensor: $T_{\mu\nu}$

All the non-gravitational fields of the universe are embedded in $T_{\mu\nu}$ of Eq. 1.1. This is generally described in terms of an enclosed energy density $\rho$ and pressure $p$. The equations of motion can be obtained using the covariant derivative, which not only accounts for the change of energy but also the change of field components with position in the space-time, given as,

$$T_{\mu\nu} = 0$$

(1.6)

This is simply restating the conservation of energy. Note that like $T_{\mu\nu}$, $G^{\mu\nu}$ is also divergenceless. Considering the evolution of a proper volume of a comoving 3-sphere, $V \sim a^3(t)$, in a FRW universe, one can obtain the following equations on the dynamics of scale factor $a(t)$.

$$\left(\frac{\dot{a}}{a}\right)^2 = \frac{8\pi}{3} \rho - \frac{\kappa}{a^2} + \frac{\Lambda}{3}$$

(1.7)

$$\frac{\ddot{a}}{a} = -\frac{4\pi}{3} (\rho + 3p) + \frac{\Lambda}{3}$$

(1.8)

$$\dot{\rho} = -3 \frac{\dot{a}}{a} (\rho + p)$$

(1.9)

The evolution of $a(t)$ explains the global expansion or contraction of the universe. The equations above involving the first and second derivative of $a(t)$ are called the Friedmann equations and form the basis of the $\Lambda$CDM model. The derivatives here are with respect to $t$. Note that the FLRW metric in Eq. 1.5 assumes complete isotropy and homogeneity of the universe, and is an exact solution to the Einstein’s field equations Eq. 1.1. A common practice to study the observable universe, which shows lumpiness, is to approach FLRW universe to zeroth order, and to extend perturbatively to address the inhomogeneities and anisotropies.

The parameter $\rho$ involves the energy density of all the components of the cosmic inventory. Each component has its equation of state relating density with corresponding pressure. For
matter, \( p = 0 \), so \( \rho_m \propto a^{-3} \); for radiation, \( p = \rho/3 \); and thus \( \rho_r \propto a^{-4} \). An additional factor of \( a \) appears in the equation for radiation, implying that this energy decreases in the comoving volume due to the expansion of the universe itself. The energy density for \( \Lambda \), \( \rho_\Lambda = \text{constant} \).

1.1.5. Redshift, Hubble parameter and age

If a light photon is emitted at time \( t_e \) by a source and is observed at \( t_o \), then the cosmological redshift \( z \) of the source is defined as

\[
1 + z = \frac{a(t_o)}{a(t_e)}
\]

which is basically the measure of relative expansion of the universe.

The expansion rate of the universe is defined in terms of the evolution of scale factor. This rate is called the Hubble parameter \( (H) \) and is given by,

\[
H(t) = \frac{\dot{a}(t)}{a(t)}
\]

A subscript 0 is usually used to represent the value at the present time, \( t = t_0 \). The inverse \( H^{-1} \) gives a rough estimate of the age of the universe. A more accurate calculation involves a correction factor that depends on the fractional contribution of different species to the energy density. The parameter \( h = H_0/(100 \text{ km s}^{-1} \text{ Mpc}^{-1}) \) is commonly used and will be followed throughout this thesis.

1.1.6. Density parameter: \( \Omega \)

The total energy density given in Eq. 1.7 constitutes all the components of the cosmic inventory. If we consider an isotropic, homogeneous, flat universe, the total energy density, \( \rho = \rho_m + \rho_r + \rho_\Lambda \). The critical density for a given cosmological time, to stop the expansion of the universe is given by,

\[
\rho_{\text{crit}} = \frac{3H^2}{8\pi}
\]
The total density is commonly expressed in fractional form by,

\[ \Omega = \frac{\rho}{\rho_{\text{crit}}} = \Omega_m + \Omega_r + \Omega_{\Lambda} \]  

(1.13)

Note that Eq. 1.13 exhibits a minimal parameter \( \Lambda \)CDM model, where curvature, \( \kappa \), is taken zero and \( \Lambda \) is assumed constant. The model can be easily extended to account for other components such as neutrinos etc. Given the value of \( \Omega \) in Eq. 1.13, the universe can be:

- \( \Omega < 1 \): Infinite, Open universe, expands forever \( \implies \kappa < 0 \)
- \( \Omega > 1 \): Finite, closed universe, leads to “Big Crunch” \( \implies \kappa > 0 \)
- \( \Omega = 1 \): Flat universe, total density equals critical density \( \implies \kappa = 0 \)

The exact contribution of the different components of the universe decides what the fate of the universe will be. For example, in a flat isotropic universe, using equation of state, \( \rho \propto a^{-3(1+w)} \) in the FLRW equations,

- Matter dominated: \( w = 0 \) yields \( a \propto t^{2/3} \)
- Radiation dominated: \( w = 1/3 \) yields \( a \propto t^{1/2} \)
- Cosmological constant dominated: \( w = -1 \) yields \( a \propto e^{Ht} \)

1.1.7. Matter-radiation equality

Primordial perturbations are believed to seed the structure growth in the universe. For a \( \Lambda \)CDM scenario, structures are formed as a result of gravitational instability. But the growth of the structures is constrained by different factors at different times. The epoch of matter-radiation equality or the equipartition of matter and radiation is important as the nature of structure growth before and after this epoch is different. The minimum scale beyond which the structure formation grows is set by the Jeans scale, which comes from the dispersion relation for the sound propagation in the early universe. Perturbations of scales
smaller than the Jeans length are supported by pressure from random velocities, i.e., those perturbations are oscillatory and growth of structure can not occur. But the perturbations larger than the Jeans scale keep growing irrespective of the scale. At the matter-radiation equality, the Jeans scale is equal to the size of the horizon, i.e., the maximum proper distance for causal contact at that time. As the matter becomes dominant, the sound speed drops and Jeans scale reduces to zero and perturbations grow on all scales.

With this brief introduction, it is now useful to introduce some of the observational probes commonly used in modern surveys and some of these will be used in the analyses in this thesis.

1.2. Distances and their measurements

Measuring distance has been the biggest concern in astronomy so far. Accurate distance measurements have been done over the past few decades with ever increasing precision, with the advent of new detectors, large data set from surveys, and rapid improvements of calibrating techniques. The discovery of the accelerated expansion of the universe ([122,134]) using distant supernovae was one of the surprising and mysterious findings. Likewise, data from CMB, galaxy lensing and clustering, all seem to support the present-day accelerated expansion of the universe, arising from the proposed “dark energy” that dominates the present energy density of the universe at \( \sim 68\% \) ([125]). The remaining segment of the cosmic energy pie is mostly occupied by “dark matter”, that is suggested to be about 26% of the total energy content, while the small remaining portion seems to be occupied by the visible matter. All these evidences have direct relationships with distance measurements at a variety of scales. Below, I briefly review the standard distance measures used in modern astronomy and cosmology, that I will use in the analyses in later chapters.

It is convinient to define Hubble distance, \( D_H = c/H_0 \). Then one can establish different distances in terms of energy density by defining a function ([74]),

\[
E(z) = \sqrt{\Omega_r(1+z)^4 + \Omega_m(1+z)^3 + \Omega_k(1+z)^2 + \Omega_\Lambda} \tag{1.14}
\]
Here, $\Omega_\kappa \equiv 1-\Omega$ is the spatial curvature component. Then, at any redshift $z$, the Hubble parameter can be expressed as $H(z) = H_0 E(z)$.

1. **Comoving distance:** The line of sight comoving distance is given by

$$D_C(z) = D_H \int_0^z \frac{dz'}{E(z')} \tag{1.15}$$

2. **Angular diameter distance:** The transverse comoving distance between two objects at the same redshift but angularly separated on the sky is given by

$$D_M = D_H \frac{1}{\sqrt{|\Omega_\kappa|}} S_\kappa(\sqrt{|\Omega_\kappa|} \frac{D_C}{D_H}) \tag{1.16}$$

where, $S_\kappa(x) = \begin{cases} 
Sinh(x), & \text{for } \Omega_\kappa > 0 \\
x, & \text{for } \Omega_\kappa = 0 \\
Sin(x), & \text{for } \Omega_\kappa < 0 
\end{cases} \tag{1.17}$

Now consider a source of proper size $D$ at redshift $z$ (say, $r = r_1$ at $t = t_1$), subtending an angle $\delta \theta$ at the telescope (origin, i.e, $r = 0$ and $t = 0$), then the angular diameter distance, $D_A = D/\delta \theta = r_1a(t_1)$ is related to the transverse comoving distance by

$$D_A = \frac{D_M}{1 + z} \tag{1.18}$$

I will revisit the angular diameter distance when I talk about baryon acoustic oscillation (BAO) as a standard ruler in cosmology.

3. **Luminosity distance:** The total flux received by an observer at origin, from a source of bolometric luminosity $L$ is given by

$$F_{obs} = \frac{L}{4\pi D_L^2} \tag{1.19}$$
where, $D_L$ is called the luminosity distance. So luminosity distance connects the observed bolometric flux to the bolometric luminosity of the object. Luminosity distance is related to the transverse comoving distance and thus the angular diameter distance via

$$D_L = (1 + z)D_M = (1 + z)^2D_A$$

(1.20)

It will be useful now to define the magnitude system and the distance modulus. The ratio of radiant flux from an object at $D_L$ parsecs (pc) away to that if the object was at 10 pc away from an observer is related to object’s magnitude as

$$\frac{F_d}{F_{10}} = \left(\frac{10 \text{ pc}}{D_L}\right)^2 = 100^{\frac{M-m}{5}}$$

(1.21)

i.e, $M = m - 5(\log_{10}D_L - 1)$

(1.22)

where M and m are called the absolute and apparent magnitudes of the object. The distance modulus, $\mu = m - M$ is thus related to luminosity distance by

$$\mu = 5(\log_{10}D_L - 1)$$

(1.23)

An additional term “$K$-correction” will be needed (e.g. [75]) in practice to account for the observer frame shift in the spectrum due to expansion, as the particular photometric observation is limited to a passband filter covering a specific range of wavelengths in the rest frame. I will revisit luminosity distance when I talk about using supernovae as standard candles in cosmology.

I skip the other distance types here, as they will not be discussed in this thesis.

This thesis is organized in the following way. Chapter 2 describes a brief overview of supernova (SN) astrophysics, their generic photometric and spectroscopic properties and their use as “standardizable” candles in cosmology. In Chapter 3, I introduce the formalism of measuring baryon acoustic oscillation (BAO) with galaxy surveys and define statistical
tools that I will be using in the analyses later. Chapter 4 will focus on the ROTSE III telescope system, data reduction pipelines, observation strategy and prompt analysis of the SN discoveries and follow up photometry that I performed during my PhD. In Chapter 5, I present an extensive study using photometry and spectroscopy of SN 2013ej. This will be followed by the light curve analysis of two other SNe: SN 2012cg and SN 2013df. In Chapter 6, I present a study of cosmological distance measurements of 12 ROTSE core collapse SNe using the expanding photosphere method and present a measurement of Hubble constant from the study. Chapter 7 briefly describes the DESI survey, hardware and software systems for the data reduction. In Chapter 8, I focus my contribution and analysis on the spectroscopic simulations and the online data reduction pipeline development. I present an analysis on measuring effects of observing conditions on the BAO scale measurement with the simulated emission line galaxies for DESI in Chapter 9. Finally I conclude the outcomes and projections of the analyses in Chapter 10.
Chapter 2
SUPERNova ASTrophysics AND COSMOLOGY

The fates of massive stars end up with bright SN explosions, potentially disrupting the entire stellar body, and releasing a total energy of $O(10^{51})$ ergs. The types and causes of SN explosions depend on the final stages and environment of the exploding star. In this chapter I’ll discuss basic astrophysics of SNe, their types and their use as “standard” or “standardizable” distance calibrators in cosmology.

2.1. Supernova phenomenology

SNe are believed to explode in broadly two different ways: Thermonuclear runaway (Type Ia) and core collapse (Type II, Ib, Ic).

2.1.1. Supernovae Type Ia

SNe Ia are among the most luminous SNe. While the exact explosion scenario is not fully settled, this category is believed to involve at least one compact white dwarf, in a binary system with another white dwarf or a main-sequence or a red-giant star. The progenitor white dwarf star is a quantum degenerate system, where the mass of the star is supported by the electron degeneracy pressure and not by the pressure due to photon or particle kinetic energy like in other typical stars. These have concluded all of nuclear burning and fully lost their outer hydrogen and helium layers. What remains is a C+O core. The mass limit for a white dwarf is set by Chandrashekhar mass limit ($M_{ch}$) of $\sim 1.4 M_\odot$. In a binary system, the white dwarf receives mass from the companion through accretion. As the mass approaches $M_{ch}$, initial carbon ignition triggers thermonuclear burning. This heats up the matter further accelerating the burning process and electrons are no longer degenerate. Because the pressure now decreases due to the change of equation of state, the resulting instability leads the star
to undergo a thermonuclear runaway in a dynamical time scale and the star explodes in the form of Type Ia SN. In a binary white dwarf system, it is believed that the two white dwarfs spiral down to merge, and if the conditions are met to ignite the carbon burning on one, it explodes as SN, fully disrupting the progenitor.

The consequence of the explosion is an extremely luminous event. In the optical bands, the light curve rises to the peak in less than 20 days gaining a brightness \( \sim 3 \) mag, rolls over and decays more rapidly for 3-4 weeks. After about a month, the light curve decays slowly over several months, mostly dominated by the radioactive decay of the synthesized material. Spectroscopically, SNe Ia are characterized by absence of H and He lines. In just about 1 week, the spectra show lines of neutral and singly ionised intermediate mass elements (e.g., Mg, Si, S, Ca) to iron-peak elements (e.g., Fe, Co). A deep absorption profile of Si II \( \lambda 6355 \) at \( > 10,000 \) km s\(^{-1} \) at maximum brightness is one of the strongest classifying feature of Type Ia spectra. The late time spectra are dominated by blends of Fe emission lines (e.g., [53]). Fig. 2.1 shows a representative light curve and a spectrum at peak brightness for a variety of SNe.

While there are uncertainties on the exact explosion mechanism, and significant scatter in the photometric and spectroscopic observables, there exists a strong correlation of the maximum luminosity with the light curve shape. A commonly used technique is to look for the correlation of maximum absolute brightness with the \( \Delta m_{15} \) (difference in observed magnitude between maximum and 15 days after maximum) parameter (e.g., [123]). More recently, robust statistical treatment using both photometry and spectroscopy have been established to perform the light curve fitting for cosmological use (e.g [13, 65]). Due to this observed tight correlation of shape with intrinsic luminosity, SNe Ia have proven to be excellent distance measures in cosmology, that led to the discovery of accelerated expansion of the universe ([122, 134]).
2.1.2. Core collapse SNe

Core collapse SNe arise from the gravitational collapse of massive stars. Stars with masses $> 8 \, M_\odot$ undergo the collapse of a core made up of heavy elements when not enough radiation pressure exists after the nuclear burning, to support the mass. Within seconds (free-fall time scale), the core is believed to collapse to a neutron star or a black hole. Most of the gravitational potential energy released comes out in the neutrino energy, which escapes easily from the ejecta. Only about 1% of the total energy is observed in the kinetic energy of the ejecta.

Core collapse SNe are suggested to occur from a variety of progenitors. SNe IIP/IIL occur from the main sequence or red super giant (RSG) stars that have retained substantial hydrogen envelope even at the time of collapse. SNe IIP show a characteristic plateau in their light curve, which typically extends for about 100 days. The plateau phase is believed to arise as the extended hydrogen outer layer that sustains optical emission through recombination as the photosphere recedes and the outer envelope cools. After full recombination, the light curve suddenly falls off as the ejecta become transparent to the photons. The late time evolution is fully powered by the radioactive $^{56}\text{Co} \to ^{56}\text{Fe}$ decay at 1 mag/100 days. Type IIL SNe show a more linear decline of the light curve after the peak. As type IIP/IIL sample is growing, substantial debate has occurred on their exact boundary of classification (see Ref. [164] and references therein). It is also suggested that these events span a more continuous population (e.g., [3]) and potentially span a wider range of stellar populations. While a smaller mass of H-envelope in a progenitor of large radius ($\sim 1000 \, R_\odot$) is suggested for IIL events (e.g., [14]), hydrodynamical calculations (e.g., [105]) have shown that smaller ejecta mass cannot reproduce the observed luminosity of IIL SNe. Studies involving hydrodynamical modelling (e.g., [160–162]) and stellar evolutionary models (e.g., [147]), along with nebular spectra modeling (e.g., [82]), have shown conflicts on the progenitor masses of these events. Lack of identification of more-massive RSG progenitors has been termed as the “red supergiant problem” ([147]).
The early spectra of SNe IIP are mostly featureless blue continua. As the SN evolves, the optical spectra show P Cygni profiles of H Balmer lines and often He I $\lambda 5876$. As the ejecta cool and the recombination front moves inward, Na I D lines get stronger and weak lines of Fe II, Sc II appear. During the nebular phase, H$\alpha$ is quite strong along with prominence of emission features of [OI], [Ca II] etc. Type IIL show similar blue featureless continua at early times. H$\alpha$ becomes strong but the evolution is faster and show less prominent absorption wing, compared to SNe IIP ([53]). See Fig. 2.1 for a representative light curves and spectra of SNe.

Other types of core collapse SNe arise as IIb, Ib, and Ic in a progressive order of the removal of their hydrogen and helium envelopes, before the explosion ([29, 53]). These events are proposed to be arising from the progenitor as either a single massive Wolf-Rayet star with main sequence mass $> 30 \, M_\odot$ undergoing huge mass loss ([55, 177]) during the main sequence or later stage or from a lower mass He star in a binary undergoing mass loss through interaction ([126]). Highly stripped SNe Ic broad line are suggested to be connected with the long duration gamma ray bursts (e.g., [26]) but this appears to be not always the case, as GRBs are not observed on all occasions. See Ref. [176] for a review.

2.1.3. Other class of SNe

Energetic gamma-rays inside very massive stars (Zero Age Main Sequence (ZAMS) $> 100 \, M_\odot$) can undergo $e^-e^+$ pair production. As a result, the equation of state changes and the radiant pressure drops with density. An instability triggers compressing the massive oxygen core to undergo thermonuclear burning, unlike the iron-core collapse. Thus, the star totally disrupts with no remnants, producing a large amount of $^{56}$Ni, which will contribute to the luminosity of the SNe over time. Although this mechanism seems straightforward, there are questions regarding the natural birth of such massive stars. Star formation models predict that population-III stars that formed very early in the universe can attain such high mass; and thus are thought to undergo pair-instability SNe (see Ref. [22] for a review).
A new category of SNe that has evolved more recently with the advent of wide area surveys, is the super-luminous SNe (e.g., [131, 148]). Although these signatures show similarity with pair-instability SNe (e.g., SN 2010kd: [28]), there are instances where they are found not to result from such mechanism (e.g., SN 2006gy). A possible explanation of their high luminosity is the collision of the ejecta material with the previously expelled matter from the progenitor ([148]).

2.2. SNe as cosmological probe

Over the last few decades, cosmological studies using SNe have been in constant highlight. Most remarkably, the discovery of accelerated expansion of the universe ([122, 134]) using SNe was a breakthrough. More avenues have been developed since then to better constrain the cosmological parameters using SNe, and the field is growing as the sensitivity of the modern telescopes is rising. In this section, I briefly discuss the approaches of using SNe for cosmology. In particular, I present the mathematical framework of using IIP SNe, that will lead to the cosmological analysis in this thesis in Chapter 6.

2.2.1. Type Ia

Although substantial variation is observed in the light curve of Type Ia SNe, they have roughly uniform peak brightness. This and their high intrinsic brightness merits them to be one of the best standard candles in cosmological distance measurements.

The scatter in the peak brightness is significantly reduced when the peak brightness is correlated with light curve shape (e.g., $\Delta m_{15}$ parameter). The sample probing the highest redshift SNe have shown $\sim$5% accuracy in the distance measurement (e.g., [13]). The standard candle method exploits measurement of the luminosity distance ($D_L$) as defined in Section 1.2. Fig. 2.2, taken from Ref. [150], shows the compiled Type Ia Hubble diagram using “Union 2.1” sample from multiple surveys with events as far as $z=1.4$.

SN Ia cosmology is dominated from various systematics from explosion mechanism, photometric calibrations, etc. Among these is also the intervening dust in the Milkyway, the
host galaxy and the space in between. While earlier surveys have mostly relied on optical photometry, several current and the next generation surveys are targeting the observations in the near IR region of the electromagnetic spectrum. It has been observed that SN Ia are better standard candles in the near-IR and the observations there are also less sensitive to dust. Upcoming surveys are expected to gain significant improvement in the precision of the dark energy equation of state (e.g., [39]) from the near-IR data of Type Ia SNe.

2.2.2. Type IIP

IIP SNe are suggested to occur from stars of mass range $\sim 8 - 16 \, M_\odot$ (e.g., [147]). So the kinematic parameter space is much broader, giving rise to broad range of peak brightness and other properties. They, however, show a strong correlation of luminosity with the ejecta velocity, from which distance can be calibrated (standard candle method (SCM), e.g., [34, 61, 69]). Another calibration approach called expanding photosphere method (EPM) (e.g., [40, 88, 112]) is a geometrical technique, where the size of the expanding photosphere as
Figure 2.2. Hubble Diagram for the Union 2.1 SN sample from Ref. [150] using 580 Type Ia SNe. Solid line is the best fit for flat ΛCDM model of the universe. With this sample, they found the best fit $\Omega_\Lambda = 0.729^{+0.014}_{-0.014}$ for a flat ΛCDM model.
a function of time is used to calibrate distance. The idea is similar to SCM, but it directly relies on the effective temperature and assumes the photosphere to be evolving isotropically, homologously as a diluted black body. The homologous expansion here simply means that the ejecta velocity is a function of radius, i.e., the outward material moves faster than inward material of the ejecta. In the case of SCM, the explicit requirement of temperature and dilution parameter as in EPM, are marginalized by the requirement of photometric color terms. However, SCM also relies on the assumption of homologous and isotropic symmetry of the ejecta.

While SCM requires multi band photometry, EPM can be performed on a single band. I present the mathematical formalism of the EPM method in Section 2.3 and analysis on ROTSE SNe IIP sample will be presented in Chapter 6. Here I discuss some of the limitations and care that should be taken for such an analysis.

IIP SNe are UV rich at early times. The degree of deviation of the expanding photosphere from a perfect black body spectrum is obtained through modeling of stellar atmosphere (e.g., [38, 42]). These studies show that the deviation can be both temperature and wavelength dependent. While temperature dependence seems to be corrected with the derived effective color temperature, the wavelength dependence is more difficult to correct because of different scattering processes in the continuum and line forming regions of the ejecta. Ref. [38] has shown that the wavelength dependence is marginal in the optical bands for first few weeks after the explosion, and EPM should be performed during these epochs using optical band photometry and spectroscopy.

EPM and SCM based Hubble diagram of Type II SNe at the highest redshift so far are shown in Fig. 2.3. The left plot is the EPM analysis taken from Ref. [61]. Solid line represents a Hubble constant of $70 \text{ km s}^{-1}\text{ Mpc}^{-1}$ in a flat cosmology ($\Omega_M = 0.3; \Omega_\Lambda = 0.7$). The right plot is for the SCM analysis, taken from [34] on a different data set from four different surveys. These results suggest that distance measurement with the current IIP sample are at about 12% accuracy and both methods yield consistent results. As higher redshift objects will be discovered in large numbers in the coming years, these techniques need
refinement and calibration methods need to be more robust to provide strong discrimination power on the cosmological models. One of the key tasks of my analysis is to refine such calibration techniques for EPM parameters from the limited data. An accurate distance measurement with limited data can be a significant advantage of EPM over other methods which need better time series follow up of individual events, potentially in multiple bands.

2.3. Mathematical formalism of expanding photosphere method

Following the procedure described by Ref. [170], the basic equation for EPM is

\[ t = D \times \left( \frac{\theta}{v_{\text{phot}}} \right) + t_0, \]  

where \( t \) is the observation time, \( D \) is the distance, \( \theta = 2R/D \) is the angular size of the photosphere, \( v_{\text{phot}} \) is the velocity of the photosphere at \( t \), and \( t_0 \) is the moment of explosion or more explicitly the moment of shock-breakout.

With the assumption of isotropic radiation from a blackbody, one can write for the observed flux

\[ f_{\lambda}^{\text{obs}} = \theta^2 \pi B_{\lambda}(T) 10^{-0.4A_{\lambda}} \]

where \( B_{\lambda}(T) \) is the Planck function for the blackbody at effective temperature \( T \). \( A_{\lambda} \) is the galactic extinction for the given band. Subscript \( \lambda \) on the parameters is to be considered as index for the observed photometric bands. SNe IIP, however, depart from a true blackbody as the thermal photons generate from deep in the atmosphere, and not from the photosphere, the last scattering surface (e.g., [17, 83]). As such, \( \theta \) in Eq. 2.1 would correspond to the thermalization layer and \( v_{\text{phot}} \) corresponds to the photosphere (optical depth, \( \tau = 2/3 \)) and the atmosphere is essentially gray (e.g., [42, 83]). To account for this, a scaling factor (also called dilution factor or distance correction factor) is introduced as,

\[ \zeta = \frac{R_{\text{therm}}}{R_{\text{phot}}} \]
ζ is treated mostly wavelength independent in the optical and infra-red and Ref. [42] has shown that it is a monotonic function of $T$ alone, at least for several weeks after explosion. Thus, the epochs in the EPM analysis are to be chosen where this condition is satisfied and wavelength dependence is marginal. Evaluation of ζ requires complex computation of a realistic model atmosphere, and comparison with the blackbody flux. I use the prescription from commonly used Ref. [38]. Although Ref. [42] is also commonly found in literature.

Accounting for this dilution correction, Eq. 2.2 can be written as

$$f_{\lambda}^{\text{obs}} = \zeta_{\lambda}(T)^2 \theta^2 \pi B_{\lambda}(T) 10^{-0.4A_{\lambda}}$$  \hspace{1cm} (2.4)$$

If we consider full extinction corrected bolometric flux, we can estimate θ from the bolometric light curve by integrating over all wavelengths and using

$$\theta = \frac{1}{\zeta(T)} \sqrt{\frac{f_{\text{bol}}}{\sigma T_{\text{eff}}^4}}$$  \hspace{1cm} (2.5)$$
where $\sigma$ is the Stefan–Boltzmann constant. Most of the times, the bolometric flux cannot be measured and any accurate calibration requires a good sampling of multiband photometry. However this is not a required condition for EPM analysis. It is a common practice to use a specific wave band, where the filter response function is convolved with black body flux, given the magnitude of the SN for that wave band. As you will see later, many SNe in the ROTSE IIP sample lack observations to yield full bolometric flux, so I use the prescription of deriving effective synthetic blackbody flux by convolving with the response function $R_\lambda(\lambda)$ of a particular photometric band,

$$b_\lambda(T) = \int_0^\infty R_\lambda(\lambda') \pi B(\lambda', T) d\lambda' \quad (2.6)$$

Then, for each observed flux and $\zeta_\lambda(T)$, $\theta$ can be solved using

$$f^\text{obs}_\lambda = \zeta_\lambda(T) \theta^2 b_\lambda(T) 10^{-0.4A_\lambda} \quad (2.7)$$

The observed flux is always supposed to be treated as the $K$-corrected flux, while the parameters $\zeta_\lambda$ and $b_\lambda$ are in the SN rest frame. $K$-correction is sufficient to account for the $(1+z)$ factors that would have appeared into the equations for high redshift cases. With this approach, the derived distance is the luminosity distance and not the angular diameter distance. For more discussion, see Ref. [61] and references therein. Determining distance using Eq. 2.1 then falls to determining $v_{\text{phot}}$ and $T$, which can be obtained directly from observations. Both $D$ and $t_0$ can be simultaneously solved by minimizing the $\chi^2$,

$$\chi^2 = \sum_j \left[ \frac{\theta_j}{v_{\text{phot},j}} - \frac{(t_j - t_0)}{D} \right]^2 \sigma_j^2 \quad (2.8)$$

where $\sigma_j$ is the uncertainty on the quantity $\theta_j/v_{\text{phot},j}$.

To summarize this chapter, SN cosmology has been very successful in probing the low to intermediate redshift ($z \sim 1$) universe. It is necessary to continue improving and building methods and analysis techniques as upcoming surveys scan larger volume of the universe.
While SNe Ia bear higher intrinsic luminosity, SNe IIP can be equally rich as they seem to be the most frequently occurring SNe ([95]). So larger upcoming surveys are expected to inflate their discoveries by at least an order of magnitude. With all the richness of SN cosmology, a dense time series observational follow up of each event is difficult with flux limited surveys. Even with the sensitivity of the most powerful telescopes present, SN may not be the ideal tool to probe much deeper universe, and one has to rely on other methods. I will next introduce a different method of studying cosmology with deeper data using galaxy surveys in the next chapter, and postpone my SNe analyses to later chapters.
After the Big Bang, the cosmic fluid included photons, baryons, dark matter and neutrinos, all of which were coupled together. The whole plasma was expanding out with sound wave excitations from initial perturbations. The radiation pressure would stop the matter from collapsing. After about 380,000 years after the Big Bang, the universe cooled enough that the ions could combine with the electrons and the radiation and matter fully decoupled. This epoch is called the recombination or the decoupling epoch, and happened at $z \sim 1100$. As a result of the acoustic waves that propagated until the decoupling epoch, oscillation patterns remained imprinted in the matter distribution and also on the temperature anisotropy of the cosmic microwave background (CMB) radiation. High resolution maps of the temperature anisotropy fluctuations have been measured by the COBE, WMAP and Planck satellites over the years. The oscillation is also measured with the galaxy clustering data (e.g. Baryon Oscillation Spectroscopic Survey (BOSS)). This signature imprint of the oscillation sets a length scale of 110 Mpc/h, i.e, about 150 Mpc comoving separation. Such an oscillation is termed as Baryon Acoustic Oscillation (BAO). The approach of redshift surveys is to measure this BAO scale at various redshifts, to infer both $D_A(z)$ and $H(z)$. If $z_1$ is the redshift of recombination, then one can write the acoustic scale ([44,78]) by

$$r_s = \int_{0}^{t_1} \frac{c_s(t)}{a(t)} dt = \int_{z_1}^{\infty} \frac{c_s(z)}{H(z)} dz$$

(3.1)

where $c_s$ is the relativistic sound speed, which depends on the photon pressure and baryon to photon ratio. The exact evolution of $H(z)$ for $z > z_1$ is sensitive to the ratio of matter density to radiation density. These have been measured from the relative heights of acoustic peaks of the CMB anisotropy power spectrum ([174]). But BAO measurements can directly
measure the expansion rate from the galaxy clustering. Measuring BAO across a wide range of redshift directly constrain the expansion rate of the universe. Before I show current state of the art measurements of expansion rate of the universe using large survey data, I present some connecting ideas from theory and present some tools to make cosmological measurements.

3.1. Linear perturbation theory

A perturbative approach is adopted to address the evolution of inhomogeneities in an expanding universe. Before decoupling, the universe was hot and dense with ionized gas. The Thomson scattering cross-section due to free electrons interacting with the CMB photons was large, such that the mean free path scale was smaller than the Hubble radius. So the electrons, ions and photons were in tight coupling potentially giving rise to long wavelength perturbations. When the nuclei combined with electrons, and photons decoupled, and the baryonic over-densities were left to gravitational instability in the absence of radiative pressure. Central dark matter perturbations, which did not interact with the fluid, were already subjected to gravitational instability but gravitational dragging developed between the dark matter and baryonic concentrations as illustrated in the Fig. 3.1. It is suggested that all species of the cosmic inventory shared similar fractional density contrast on the primordial perturbations (e.g., [45]).

I briefly describe here a linear perturbative theory from a Newtonian approach. Let $x = r/a(t)$ be the comoving spatial coordinate. Then the peculiar velocity field, $v = \dot{r} - Hr$ describes the motion departing from Hubble expansion. Let $\phi(x, t)$ be the gravitational potential and $\rho(x, t)$ be the matter density. When the mean free path of the particles is small, we can write the Euler equation for the ideal fluid following Ref. [30] as

$$\frac{\partial(a\mathbf{v})}{\partial t} + (\mathbf{v} \cdot \nabla_x)\mathbf{v} = \frac{-1}{\rho} \nabla_x p - \nabla_x \phi$$

(3.2)
Assuming an irrotational velocity field, one obtains the continuity equation conserving matter and the Poisson equation for the gravitational potential as,

\[
\frac{\partial \rho}{\partial t} + 3H \rho + \frac{1}{a} \nabla_x (\rho \mathbf{v}) = 0 \tag{3.3}
\]

\[
\nabla^2_x \phi = 4\pi G a^2 (\rho - \rho_0) = 4\pi G a^2 \rho_0 \delta \tag{3.4}
\]

where, \(\rho_0 = \bar{\rho}(\mathbf{x})\) is the mean density and \(\delta(\mathbf{x}) = \frac{\rho(\mathbf{x}) - \bar{\rho}(\mathbf{x})}{\bar{\rho}(\mathbf{x})}\) is the density contrast or the overdensity field.

The Euler, continuity and the Poisson equations above can be linearised by expanding \(\rho\), \(\mathbf{v}\) and \(\phi\) perturbatively and taking only the first order terms. Again this is true when perturbations are small. One then obtains ([30]),

\[
\ddot{\delta} + 2H \dot{\delta} - \frac{3}{2} \Omega H^2 \delta = 0 \tag{3.5}
\]

where \(\Omega\) describes the growth of structure ([120]). Considering a flat matter dominated universe, this yields two independent solutions for \(\delta\):

\[
\delta(\mathbf{x}, t) = D_{\pm}(t) \delta(\mathbf{x}), \tag{3.6}
\]

with growing mode \(D_{\pm}(t) \propto a(t) \propto t^{2/3}\) and the decaying mode \(D_{-}(t) \propto t^{-1}\)

3.1.1. The correlation function

The solution found above for the growth of density field is for a single mode. Superposition of such modes of different amplitudes demands statistical treatment of the perturbation, and this is why the observations from surveys have to be treated statistically. The density field at early time constitutes adiabatic perturbations, and the fluctuations are close to Gaussian ([124]). This means, the complete statistical description is obtained from the covariance,
i.e. the two-point correlation function in this case, which is defined as

$$\xi(x_1, x_2) = \langle \delta(x_1) \delta(x_2) \rangle$$

(3.7)

Due to the statistical isotropy and homogeneity of the universe, it only depends on the distance between the points.

$$\xi(x_1, x_2) = \xi(x_1 - x_2) = \xi(|x_1 - x_2|)$$

(3.8)

If one considers a single primordial perturbation, then as sound waves propagate outward, a spherical wave of baryons+photons is driven away. As the photons decouple from the baryons, the shell of baryons is left around the original concentration of dark matter at a radius set by the sound horizon, also referred to as the surface of last scattering. These overdense regions (both the central peak and shell) grow over time via gravity ([44, 45]). So when looked at the correlation function of the distribution of sufficiently well-sampled galaxy distributions, one expects to see a small overdensity above a smooth background at the sound horizon, i.e., BOA scale. Fig. 3.1 shows the development of mass profiles of different species as the sound propagates in the early dense universe through a series of stages involving decoupling of each of the species from the fluid at different times and at much later stage where structure formation takes place at the density peaks.

3.1.2. The power spectrum

The Fourier transform of the overdensity field $\delta(r)$ is given by

$$\delta(k) = \int \delta(r) e^{i k \cdot r} d^3 r; \delta(r) = \int \delta(k) e^{-i k \cdot r} \frac{d^3 k}{(2\pi)^3}$$

(3.9)

Then, the power spectrum is given as the covariance of the Fourier modes,

$$P(k_1, k_2) = \frac{1}{(2\pi)^3} \langle \delta(k_1) \delta(k_2) \rangle$$

(3.10)
Figure 3.1. Acoustic peaks generated using Linear theory to point-like primordial overdensity. Figure is taken from Ref. [174], that was reproduced using Ref. [45]. In the subplots are shown snapshots of perturbed mass profiles at different redshifts as the perturbations evolve for four species: dark matter (black), baryonic matter (blue), radiation (red) and neutrinos (green). After the photon-baryon decoupling, the baryons and CDM overdensities undergo gravitational instability. Thus at late times, galaxy (and thus halo) formation is more likely to occur near the origin, and at the spherical shell at comoving radius of 150 Mpc.
Using property of isotropy and homogeneity,

\[ P(k_1, k_2) = \delta_D(k_1 - k_2)P(k_1) \] (3.11)

With \( P(k) \) the Fourier transform of the correlation function,

\[ P(k) = \int \xi(r)e^{ik \cdot r}d^3r; \quad \xi(r) = \int P(k)e^{-ik \cdot r}\frac{d^3k}{(2\pi)^3} \] (3.12)

Thus \( \xi(r) \) and \( P(k) \) form a Fourier pair and the information content is same. Often, it becomes convenient to measure one over the other ([66]). Because the Dirac delta function appears naturally in Eq. 3.11, each Fourier mode propagates independently.

The original primordial power spectrum (inflationary models suggest a power law form: \( P_0(k) \propto k^n \) (e.g [64])) is connected to the later power spectrum via a wave number dependent function called a Transfer function, \( T(k) \) ([30]). All the model dependence is generally embedded into this function. The inflationary power spectrum mostly survives for CDM models, while strong dissipation of power is seen in a relativistic Hot Dark Matter (HDM) scenarios. Late time structure formation are thus completely different for the two models. In the linear regime, \( \delta(k) \ll 1 \), and the power spectral shape is preserved at all scales larger than the Jeans length ([30]).

In the following section I describe how these statistical estimators are measured in the redshift surveys. In the later section, I highlight some of the theoretical challenges from non linearity of clustering that may impact the BAO scale measurement.

### 3.2. Measurement of \( \xi(r) \) and \( P(k) \) from galaxy surveys

Galaxies are not randomly spread in space, but they make clusters and voids. With 2-pt correlation function, we measure the deviation from some uniform random distribution. What this means is that if \( \Delta V_1 \) and \( \Delta V_2 \) are two small volumes and \( n = N/V \) is the average number density, then the joint probability of finding one galaxy in \( V_1 \) and another in \( V_2 \)
Volumes is given by

\[ \Delta P = [1 + \xi_g(r)] \frac{\Delta V_1}{V} \frac{\Delta V_2}{V} \]  (3.13)

Thus \( \xi_g(r) \) is a measure of excess probability that is zero if galaxies were uniformly distributed, positive if galaxies cluster, and negative if galaxies avoid clustering.

In galaxy surveys, one actually observes the position of a representative sample of galaxies in redshift space. In linear perspective, the convention is that an overdensity in a matter field is traced by the distribution of galaxies. But the observations in surveys have spatially varying noise. This has to be correctly understood before an inference is made on the clustering measurements.

If \( \bar{n}(r) \) denotes the expected mean number density of galaxies at position \( r \), after applying the mask of the survey that carries the information of target selections for the survey, and \( n(r) \) is the observed number density then the overdensity of the galaxies can be written as

\[ \delta(r) = \frac{n(r) - \bar{n}(r)}{\bar{n}(r)} \]  (3.14)

Then, the two point correlation function is given as

\[ \xi(r_{ij}) = \langle \delta(r_i) \delta(r_j) \rangle \]  (3.15)

In practice, the survey mask is described by a random catalog of objects. Then, the correlation function estimation becomes a pair counting problem. The most commonly used estimator in recent surveys (e.g., BOSS) is that of Landay and Szalay ([91]).

\[ \bar{\xi}(r) = \frac{DD(r) - 2DR(r) + RR(r)}{RR(r)} \]  (3.16)

where \( DD(r) \), \( DR(r) \) and \( RR(r) \) represent the number of galaxy-galaxy, galaxy-random and random-random pairs in a bin with center \( r \); respectively normalized by maximum possible number of pairs in each case. It is common to extend this approach of computing correlation function in the radial and transverse to the line of sight direction and using wedges.
with respect to the line of sight to measure anisotropic correlation function (e.g. [85]). Measurement of the spherically averaged correlation function on the BOSS DR9 CMASS galaxy sample is shown in Fig. 3.2, taken from Ref. [4]. The parameter $\alpha$ gives the relative shift on the BAO scale from a fiducial cosmological model. A value $\alpha = 1$ means the measurement is consistent with the assumed fiducial cosmology. On the right panel is shown the spherically averaged power spectrum from the same paper. The inset shows the galaxy power spectrum divided out by a smooth model without BAO (no-wiggle model).

It is now useful to point out some ideas on how the non-linear effects can modify the BAO feature, and how they can be corrected.

### 3.3. Non-linear evolution of structure

Structure formation and growth are highly non-linear at small scales and the galaxy-dark matter mass tracing becomes complicated. Bulk motions can smear the initial separation of overdensities and peculiar velocities can also smooth the BAO feature. These effects can seriously distort the location and shape of the BAO feature, thus restricting the ability to use them as a standard ruler (e.g. [172]). Ref. [45] showed that the large scale of acoustic peaks are less sensitive to non-linear evolution. From the comparision of dark matter particles’ comoving separation at high redshift and at $z = 0$, it was shown that the rms of the separation for high $z$ is only $\sim 10 h^{-1}$Mpc, compared to $8 h^{-1}$Mpc set by Silk damping.

Fig. 3.3 shows that the peak in the correlation function broadens with lower redshift in N-body simulations ([139, 174]). The right panel shows the effect in the real space matter power spectrum on the simulation from Ref. [140]. The broadening of the BAO feature at lower redshift is equivalently shown by decreased amplitude of the wiggles for higher wave number. This effect is usually treated by adding an rms of pairwise displacement, $\Sigma_{NL}$, in quadrature with the width predicted from the linear model. Also in the power spectrum, higher mode power is damped compared to the linear model by a factor $e^{-k^2\Sigma_{NL}^2}$. The power is boosted at small scales because of non-linear clustering. In real survey measurements, it
Figure 3.2. Left: Spherically averaged correlation function of BOSS DR9 CMASS galaxy sample from Ref. [4] using Landy and Szalay estimator. The parameter $\alpha$ is consistent with value 1 suggests the consistency of the BAO scale measurement with the fiducial cosmology. BAO peak is observed at the middle of the plot at $\sim 110$ Mpc/h. Right: Spherically Averaged power spectrum from the same paper on the same data, calculated using Ref. [52]. The inset shows the galaxy power spectrum divided by the smooth no-wiggle model.
is common to marginalize over any broadband scale-dependent biases, but the BAO scale itself is assumed to be not shifting significantly because of broadening, or the linear model has to be corrected ([174]). Although the scale may not change significantly, the broadening decreases the precision of measurement.

Reconstruction techniques have been used to sharpen the BAO and improve precision. Studies using both mock data (e.g., [114]) and in particular survey data (e.g., [115]) have been effective to reconstruct the BAO feature and reducing the error bars by a factor of 1.7 (for SDSS-II Luminous Red Galaxy sample). However no such improvement was observed in the CMASS sample. So the efficacy of reconstruction has not been universally justified ([4]). Before wrapping this chapter I give a brief review of cosmological sensitivity from the existing surveys.

3.4. Current state of the art

In the last two decades, astronomical surveys presented us an unprecedented set of measurements that massively boosted our understanding of the universe. This has also raised
a series of new fundamental questions that warrant better and more sensitive tools to be answered. Through multiple probes spanning the nearby and far away universe, the astronomical surveys have been able to put strong constraints on the cosmological models but the full history of the dynamics of the universe require much deeper and wider probe of the cosmos. The sensitivity of the current experiments lack enough discriminating power on, for instance, the equation of state parameter \( w \) of the dark energy models (e.g., \([36]\)). A small change in \( w \) can significantly impact the cosmological evolution. The existing data points at different redshift bins are not dense enough to precisely describe the expansion history of the universe. Current state of expansion rate from few of the surveys is shown in Fig. 3.4. A densely sampled expansion history over a wide range of redshift from a single experiment will not only have an advantage of zero cross-survey systematics but also be invaluable in understanding these deep issues. The upcoming Dark Energy Spectroscopic Instrument (DESI) exactly aims to answer such questions increasing the statistics by an order of magnitude over a wide range of redshifts. DESI is expected to constrain the cosmological parameters at sub-percent level precision. Through BAO and Redshift Space Distortion (RSD) measurements, DESI will help understand the nature of dark energy and structure growth, by 3D mapping of the matter distribution for about the last 10 billion years.

I will follow up with the description of DESI experiment and its expected sensitivity in Chapter 7, and BAO analysis for DESI simulated galaxy samples in Chapter 8. Before that, I present the ROTSE III telescope system, its SN program at SMU and the SN analysis that I conducted during my PhD in next two chapters.
Figure 3.4. Current state of expansion rate measurements using $H_0$ with Hubble Space Telescope (blue), BAO with SDSS galaxies (black), BAO with BOSS galaxies (red square), BAO with BOSS Lyman-$\alpha$ forest (red circle) and BOSS Lyman-$\alpha$ - quasar cross correlation (red x). Plot is taken from Ref. [36].
Both wide and deep sky surveys have advanced remarkably over the last two decades. Robotic instruments have been crucial in systematically observing transient events, which are and have been of fundamental interest both in astronomy and cosmology. The Robotic Optical Transient Search Experiment (ROTSE) ([2]) telescope system provided a unique global coverage for both SN and GRB optical observations from 2003–2012, and from Texas since then. In this chapter, I give an overview of the ROTSE telescope system, operation and scheduling, its use in the supernova search, follow up spectroscopy with larger telescopes, and the ROTSE SN sample.

4.1. ROTSE telescope system

The current ROTSE system is the third generation of the experiment. Until early 2013, the experiment operated in 4 nodes: ROTSE-IIIa - Australia; IIIb - USA; IIIc - Namibia; IIIId - Turkey. All of the sister telescopes are clones in design and instrumentaion. While primary operation and control of the global system was managed by the University of Michigan (PI: Carl Akerlof), ROTSE collaboration has been a global collaboration. The SMU team joined in the formal collaboration in the fall of 2012. In early 2013, IIIa and IIIc systems were decommissioned, and SMU took over the ownership of the IIIb telescope at McDonald Observatory, Texas. While the ROTSE IIIb has been the most operational since, efforts are being made to bring the IIIId back in normal operation, and signs are encouraging. From fall 2012 to early 2016, I was the primary person for the ROTSE IIIb telescope operation, maintenance and survey data analysis.
4.1.1. Telescope and design

Each ROTSE telescope system is built inside a stand-alone enclosure, including the telescope mount and optical assembly, camera, clam-shell cover, a weather station and the computing facilities. While the mount control computer is separate, the computing system involves two additional computers, one to control the central data acquisition system (DAQ) for the telescope, and another is a dedicated quasi-realtime SN analysis engine. The online GRB analysis is also performed from the main control computer of the system.

ROTSE III telescopes each have a 0.45m primary mirror, fully robotic, Cassegrain system. The corrector is all-refracting and provides a final focal length of 0.85m, the focus located at \( \sim 75 \) mm above the primary vertex. The telescope is designed for fast responses to the Gamma-ray Burst (GRB) triggers from satellites like Swift, Fermi, HETE-2, and INTEGRAL. Not only are the GRB optical counterparts short lived by nature, their astrometric uncertainty is also quite large. ROTSE telescopes have proven ideal instruments to address both of these needs. The equitorial fork mount holds the 18 inch optical tube and provides excellent slew torques, accurate tracking and good mechanical tolerances. The maximum slew acceleration is \( 16.4^0 \text{ s}^{-2} \) in RA and \( 20.6^0 \) along the Dec axes. The maximum slew velocity of \( 35^0 \) provides the entire horizon to horizon coverage in 8s, thus making it ideal to observe the ephemeral optical signal from GRBs as quickly as possible. The positional uncertainty is also addressed by the wide \( 1.85^0 \times 1.85^0 \) field of view (FOV).

The ROTSE camera has a Marconi 2045 \( \times \) 2049 back-illuminated thinned CCD with pixel size of 13.5 \( \mu \)m, and photometry performed is unfiltered (open-CCD). The pixel scale is 3.28", which undersamples the seeing PSF, but was designed to attain the desired field coverage. The detector is supplied with a cooling system to operate the imaging in a stable temperature of \( \sim 40^0 \text{C} \), that can be routinely reached at ambient temperature of 20\(^0\text{C}\). This cooling is achieved with a propylene glycol heat transfer loop driven by an air-cooled recirculator. The camera readout is \( 8e^- \) and the images are totally sky dominated in 5s exposures. The typical limiting magnitude for a 60s exposure image is about magnitude 18.5. As the CCD is operated in unfiltered mode, the quantum efficiency spans a broad
range in near UV, optical and near IR ranging from 3000 to 10000 Å; with the peak response of approximately 85% at about 5500 Å ([132]). Fig. 4.1 shows the CCD quantum efficiency of ROTSE unfiltered CCD. Overplotted are the broadband SDSS $ugriz$ filter response curves for comparison.

4.1.1.1. Operation

In this section, I briefly describe the generic telescope operation mode, calibration and science frames, targetting, cadence of observing, configuration and observing priorities for the ROTSE telescopes.

Each ROTSE telescope is operated in automated mode by default. While user intervention is facilitated any time, this is not needed in general. The full automation is facilitated by the ROTSE DAQ system, which comprises an integrated system interconnected with several daemon interfaces that are independently specific for a distinct purpose of the telescope system. A schematic diagram of ROTSE DAQ system is shown in Fig. 4.2.

**Calibration images:** Before observing starts for a given night, a series of calibration frames are taken. Twilight sky flats are normally taken during the evening twilight. These twilight flats will be used to perform a pixel level calibration to remove any disproportionality of pixel sensitivity due to imaging effects such as vignetting, that arises from reduction in brightness at the periphery relative to image center. A series of dark exposures are taken with configured exposure lengths of 5, 20 and 60 seconds. The dark exposures will be used to subtract a pedestal (bias) and quantify the dark current in the electronics. The exposure lengths are chosen such that the integration time matches with the science images.

**Survey imaging and target fields:** Science images refer to the field images on the sky. The survey basically includes two major modes: a GRB optical follow-up mode and a SN search mode. Each science observing is assigned a unique field ID, and supplied with the definite observing configuration including exposure times, observing cadence, target coordinates,
Figure 4.1. Response function of ROTSE open-CCD. SDSS $ugriz$ band filters curves are plotted for comparison.
Figure 4.2. A schematic diagram of ROTSE data acquisition system. Figure is taken from Ref. [2] and shows the interconnection of several daemons with the main “rotsed” daemon.
priority and so on. During a normal operation of a GRB optical follow-up, first, a series of images of exposure lengths 5s are taken to address the potentially rapid evolution of optical spectral energy distribution (SED) of the GRBs. Then a series of 20s exposure are taken, followed by several 60s exposures. The number of exposures depend on the time spanned since trigger was received and the observing conditions at that time. In the SN search mode, imaging is performed pointing at various observable target fields that are already configured. These target fields are chosen to cover some of the brightest nearby clusters and some individual galaxies that are supplied in the configuration. Occasionally, some specific targets (e.g., variable stars or a nova) are also configured for observing over a designated period of time.

**Observing cadence and priorities:** In an automated mode, the telescope usage involves GRB triggers with the highest priority at any time, with the remaining time for SN search. Although there is full flexibility in the specific configuration for each exposure type, adjustments are rarely needed. The SN target fields are generally configured to have different priorities given the galaxy number density in a given field. The SN fields are always observed with a pair of exposures, each with 60s. This is repeated half an hour later and the whole chain is repeated again after 2 hours. So most of the fields will have 4 pairs of images per night. A few fields have more than 4 pairs while others have only 2 pairs of imaging. This relies on the selection criteria in the scheduler and sometimes a seasonal dependency of observing. At SMU, we have practiced a configured follow up scheduling of events of interest, with more or less arbitrary choice of configuration parameters. These events are selected in accordance with the ongoing science projects within the collaboration, after their discovery and classification have been publicly announced.

**4.2. ROTSE data processing system**

The ROTSE III system has adopted a standard photometric data-reduction pipeline. In this section, I briefly describe the data processing of raw ROTSE images for SN analysis.
4.2.1. Data reduction pipeline

In the automated fashion, the data reduction pipeline processes each raw image to subtract dark and pixel-flat as part of the initial calibration or preprocessing. The raw image is not stored on disc but the data after these two calibration steps are saved as FITS\(^1\) files. An object list file of the sources identified by sextractor\(^2\) is created and saved to disc. Subsequently, the identified sources are calibrated to the magnitudes from USNO\(^3\) A2.0 catalog by default. This step uses an algorithm matching triplets of stars and determining rotation, scaling and warping to get the coordinates given the CCD projection onto the sky. The calibration parameters such as readnoise, calibration file names, sextractor parameters and the image to sky mapping matrix terms etc. are all propagated into the headers or other extensions of the FITS files. The sectractor derived parameters include average FWHM of the sources’ point spread function (PSF), measured sky background etc. The ROTSE operation manual\(^4\) gives details of the ROTSE Pipeline.

The automated pipeline performs photometry over the field in quasi-real time. This is used for data quality assurance (QA) and the QA metrics are displayed on the web interface for real time monitoring of the data, and for offline analysis as well.

4.2.2. Photometry

While an aperture photometry is sufficient for transients such as GRBs, this is not usually the case for SNe, where the light from the host galaxy can be significant and needs to be correctly subtracted. After the raw images are read out and corrected for dark and flat-field, the GRB prompt burst images are subjected to photometry and calibrated to the USNO A2.0 catalog and compiled to the lists. As more than one list is available, a match

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\(^1\)https://fits.gsfc.nasa.gov/fits_documentation.html

\(^2\)https://www.astromatic.net/software/sextractor

\(^3\)http://www.usno.navy.mil/USNO

\(^4\)http://www.rotse.net/equipment/docs/rotsedocs.pdf
structure is created. It is a data structure of calibrated lists that contains the information of the identified and unidentified sources and transients. A time series scanning of the GRB optical counterparts can be easily detected and subject to further analysis for discovery etc.

SNe data analysis involves a different pipeline. First the images are astrometrically alligned and coadded in the set of first four, the last four and the full stack of images for the night. The final coadded image is then subjected to the differencing algorithm (see Section 4.2.4.2), where it is compared to a template image of the field taken at different time. The idea is to perform a subtraction of the template image from the new image, such that any significant detection in the residual image can be identified as a potential SN candidate. Any photometric detection $> 5\sigma$ is posted in the web interface for cross checking against various catalogs for variable stars, asteroids, comets etc. Once no record for the transient is found accross several catalogs, an offline analysis is performed taking images over a period of time. This will help confirm whether the detection is indeed a SN candidate or an unknown source such as cosmic ray or just a subtraction artifact due to complicated underlying host background. If the event is promising and passes all the selection cuts, a spectroscopic observation is requested for confirmation. Once confirmed, the discovery is publicly announced through circulars such as Central Beareu of Electronic Telegram (CBET) or The Astronomical Telegrams (ATel).

The final photometry is performed using relative photometry with the field stars close to the event that are matched against the catalog for each epoch. The systematic uncertainties due to observational effects are propagated to the final photometry.

4.2.3. Spectroscopic follow up

The gigantic 9.2m Hobby Eberly Telescope (HET), located close to the ROTSE III b site has been the most used spectroscopic instrument to obtain spectra for the ROTSE SNe. This is true to make a confirmation and thus discovery or time series spectroscopic follow up as the SNe evolve. Note that there are other instruments from which the optical, UV and IR spectra have been obtained and studied for ROTSE SNe. As the SN candidate is
identified, the coordinates are passed to the HET and a spectrum can be obtained as early as possible. Usually, a pair of observations is requested with the low resolution spectrograph (LRS) [73] using the GG385 filter and 2" slit. This yields a spectrum devoid of order overlap from the grating in the range 4100-7700Å. The spectra are then reduced from the 2D CCD image with pipeline scripts in Image Reduction and Analysis Facility\(^5\) (IRAF) and Interactive Data Language\(^6\) (IDL). Arc lamp spectra are reduced for wavelength calibration, and standard stars are compared to get the final flux calibration. While the above set up is mostly sufficient to classify a new event, the subsequent follow up has been obtained for wider spectral coverage using a different setting of filter type, depending upon the object’s SED. In the next section, I give a brief overview of SN programs conducted with ROTSE III telescopes.

4.2.4. Overview of supernova search with ROTSE

4.2.4.1. Texas Supernova Search - University of Texas

The first use of the ROTSE telescope for SNe was done by the Texas Supernova Search (TSS) program [129] starting from late 2004. This involved discovery and follow up of SNe with the ROTSE IIIb telescope, and scheduled spectroscopic follow up by the nearby Hobby-Eberly Telescope (HET). TSS discovered about 30 SNe in the first two years. They were able to build a well observed nearby SNe sample, occasionally with spectroscopy at high cadence. The primary objective was to catch the SNe at earliest phases, where the sample was very low. This is still the case today. One of the important and exciting findings of the TSS program was the discovery of Superluminous Supernova (SLSNe) [130].

The search fields were selected primarily from the galaxy clusters: Coma, Virgo, Ursa Major, Abell, and Perseus. The frequency of observations varied with galaxy number density

\(^5\)http://iraf.noao.edu/

\(^6\)http://www.harrisgeospatial.com/SoftwareTechnology/IDL.aspx
and season and the shared observing time between the University of Texas and the University of Michigan. When optimal, a few selected nearby galaxy samples were also observed. The galactic depth for this selection is about magnitude 18 in $R_b$ and. See Chapter 4 of Ref. [129] for details of TSS search fields.

4.2.4.2. ROTSE Supernova Verification Project - University of Michigan

After the TSS program, the ROTSE Supernova Verification Project (RSVP) program was designed as the successor of TSS, but extended the supernova search with other ROTSE telescopes as well. RSVP used a majority of the target fields from TSS, but also expanded the coverage towards the southern sky because of availability of two telescopes in the southern hemisphere. These fields were overlapped with the Sloan Digital Sky Survey (SDSS) when appropriate. In total there were 323 fields, with highest efficiency of search space-time volume of about 2654 Mpc$^3$ year with IIIb telescope. See Chapter 4 of Ref. [180] for details of RSVP observations. A key difference with TSS from the SN data processing point of view was that the TSS adopted an image differencing software from the Supernova Cosmology Project, while RSVP had to redevelop it. The new RSVP differencing software showed similar performance as TSS software, and was adopted for the online and offline photometric data reduction of SN data.

By September 2009, RSVP found 46 SNe including one of the exceptionally luminous SNe: SN 2007if. Spectroscopic follow ups were continued with the HET telescope as available. The RSVP SNe were discovered about equally with ROTSE IIIb and IIId, and 2 each from IIIc and IIIa. While the GRB responses from all four telescopes have been somewhat uniform, the SNe discovery and follow up imaging is dominated by the IIIb and IIId telescopes in the northern hemisphere. A few of the SNe discovered have occurred in the common fields and thus imaging took place with both telescopes.
**Supernova observation and follow-up:** Beginning late 2012, SMU took on the ownership of the ROTSE IIIb telescope at McDonald Observatory in west Texas, USA. SMU also took the control of the observations and operations of the ROTSE web servers, and monitoring daily images through the web. These facilities, along with analysis codes were available from the previous RSVP and TSS projects. I was first trained to perform these operations and then partner with SMU team members running the IIIb telescope. Beginning early 2013, I took responsibility of operating and maintaining the ROTSE IIIb telescope, but also daily monitoring of the nightly images from ROTSE IIIId as well.

I did not add any fields to the existing target fields, but built up a follow up strategy for the interesting candidates, whether or not they were discovered by the ROTSE telescopes. Particularly bright, young Ia and IIP events were selected soon after their discovery and after astrometric observational constraints were confirmed, they were scheduled for observation. In most cases, the observing cadence was left identical with SN search fields but priorities were tuned depending on the SN age. I continued to name the new SNe follow up fields as *vsp* fields, a ROTSE name given to temporarily observe transient events such as variable stars, nova etc. This scheduling proved very useful to monitor events by increased sampling of light curve when the transient was undergoing rapid dynamical changes. (e.g. [40]). During this initial campaign, I discovered or performed prompt analysis of 10 SNe in the ROTSE IIIb fields, and followed up about 15 others.

**Texas Supernova Spectroscopy Survey:** I also accumulated the SNe spectra from 2003 to 2013, organized and archived into the online database WISEREP ([178]) as a part of the Texas Supernova Spectroscopic Survey ($TS^3$). These spectra mostly include, but are not limited to, the ROTSE discovered events. Spectra were mostly taken with HET, but were scattered within the members of the collaboration over the years. With this work, both published and unpublished spectra are more organized within the database. The previously
published spectra are publicly available from the website https://wiserep.weizmann.ac.il/. This sample includes more than 600 spectra of almost 200 SNe, but this list is not complete.

Hereafter, TS$^3$ refers to the SN observation and analysis after late 2012 that includes both photometry and spectroscopy.

4.3. ROTSE Supernova search at SMU

After a SN candidate is found and the subsequent offline time series analysis finds convincing evidence for a potential discovery, telescope time with HET or otherwise is requested for a spectrum. In most cases, visual inspection of the reduced spectrum is sufficient to confirm a new SN. A prompt analysis is performed for preliminary SN redshift measurement, and classification using SNID ([15]) template fitting code. SNID statistically compares the new spectrum with a library of SN spectra at different epochs and gives out a set of best matching redshift, age and classification with a confidence level for each match. Additionally a set of crude measurements of some of the commonly observed parameters such as expansion velocity, presence of high velocity lines etc. are also publicly announced and encouraged for follow up observations as warranted. Table 4.1 lists all the SNe observed with the ROTSE IIIb telescope in the ROTSE SN search fields since fall 2012. Table 4.2 lists the SN events followed up outside of ROTSE search fields as vsp targets since then.

4.3.1. SNe with search observations and follow ups

SN 2012fb: SN 2012fb was discovered on 17 Sept., 2012 with the ROTSE IIId and ROTSE IIIb telescopes (CBET 3237). The object, located at RA = 01$^h$50$^m$51.23 and Dec = +33°08′25.94 (J2000), was confirmed to be a normal Type Ia event, from the spectrum obtained on Sept 22, 2012, with the HET LRS spectrograph. With abundance of Si II and Fe II features, the event closely matched with normal type Ia- SN 1994D at a few days after maximum, in the SNID ([15]) fit. Redshift obtained from the narrow H-alpha feature is $∼$0.038, which is consistent with the photo - $z$ 0.0405 $±$ 0.0134 of proposed host galaxy SDSS
### Table 4.1. SNe in search observations and follow-up by ROTSE IIIb telescope in 2012-2015

<table>
<thead>
<tr>
<th>SN Name</th>
<th>Disc. Date</th>
<th>RA</th>
<th>DEC</th>
<th>Type</th>
<th>Redshift</th>
<th>ATel/CBET</th>
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<tbody>
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<td>01:50:51.23</td>
<td>+33:08:25.94</td>
<td>Ia</td>
<td>0.0405</td>
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<td>2012-11-20</td>
<td>13:00:36.10</td>
<td>+27:34:24.64</td>
<td>Ia</td>
<td>0.0170</td>
<td>CBET 3319</td>
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<td>SN 2013X</td>
<td>2013-02-06</td>
<td>12:17:15.19</td>
<td>+46:43:35.94</td>
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<td>0.0326</td>
<td>CBET 3413</td>
</tr>
<tr>
<td>SN 2013ag</td>
<td>2013-03-02</td>
<td>12:51:35.02</td>
<td>+26:37:45.41</td>
<td>Ia-HV</td>
<td>0.0212</td>
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<td>2013-04-05</td>
<td>12:36:27.67</td>
<td>+11:45:28.1</td>
<td>Ia</td>
<td>0.0658</td>
<td>ATEL 4965/CBET 3470</td>
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<td>SN 2013bu</td>
<td>2013-04-29</td>
<td>22:37:02.14</td>
<td>+34:24:05.72</td>
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<td>0.0027</td>
<td>ATEL 5005/ CBET 3498</td>
</tr>
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<td>SN 2013dz</td>
<td>2013-07-11</td>
<td>02:58:24.40</td>
<td>+36:17:03.50</td>
<td>In</td>
<td>0.0490</td>
<td>CBET 3589</td>
</tr>
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<td>2013-07-25</td>
<td>01:36:48.16</td>
<td>+15:45:31.00</td>
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<td>0.0022</td>
<td>CBET 3606, 3609</td>
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<td>SN 2013ab</td>
<td>2013-10-30</td>
<td>01:50:46.32</td>
<td>+33:06:35.6</td>
<td>Ia</td>
<td>0.035</td>
<td>CBET 3746</td>
</tr>
<tr>
<td>SN 2014J</td>
<td>2014-01-22</td>
<td>09:55:42.14</td>
<td>+69:40:26.0</td>
<td>Ia</td>
<td>0.0008</td>
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<td>SN 2014L</td>
<td>2014-01-26</td>
<td>12:18:48.68</td>
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<td>0.0080</td>
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<td>SN 2014W</td>
<td>2014-01-25</td>
<td>11:54:42.29 +44:01:18.10</td>
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<td>0.0363</td>
<td>ATEL 5832/ CBET 3819</td>
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</tr>
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<td>SN 2014dt</td>
<td>2015-01-29</td>
<td>12:21:57.71</td>
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<td>Iax</td>
<td>0.0052</td>
<td>CBET 4011</td>
</tr>
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### Table 4.2. SNe followed up outside search fields in 2012-2015

<table>
<thead>
<tr>
<th>SN Name</th>
<th>Disc. Date</th>
<th>RA</th>
<th>DEC</th>
<th>Type</th>
<th>Redshift</th>
<th>ATel/CBET</th>
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<td>+09:53:12.3</td>
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<td>+31:13:38.3</td>
<td>IIb</td>
<td>0.0024</td>
<td>CBET 3557</td>
</tr>
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<td>SN 2013dh</td>
<td>2013-06-12</td>
<td>15:30:01.09</td>
<td>+12:59:12.9</td>
<td>Ia-91T</td>
<td>0.0134</td>
<td>CBET 3561</td>
</tr>
<tr>
<td>SN 2013dy</td>
<td>2013-07-10</td>
<td>22:18:17.60</td>
<td>+40:34:09.6</td>
<td>Ia</td>
<td>0.0039</td>
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</tr>
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<td>0.0142</td>
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<td>MASTER OT J072940+141425</td>
<td>2015-02-11</td>
<td>07:29:40.10</td>
<td>+14:14:25.5</td>
<td>IIP</td>
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<td>SN 2015bq</td>
<td>2015-02-14</td>
<td>12:35:06.37</td>
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<td>Ia-91T</td>
<td>0.0282</td>
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<td>+29:54:52.5</td>
<td>Ia-91T</td>
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<td>II</td>
<td>0.0175</td>
<td>ATEL 7320</td>
</tr>
</tbody>
</table>
J015051.24+330823.0. From the absorption minimum of Si-II 635.5 nm feature, a photospheric velocity of $\sim 12000$ km/s is obtained. ROTSE detected SN 2012fb until about 25 days after maximum. Images at about 40 days past maximum show no detection at limiting magnitude of 17.5.

**SN 2012ha:** SN 2012ha was another Type Ia event, first observed in ROTSE IIIb images on Nov 20, 2012. The discovery was announced after the confirmation with the spectrum on Nov 29, 2012; obtained using HET (CBET 3319). Cross-correlation of the spectrum with SNe spectral templates using SNID confirmed the age $\sim 30$ days after maximum. The object was found at RA = 13$^h$00$^m$36$^s$.10, Dec = +27$^\circ$34$'$24$''$.6 (J2000); at 6$.3$ east and 2$.6$ south from the core of the proposed host galaxy 2MFGC 10318, at $z = 0.01700 \pm 0.00001$ in the Virgo cluster. SN 2012ha was observed with ROTSE until 110 days after maximum, where its magnitude was $\sim 17$.

**SN 2013X:** SN 2013X was discovered by ROTSE IIIb on Feb. 6, 2013 at RA 12$^h$17$^m$15$^s$.19, Dec = 46$^\circ$43$'$39$''$.7 (J2000). The object was located 4$''$.9 east and 3$''$.8 south of the presumed host galaxy SDSS J121714.86+464339.7, which has a spectroscopic redshift of 0.03260 $\pm$ 0.00001 (CBET 3413). A spectrogram obtained on Feb 10 with the 9.2m HET, when fitted over SNID, indicated it to be a 1991T- like Type Ia event about 10 days past maximum. The minimum of the Si II $\lambda 6355$ line, after removing the host galaxy recession, was found to be about 11300 km/s. The last ROTSE detection of SN 2013X was at $\sim 45$ days after maximum at magnitude 19.1.

**SN 2013ag:** SN 2013ag was also discovered with ROTSE IIIb at unfiltered magnitude 16.0 on March 2, 2013. The transient was found at RA = 12$^h$51$^m$35$^s$.02, Dec = +26$^\circ$37$'$45$''$.4 (J2000), off centered by 14$''$.9 in the west and 1$''$.1 in the north of the presumed host galaxy SDSS J125135.41+263744.09, at a redshift of 0.02129 $\pm$ 0.00004 (CBET 3428). Spectrum obtained on March 3, 2013 was cross-correlated with SNID templates, which confirmed SN 2013ag
to be a high velocity type-Ia SN, with the best match with SN 2002bo about 2 days before maximum light. Absorption minimum of Si-II 635.5 nm line showed a blueshift by 12600 km/s. SN 2013ag was observed with ROTSE IIIb until \( \sim 35 \) days after maximum.

**SN 2013be:** SN 2013be was discovered by the THU-NOAC Transient Survey (TNTS)\(^7\) on April 5, 2013 at RA = 12\(^{h}\)36\(^{m}\)27\(^{s}\).67, Dec = +11\(^{\circ}\)45\(^{\prime}\)28\(^{\prime\prime}\).1. (CBET 3470) It was spectroscopically identified as Type Ia with the spectrum taken by the Nearby Supernova Factory II\(^8\) on April 9, 2013. This object was substantially dim for ROTSE IIIb sensitivity, but after the object increased in brightness, it was observed for a few days. A follow up HET spectrum of SN 2013be was obtained on April 8, 2013. The redshift of the host galaxy was found to be 0.0658.

**SN 2013bu:** Although SN 2013bu was not discovered by ROTSE, it was detected by the automated pipeline of ROTSE IIIb at an unfiltered magnitude 15.5 on April 29, 2013; a week after the first public report of its detection. The object is located at RA = 22\(^{h}\)37\(^{m}\)02\(^{s}\).17 and Dec = +34\(^{\circ}\)24\(^{\prime}\)05\(^{\prime\prime}\).2 (J2000). The object was identified to be a young IIP event, located about 25\(''\) west and 52\(''\) south from the center of the host NGC 7331 (CBET 3498, ATel 5005). HET took a nebular spectrum of SN 2013bu on Aug. 3, 2013. SN 2013bu was observed with ROTSE IIIb for about 160 days after explosion.

**SN 2013dz:** SN 2013dz is the only IIn SN, discovered by ROTSE IIIb during the SMU era. The object was first detected on July 11, 2013, at magnitude about 16. It is located at RA = 02\(^{h}\)58\(^{m}\)24\(^{s}\).40, Dec = +36\(^{\circ}\)17\(^{\prime}\)03\(^{\prime\prime}\).5 (J2000); which is 2\(''\)8 west and 1\(''\).5 north of the apparent host galaxy SDSS J025824.58+361701.9 (CBET 3589). A spectrum was obtained on July 13, using HET LRS, which under cross-correlation with SNID templates confirmed

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\(^7\)http://astro.tsinghua.edu.cn/

\(^8\)http://snfactory.lbl.gov
classification as IIn. The spectrum exhibited moderately narrow (300-800 km/s) Balmer emission lines at \( z = 0.0490 \), and a blue continuum. With ROTSE IIIb, I performed continued observation of SN 2013dz for more than 150 days after maximum.

**SN 2013ej:** SN 2013ej was one of the nearest IIP SNe found at RA = 01\(^h\)36\(^m\)48\(^s\).16, Dec = 15\(^0\)45'31".0 in the spiral arm of the galaxy M74. It was originally discovered by KAIT on July 25. From a closer inspection of ROTSE IIIb observations, I found that ROTSE data was 1 hr. 40 min earlier, resulting in a pre-discovery detection (CBET 3606). A substantial photometric and spectroscopic follow up of SN 2013ej was carried out by ROTSE IIIb and HET. A detailed study of SN 2013ej combining data from different telescopes spanning UV through IR is presented in Ref. [40], and Chapter 5 of this thesis. ROTSE IIIb data extend up to 215 days after maximum brightness.

**SN 2013hb:** SN 2013hb was discovered by Catalina-Real Time Transient Survey (CRTS) as SNhunt214 on Oct 10, 2013. It was found at RA = 01\(^h\)50\(^m\)46\(^s\).32, Dec = +30\(^\circ\)06′35″.6 in a crowded field, near IC 1773. Even though it lands in the ROTSE search fields, the online pipeline did not detect it potentially due to lower detection efficiency of the RSVP image subtraction analysis in a crowded field. This is the limitation of the software and a demerit of a large pixel size. The type was confirmed as Type Ia, at a redshift of 0.035 (CBET 3746). SN 2013hb was detected by ROTSE IIIb until \(~50\) days after maximum brightness.

**SN 2014J:** SN 2014J was another very nearby Type Ia SN, found at RA = 09\(^h\)55\(^m\)42\(^s\).14, Dec = +69\(^\circ\)40′26″.0, which is 54″ west and 21″ south of the center of the cigar galaxy M82. Because of complexity and huge galaxy background, ROTSE online search alerts were not triggered until 1 week after explosion. ROTSE IIIb had the earliest detection in the pre-discovery image of Jan 15 at magnitude 13.5. No detection was observed on Jan 7th at a shallow limiting magnitude of 14.6 (CBET 3792). Low resolution spectra taken on Jan 22 by multiple groups (CBET 3792) confirmed the discovery of Type Ia event, about 1-2 week
before maximum. Highly sensitive imaging from iPTF on Jan 12 at limiting magnitude of 20.4 showed no detection at the SN target. So Jan 15 detection can provide an upper limit to the age by 3 days after explosion.

**SN 2014W:** Using ROTSE IIIb data, I discovered SN 2014W at RA = 11^h^54^m^42^s^.29, Dec = 44^0^01^′^18^″.1, in the unfiltered images (mag ~17.5) taken on Jan 25. A spectroscopic confirmation was reported in CBET 3819 with a 500-1000 nm spectrum obtained on Mar 1 using 10m Keck II telescope. The spectral cross correlation using SNID found the best match with Type Ia SN, at 2-3 weeks past maximum brightness. Narrow emission lines from the host confirmed a host recession velocity of ~10,900 km/s.

**SN 2014dt:** SN 2014dt was discovered by Koichi Itagaki on Oct. 29, 2014 at RA = 12^h^21^m^57^s^.57, Dec = +04^0^28^′^18^″.5, which is 34″ east and 7″.2 south from the center of the host galaxy M61. This field was not observed by ROTSE IIIb until Jan. 21. The object was detected by the online pipeline on Jan 29. So I scheduled a follow up *usp* imaging from early Feb., 2015. The event was spectroscopically classified on Oct 31, 2014 as a peculiar Iax type event, about a week after maximum brightness (ATel 6639). ROTSE IIIb observed SN 2014dt for about next 100 days.

4.3.2. SNe with follow up cadence data from ROTSE IIIb

In the *TS^3* program, I started a follow up strategy of interesting SNe that would not be occuring in any of the ROTSE SN search fields, but were otherwise observable. For these events, I observed sufficiently late until the SN optical signal was below the ROTSE sensitivity, so as to construct a template image, needed for subtraction of the background host light.

**SN 2013ab:** SN 2013ab was discovered by LOSS ([54], CBET 3422) on Feb 17, 2013 at RA = 14^h^32^m^44^s^.49, Dec = +09^0^53^′^12^″.3, about 15″ east and 2″ south of the center of the
Figure 4.3. ROTSE light curve and a representative HET spectrum for Top: SN 2012fb, shown spectral epoch is labelled as $S$, and limiting magnitudes for the nights when object was not detected, are shown by downward arrows on the left panel (here and after) Middle: SN 2012ha and Bottom: SN 2013X. Light curves are shown relative to approximate maximum derived from SNID or otherwise.
Figure 4.4. ROTSE light curve and a representative HET spectrum for Top: SN 2013ag, Middle: SN 2013be and Bottom: SN 2013bu. Light curves are shown relative to approximate maximum derived from SNID or otherwise.
Figure 4.5. ROTSE light curve and a representative HET spectrum for Top: SN 2013dz, Middle: SN 2013ej and Bottom: SN 2013hb. Light curves are shown relative to approximate maximum derived from SNID or otherwise.
Figure 4.6. ROTSE light curve for Top: SN 2014J, Middle: SN 2014W and Bottom: SN 2014dt. Light curves are shown relative to approximate maximum derived from SNID or otherwise. Spectrum shown for SN 2014W was observed with DEIMOS spectrograph on Keck II telescope.
host galaxy NGC 5669. A spectrum taken on Feb 22, 2013 with the HET LRS confirmed the object to be IIP type (ATel 4839). I triggered ROTSE IIIb follow up of SN 2013ej on Feb 22, after the spectral classification. A template image was constructed taking 17 images from May 2014. ROTSE IIIb detected SN 2013ab for about 70 days after the discovery.

**SN 2013df:** SN 2013df was discovered on June 7, 2013 by ISSP$^9$ (CBET 3557) at RA = 12$^h$26$^m$29$^s$.33, Dec = +31$^\circ$13$'$38$''$.3, hosted in nearby galaxy NGC 4414. It was spectroscopically identified as a IIb type SN. Following the discovery and observed increasing brightness, I scheduled SN 2013df a targeted *v*sp followup for ROTSE IIIb. Follow up observations of Type IIb SN 2013df was done for about 200 days since explosion ([152]).

**SN 2013dh:** SN 2013dh was discovered on June 12, 2013 by LOSS at RA = 15$^h$30$^m$01$^s$.09, Dec = +12$^\circ$59$'$12$''$.9. It was found about 3$''$.7 east and 8$''$.6 south from the core of the host NGC 5936. A total of 5 spectroscopic follow up observations of the object were obtained with the HET LRS spectrograph. ROTSE IIIb detected the event until about 80 days after the explosion.

**SN 2013dy:** SN 2013dy was also discovered by LOSS on July 10, 2013 at RA = 22$^h$18$^m$17$^s$.60, Dec = 40$^\circ$34$'$09$''$.6. The object was found to be 2$''$.1 west and 24$''$.9 north of the host galaxy NGC 7250. SN 2013dy was discovered very early in the light curve phase and spectroscopically confirmed as Ia (CBET 3588). It was one of the brightest event of 2013 reaching a peak magnitude of about 12.7. ROTSE IIIb observed SN 2013dy for nearly 200 days after peak. A set of 7 HET LRS spectra were obtained for the event within -3 to +15d of $B^\prime$ max and published in Ref. [116].

Using the image differencing technique adopted from RSVP, I observed potential areas of improvement in the detection efficiency and reduction of photometric residuals. In 2014,

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$^9$italiansupernovae.org
Figure 4.7. ROTSE light curve and a representative HET spectrum for Top: SN 2013ab, Middle: SN 2013df and Bottom: SN 2013dh. Light curves are shown relative to approximate maximum derived from SNID or otherwise.
I started writing a new image differencing code and, over time, have analyzed tens of SN photometry that have shown improvements in both the detection efficiency and the final light curves. As of now, the new code has only been used for offline analysis and the automated pipeline still uses the original differencing code from RSVP. In Section 6.1, I will explain in detail the algorithmic prescription of the new image differencing code, and present some performance results from the new code. All the SN light curves shown in this chapter were reduced using the new code.
In this chapter, I present photometric and spectroscopic analyses of three individual SNe observed by ROTSE IIIb. These analyses were performed to understand the astrophysical properties of a variety of SNe at an individual and a class level. SN 2013ej is studied thoroughly from very early to very late times while SN 2012cg and SN 2013df are limited to photometric analysis addressing very early and very late time behaviour of the two events respectively.

5.1. Extensive spectroscopic and photometric analysis of Type IIP SN 2013ej

SN 2013ej was one of the brightest and the nearest SNe of 2013. Because it was a IIP event, it remained bright for about 100 days after explosion. This provided an excellent opportunity to accumulate data from many different telescopes over a wide range of energy. With the wealth of photometric and spectroscopic data of SN 2013ej, I present an extensive analysis delivering a variety of measurements from the SN. This analysis was motivated towards a thorough understanding of IIP properties and as a preparation for a follow up cosmological study using these events. I also combined other well sampled IIP objects from the literature to measure the systematic dispersion of their properties and seek to establish a better calibration method for cosmological study to follow.

I combined photometric data sets of SN 2013ej from several telescopes from ground and space, spanning unfiltered, $UBVRIJHK$ and $ugriz$ filters. Optical and UV spectra were obtained from four different telescopes. With such densely sampled data, I was also motivated towards establishing an empirical bolometric calibration from ROTSE data. An accurate calibration technique would facilitate better measurements of the properties of similar events that have sparse data sample. Thus a high quality dense data set not only helps...
understand the IIP explosion mechanism better but also aids to understand the systematics as they are used as cosmological probes (e.g., \[34, 40, 61, 69, 112, 128\]).

5.1.1. Discovery, classification and early observation

SN 2013ej was discovered on 2013 July 25, with the 0.76-m Katzman Automatic Imaging Telescope (KAIT) at Lick Observatory in the nearby spiral galaxy M74 ([87]). As a brightening transient source in such a nearby galaxy, SN 2013ej captured a lot of attention as one of the closest and earliest discovered SNe ever. Further constraints on its age were provided by the prediscovery photometry announced with the Lulin telescope ([92]) and the ROTSE IIIb telescope at McDonald Observatory ([41]). Spectra obtained with the Hobby Eberly Telescope (HET), and the KAST spectrograph at Lick Observatory clearly indicated a young core-collapse event, showing high statistical confidence in a SN IIP classification when compared with a series of template spectra using SNID. Photometry in \(BVR\) bands from July 24.8, which is 15 hr earlier than the first ROTSE IIIb detection, and the youngest observed phase, was announced in Ref. [92]. Furthermore, a nonphotometric report of a prediscovery observation on images from July 24.125 was also announced by C. Feliciano on the Bright Supernovae website\(^1\). Imaging on 2013 July 23.54 by ASAS-SN showed no emission at \(V = 16.7\) limiting magnitude ([141]).

Analysis of the first month of photometry and spectroscopy was published by Ref. [165], who observed a rather slowly evolving SNe IIP. A moderately strong feature blueward of H\(\alpha\) was a somewhat peculiar observation and was suggested to be a Si II line. Pre-explosion images obtained with the \textit{Hubble Space Telescope (HST)} were analysed by Ref. [59], from which they suggested two possible progenitors, favoring the redder source to be the more likely candidate. Their analysis put a progenitor \textit{ZAMS} mass constraint of 8–15.5 M\(_\odot\). Later on, hydrodynamic modeling was performed by Ref. [79], which also supported the progenitor to be a red supergiant of mass of 12–13 M\(_\odot\) before explosion. From the observed steep plateau in the optical bands and the high velocity of strong H I lines, Ref. [19] favored

\(^1\)http://www.rochesterastronomy.org/supernova.html
SN 2013ej as a IIL event. But the IIP/IIL classification boundary has been debated as discussed in Chapter 2.

In the following sections, I will present an in-depth photometric and spectroscopic analysis of SN 2013ej, followed by its distance estimation, and measurement of kinematic parameters from these observations. With an extensive data set, I will present a wealth of measurement of physical parameters from the explosion. The dates quoted hearafter are relative to the derived explosion epoch MJD 56496.9, unless otherwise explicitly mentioned.

5.1.2. Observation and data reduction

5.1.2.1. Photometry

The first automated detection of SN 2013ej with the ROTSE IIIb telescope was only made on 2013 July 31.36. The July 14.42 precursor image rules out any emission at a relatively shallow limiting magnitude of 16.8. The first good-quality observation of the field took place on July 25, and with a more careful offline analysis of the observations, I obtained the earliest $> 5\sigma$ detection at July 25.38, which was 100 minutes before the announced discovery epoch Ref. [41] from KAIT. For SN 2013ej, I scheduled parallel $v_{sp}$ follow-up observations with ROTSE IIIb to achieve more densely sampled photometry. This follow-up continued for more than 200 days as shown in Fig. 5.2, at which point SN 2013ej passed beyond ROTSE IIIb detection.

The ROTSE data were initially reduced using an image-reduction pipeline Ref. [181], followed by a DAOPHOT-based point-spread-function (PSF) photometry technique Ref. [149]. In the first attempt, I observed significant photometric artifacts and reduced efficiency using the image differencing technique. Thus, an aperture photometry of SN 2013ej (e.g., [98]) was performed as the SN occurred in the outskirts of host galaxy M74. I chose an aperture size of 1 FWHM of the median PSF on each image for the signal region, and an annulus of inner and outer radii of 2 and 4.5 times the FWHM was considered for measuring the background sky. To correct for any host contamination, a reference template image was
smeared to reflect the median PSF of each epoch, and the underlying host contribution inside the aperture was subtracted. For the observed timescale, the ROTSE IIIb PSF was at 3–4″. The final photometry was obtained by calibrating the derived relative flux to USNO B1.0 R band catalog.

Broadband photometry from +8 d to +130 d was obtained with the 60/90 cm Schmidt telescope of the Konkoly Observatory at Piszkesteto Mountain Station, Hungary, in the Bessell BVRI filters. Fig. 5.1 shows a BVI color combined field around SN 2013ej. Observations were also done in Sloan g′r′i′z′ filters at the Baja Observatory, Hungary, with the 50 cm BART telescope equipped with an Apogee-Ultra CCD.

Image Reduction and Analysis Facility (IRAF) routines were used to reduce the photometry of the Konkoly data. Obtained instrumental magnitudes were calibrated to the standard Johnson-Cousins system via local standards stars shown in Fig. 5.1 tied to Ref. [90] standards on a photometric night. The g′r′i′z′ data obtained from Baja Observatory were calibrated picking ~ 100 stars passing a magnitude cut 14 < r′ < 18 within the ~ 40 × 40 arcmin² field of view around the SN, taken from the Sloan Digital Sky Survey (SDSS) DR12 catalog.

Observations with multi-channel Reionization And Transients InfraRed camera ([24, RATIR;]) mounted on the 1.5 m Johnson telescope at the Mexican Observatorio Astronómico Nacional on Sierra San Pedro Mártil in Baja California, México ([173]) were also obtained for SN 2013ej. A series of 80 s exposures in the ri bands and 60 s exposures in the ZYJH bands spanned from +3 d to +125 d. Coaddition, reduction and calibration were done with standard CCD and IR processing techniques in IDL and Python, and using SExtractor and SWarp. Final calibration was done against field stars with reported fluxes in both 2MASS ([146]) and the SDSS DR9 Catalog ([1]).


Along with unfiltered CCD data extending to +213 d, an extensive $BVRI$ imaging was done by the KAIT and the Nickel 1-meter telescope located at Lick Observatory starting from July 30. The broadband data extends long in the tail up to +461 d. (see Fig. 5.2 and Fig. 5.15). KAIT and Nickel data were reduced using PSF photometry ([62]) using DAOPHOT. Instrumental magnitudes were transformed to the Landolt system$^4$ and calibration was done against field stars in the APASS$^5$ catalog. Unfiltered photometry was calibrated to APASS $R$ band magnitudes. As the object was extremely bright and far off the host core, no image subtraction was performed on the KAIT data.

The $UVOT$ instrument onboard the NASA $Swift$ space telescope also observed SN 2013ej through the $uvw2$, $uvm2$, $uvw1$, $u$, $b$, $v$ filters from +7d to +138d. This dataset was obtained

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$^4$http://www.sdss.org/dr7/algorithms/sdssUBVRITransform.html  
#Lupton2005

$^5$http://www.aavso.org/apass
from the *Swift* archive\(^6\). Also see Ref. [165], [79], [19] for independent usage of *Swift* data. This reduction yielded brighter magnitudes than [79] beyond +30 d in the \(uvw_2\), \(uvm_2\) and \(uvw_1\) filters. Noting that \(uvw_2\) and \(uvw_1\) filters have extended red tails (e.g., [48]) and their fluxes have a marginal contribution to the total flux after +30 d ([40]), these bands were ignored after +30d (see Section 5.1.3.4). All photometric data are published in Ref. [40]. Here, the light curves are plotted in Fig. 5.2.

**5.1.2.2. Spectroscopy**

A time series of 17 low-resolution optical spectra spanning from +8 d to +135 d were obtained for SN 2013ej. These were obtained using the Marcario Low-Resolution Spectrograph (LRS; [73]) on the 9.2 m Hobby-Eberly Telescope (HET) at McDonald Observatory, the dual-arm Kast spectrograph ([103]) on the Lick 3 m Shane telescope, and the DEep Imaging Multi-Object Spectrograph (DEIMOS; [49]) on the Keck-II 10 m telescope.

Standard CCD processing and reduction techniques (e.g., [144]) were used and the final data were extracted with the optimal algorithm of Ref. [76]. Wavelength calibration was performed from low-order polynomial fits to calibration-lamp spectra. Further fine tuning for small wavelength shifts were done by cross-correlating a template sky to the night-sky lines extracted along with the SN spectra. Spectrophotometric standard-star spectra were used against the data in order to flux calibrate the spectra and to remove telluric lines ([100,171]). Table 5.1 provides a log of the obtained spectra and Fig. 5.3 plots them. All of these spectra are archived on WISEREP and are publicly available from the database.

UVOT/UGRISM onboard *Swift* took near-UV spectra of SN 2013ej in the wavelength range 2000–5000 Å in between +8 d and +16 d. These early time UV spectra were collected from the *Swift* data archive, and then reduced using the *uvotimgrism* task in *HEAsoft*\(^7\). The UGRISM spectral observations are logged in Table 5.2 and are shown in Fig. 5.11.

\(^6\)http://heasarc.gsfc.nasa.gov/cgi-bin/W3Browse/swift.pl

\(^7\)http://heasarc.nasa.gov/lheasoft/
Figure 5.2. Left: Open CCD and broadband photometric light curves of SN 2013ej. BVRI and unfiltered data from KAIT are shown by empty circles, and Konkoly BVRI and ROTSE observations are shown in solid circles. Swift data points are plotted in filled squares. Right: SDSS $g'r'i'z'$ photometry from Baja Observatory and RATIR optical and near-IR photometry for SN 2013ej.
Table 5.1. Observing log of SN 2013ej optical spectra ([40])

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### Table 5.2. Observing log of *Swift* UVOT/UGRISM spectra for SN 2013ej

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### Figure 5.3

**Left:** Time series optical spectral data of SN 2013ej. Phases quoted are in days relative to explosion time (MJD 56496.9). The observing log is shown in Table 5.1. **Right:** The early time near UV spectra SN 2013ej obtained with *Swift*/UGRISM. Feature identifications are based on Ref. [23].
5.1.3. Photometric analysis

As narrow Na I D lines from the host galaxy are commonly used to derive the extinction correction, these lines were not observed in SN 2013ej spectra in this sample or published spectra (e.g., [19]). Ref. [165] showed negligible reddening from M74 in the direction of SN 2013ej. Therefore, I also did not apply any reddening correction from the host extinction. However, the Milky Way reddening value of $E(B-V) = 0.061$ mag was used following Ref. [138] and considered the total extinction.

5.1.3.1. Constraining shock breakout

After the core collapses in a IIP SN, the gravitational waves and neutrinos are believed to escape soon, while the electromagnetic signal remains trapped in the envelope. The time scale for the hydrodynamic shock front to reach the photosphere can be from hours to days. Only then, the photon emission from the explosion can be observed. This epoch begins the shock breakout phase and marks the time of first light. Tight constraints of shock breakout phase can provide precise knowledge of the kinematics before and after the explosion. This is also important to reduce the systematic uncertainty on the distance measurements through methods such as EPM (see Section 2.3).

Multiple studies have attempted to model the shock breakout of the compact progenitor of SN 1987A (e.g., [43]). Ref. [50] have shown that the breakout peak varies with the envelope mass and density structure. Thus very early light may provide clues on the envelope structure of massive stars. More recent theoretical studies (e.g., [110, 151]) have attempted to understand the shock breakout of SNe II in a variety of progenitor scenarios. Ref. [32] have argued that aspherical shocks can alter the times when the shock breaks, which can result in a very different outcome than a spherical shock from a spherical star.

To estimate the shock breakout, I use multiple datasets during the first few days after the discovery. This required a separate calibration as the earliest data points are still sparse and not concurrent. ROTSE and KAIT unfiltered data are both calibrated to $R$ magnitudes. These are first combined with the earliest prediscovery detection in $R$ band from Lulin
Observatory. There can still be a systematic offset as the response functions of unfiltered
CCD are quite different from $R$ filter. The differences are evident even in the KAIT and
Konkoly $R$ band photometry of SN 2013ej. Assuming the offsets of Lulin $R$ to be at the same
level as KAIT and Konkoly $R$ band data, I calculated the average of differences of KAIT $R$
and Konkoly $R$ magnitudes with ROTSE unfiltered magnitudes between $+30$ d to $+90$ d,
and added this as a correction to the Lulin data point as an effective correction to bring it
to the ROTSE system. The root-mean square (RMS) of KAIT $R$ and Konkoly $R$ magnitude
differences was added as an additional systematic uncertainty to the Lulin data. Note that
the Lulin data point already has the statistical uncertainty higher than 0.2 mag. Further
potential systematic uncertainty arising from translating the calibration from plateau to
early time is expected to be smaller than this. Note that Ref. [25] found 0.1 mag correction
in their study of Gamma ray bursts using KAIT unfiltered and Lulin $R$ band magnitudes.
Both ROTSE unfiltered and KAIT unfiltered data track $R$ band magnitudes quite well and
they have been cross-calibrated to the same unfiltered source in the past in independent early
time studies of SN Ia and SN IIP (e.g., [132, 183]). These results are convincing evidence
that above calibration is within the statistical consistency of those studies.

For the first night of ROTSE detection, the observations had a time granularity of 2 hr
between coadds of two sets of images. So instead of coadding the full night images, as the
SN was still young ($\sim 1$ d after explosion), I left the coadds separate as a significant rise
could be detected even in only 2 hrs. With the combined data, I performed a least square
fit of the observed fluxes with a simple power law model

$$F(t) = A (t - t_0)^\beta,$$

where $A$ is a constant, $t_0$ is the time of shock breakout, and $\beta$ is the power-law index.

Any analytic functional form for the IIP evolution has not been well established, as the
rise is very short lived and it is hard to get a well sampled light curve that early. A $t^2$ rise
model has been tested many times for SN Ia (e.g. [98, 183]) taking data up to $\sim 2$ weeks
after the explosion. There, it is more reasonable to expect a steady temperature as the heat
loss from cooling is compensated by radioactive heating. In the case of IIP, however, the
adiabatic cooling is expected to significantly drop the temperature. A $t^2$ model fit on very
early IIP data have been done before (e.g., [132]), but there is significant uncertainty on this
approach. Noting that the data and limits are very tight for SN 2013ej, I perform a fit using
the power law model of Eq. 5.1 on data $< +2$ d. Fig. 5.4 shows the fits obtained where the
data points are relative to July 24.8, the Lulin observation epoch. First the power index $\beta$
was kept free, which yielded $\beta = 4.83$, and $t_0 = -2.19$ with $\chi^2/dof = 2.80$. Next $\beta$ was kept
fixed at value 2, which gave $t_0 = -0.90 \pm 0.25$ days with better $\chi^2/dof$ value of 1.47. This
fitted value of $t_0$ with fixed $\beta$ corresponds to July 23.9 $\pm$ 0.25. Deviation of $\beta$ from 2 might
indicate asymmetry in the explosion itself or something else which can not be settled with a
simple model. For SN 2013ej, the power index of 4.83 can not be ruled out with the available
data (see Fig. 5.4), $t^2$ model was favored from better $\chi^2/dof$ and observed consistency with
all reported early detections and limits. MJD 56496.9 $\pm$ 0.3 was thus taken as the estimated
epoch of shock breakout.

5.1.3.2. Unfiltered and broadband photometry

Like other SNe IIP, SN 2013ej light curves show distinctive signatures of the phase evo-
lution in all the optical bands. Because the timescale of photon diffusion is much longer
than the expansion time of relativistic ejecta, a rapid cooling occurs from the outside. As
the temperature subsides, H-ions start to recombine with electrons, giving rise to an inward
receding ionisation wave. The outcome is a linear, slowly declining plateau. The photosphere
is believed to be contiguous with the ionisation front. In about 100 days, all the hydrogen
fully recombines and photosphere further recedes into the inner denser core. This is the
point where light curve suddenly falls off, and is powered mostly by the radioactive decay of
the synthesized material.

The unfiltered and broadband light curves of SN 2013ej are shown in Fig. 5.2. The two
$BVRI$ sets start from around the peak, extend through the characteristic plateau phase for
about 100 days, and fall off to the radioactive tail phase. The peak for SN 2013ej occurs
Figure 5.4. Early rise of SN 2013ej light curve. Multiple datasets are calibrated to ROTSE magnitudes. Solid lines represent power law \((Eq. \ 5.1)\) fits performed on data \(< +2 \, \text{d}\). The triangle is the upper limit from July 23.54 for \(V \approx 16.7\) mag. The dashed line is the detection on July 24.125 but with no photometry. The inset shows the projection of the emission, given a floating (green) and a fixed (blue) value of index \(\beta\).
at +18 d in ROTSE light curve. This corresponds to an absolute peak magnitude of −17.5 (see Section 5.1.4.5 for distance). The peaks occur on +12.5 d in $B$, +15.5 d in $V$, +19.5 d in $R$, and +20 d in $I$ in both Konkoly and KAIT broadband lightcurves. A secondary peak was observed by Ref. [18] in SN IIP 2012aw at about +50 d in the $V$ band (local minimum around +42 d). This is suggested as indication of the end of free adiabatic cooling. No such feature is observed in SN 2013ej. The exact epoch of the advent of the plateau phase is thus uncertain. The light curves decline by 0.038, 0.021, 0.016, and 0.012 mag per day in $B$, $V$, $R$, and $I$ bands from their respective peaks until +90 d. Our $V$ band slope is steeper compared to the 0.017 mag day$^{-1}$ given by Ref. [19] for SN 2013ej. We suspect this to be a sampling issue in a sparsely sampled light curve with a less well constrained peak time.

When compared to an ensemble of SNe IIP (see Fig. 5.5), SN 2013ej shows one of the steepest plateaus. Ref. [18] showed For SN 2012aw, a decline by 1.74 mag by +104 d in the $B$ band and no significant change of brightness in the $R$ band was observed by Ref. [18]. Similar evolution for classic SN IIP 1999em during the plateau was observed by Ref. [93]. A more energetic event, SN 2004et showed faster decline of $\sim 2.2$ mag by +100 d ([18]). For SN 2013ej, the decline rate is systematically greater in all bands. For comparison, I show light curves of a recent IIL type event SN 2013by from Ref. [166] and the archetype SN 1980K from Ref. [8]. The fall of the $V$ band light curve by $\sim 0.75$ mag in +50d puts SN 2013ej in the SN IIL category of Ref. [51], where the classification cut is 0.5 mag. A fall of $1.46 \pm 0.06$ mag in the $V$ band was observed for IIL SN 2013by by Ref. [166], but they also pointed out the SN IIP-like behavior in its plateau-tail drop. A handful of events previously identified as IIP were shown to have $V$ band peak to +50 d drop higher than 0.5 mag; which would put them in the SN IIL category of Ref. [51].

In spite of a relatively steep plateau, SN 2013ej shows a significant sudden drop at the plateau to tail transition. This level is not a signature of any IIL like event, so I treat this as a IIP event. Hydrodynamic models (e.g [11]) show that a steep plateau signifies less extensive mixing from $^{56}$Ni. Higher Ni yield causes more radioactive heating, thus causing early flattening in the plateau. The steep plateau of SN 2013ej may suggest low Ni production.
or the ejecta may have undergone an inefficient thermalization.

The tail phase puts a direct constraint on the radioactive material production. Section 5.1.4.6 presents the measurement of Ni produced. Here I present the behaviour of evolution in different photometric bands. Data from +120 d to +183 d and from +183 d to +461 d are separately fitted to a line. +183 d was chosen as the break time of the observed late time behavior (see Section 5.1.4.6). The decay rate and the respective $\chi^2/dof$ of the fits are presented in Table 5.3. The light-curves decline more steeply in all bands before +183 d. A slope shallower than $^{56}\text{Co} \rightarrow ^{56}\text{Fe}$ decay rate is observed only in the $B$ band after +183d. Note that SN 2006bp had a tail decline rate of $0.73 \pm 0.04$ mag per 100 d ([132]), which is smaller than the radioactive decay considering full trapping of gamma rays. Contrary to this, SN 2013ej exhibits the opposite behavior.

5.1.3.3. Color and temperature evolution

Like other SNe IIP, SN 2013ej exhibits a rapid color evolution in the blue in the first 30 days, as shown in Fig. 5.6. However, the $V - R$ and $V - I$ colors are smooth and slowly rising as this range show slower evolution of spectral energy distribution (SED). As the temperature falls below 6000 K (see below), the evolution trends are more alike. Fig. 5.6 shows the color evolution in the UV through NIR sets on the top panel. The post plateau $V - R$ and $V - I$ colors show a rapid rise, because of greater decline of flux in $V$ than in the $I$ band. This transition is however evident in both the optical and NIR colors. The SED evolution is overplotted with contemporaneous UV and optical spectra in the bottom left panel of Fig. 5.6. Both color and SED evolution ascertain the general characteristic of SNe IIP: a rapid decline of the UV flux ensued by decrease of the continuum slope during the plateau in the optical, also observed as continuous reddening of the optical colors.

Temperature evolution is derived using blackbody fits of the $BVI$ SEDs from the KAIT and Konkoly photometry. $R$ band is excluded to avoid potential contamination from strong H$\alpha$ line. UV bands are also not included as they heavily bias the fits from blending of metallic lines at shorter wavelengths. For SN 2013ej, the temperature drops from 12,500 K
Figure 5.5. $V$ band light curves of IIP and IIL SNe, except for SN 2006bp (ROTS E). SN 2013ej shows a systematically steeper plateau, ensued by a sharp drop around +100d. SNe IIL 2013by and 1980K show linear decline post peak, but the drop is not as sharp at the late time. Also, SN 2013ej exhibits a steeper light curve decline in the tail.
Table 5.3. Radioactive decline rate of SN 2013ej ([40]). $\chi^2$/dof are given inside paranthesis.

<table>
<thead>
<tr>
<th>Band</th>
<th>+120 d to +183 d (mag/100 days)</th>
<th>+183 d to +461 d (mag/100 days)</th>
</tr>
</thead>
<tbody>
<tr>
<td>KAIT $B$</td>
<td>1.15 ± 0.08 (0.37)</td>
<td>0.75 ± 0.02 (1.75)</td>
</tr>
<tr>
<td>KAIT $V$</td>
<td>1.46 ± 0.04 (1.01)</td>
<td>1.10 ± 0.02 (2.69)</td>
</tr>
<tr>
<td>KAIT $R$</td>
<td>1.54 ± 0.04 (2.66)</td>
<td>1.30 ± 0.01 (3.92)</td>
</tr>
<tr>
<td>KAIT $I$</td>
<td>1.63 ± 0.03 (0.78)</td>
<td>1.18 ± 0.02 (2.96)</td>
</tr>
<tr>
<td>KAIT unfiltered</td>
<td>1.51 ± 0.02 (0.88)</td>
<td></td>
</tr>
<tr>
<td>ROTSE unfiltered</td>
<td>1.64 ± 0.07 (1.13)</td>
<td></td>
</tr>
<tr>
<td>KAIT $BVRI$</td>
<td>1.36 ± 0.09 (0.48)</td>
<td>1.06 ± 0.02 (2.15)</td>
</tr>
</tbody>
</table>

at +8 d to 6400 K at +24 d. This steep drop encapsulates rapid adiabatic cooling during the first few weeks. After this, temperature declines more slowly to 4000 K at +100 d as shown in the bottom right plot of Fig. 5.6. This smooth slow decline reflects the nearly steady temperature cooling from diffusion of photon energy during recombination as dictated by the atomic physics.

5.1.3.4. Bolometry

Bolometric flux from the full integration of fluxes over all wavelengths allows one to put strong constraints on explosion parameters. While UV flux is a small fraction for stripped core SNe Ia and SNe-Ib/c, it is dominant in the SNe-II at early times. Emission flashes are expected after the shock breaks out in the UV and X-ray bands. However, after a few weeks of rapid SED evolution, the UV fraction is low and the light output is dominated by the optical emission.

A full bolometric luminosity is generally impossible to obtain in practice. Availability of extensive data from UV through NIR for SN 2013ej provides unique opportunity to obtain the most accurate bolometric flux. In this particular case, where there are observations available
Figure 5.6. Top: Evolution of the optical and near-IR colors of SN 2013ej. The RATIR $J - H$ and $Y - Z$ colors have been converted to the Vega system for consistency. Bottom left: Spectral energy distribution (SED) evolution of SN 2013ej. Broadband photometry from the UV, optical, and near-IR are plotted with filled circles. The horizontal bars represent the FWHM of respective pass band filters. Overplotted are the spectra at similar epochs. Bottom right: Color temperature evolution derived from the Planck function fits of the $BVI$ fluxes using KAIT (green points) and Konkoly (blue points) datasets. Estimates from Ref. [165] are also shown for comparison in red points. For all the measurements, a dereddening with $E(B - V) = 0.061$ mag is applied.
from unfiltered CCD sensitive to UV through NIR, there is an additional benefit to establish a bolometric calibration relation for the class of SNe. I have used the distance of $9.0^{+0.4}_{-0.6}$ Mpc (see Section 5.1.4.5) to derive fluxes in SN frame. In practice, a pseudo-bolometric estimate from limited broad band coverage (e.g., $BVRI$, $UBVRI$ etc) are commonly derived. For the context here, bolometric should be understood as $Swift$ $uvw2$ through $K$ band integrated flux. Any flux outside this range is expected to be at the subpercent level contribution for the epochs considered (see below).

The left panel of Fig. 5.7 shows derived pseudo-bolometric and bolometric light curves of SN 2013ej. The right panel shows the fractional contribution from the UV, optical and NIR bands to the estimated bolometric flux. For calculating bolometric flux, any contribution from the $uvw2$, $uvm2$ and $uvw1$ bands beyond +30 d are estimated at $< 1\%$ and ignored here because of complication of red leaks pointed earlier. For the UV and NIR fraction in the late time, where observations are only available for the optical bands, I use a scaling relation derived in the earlier tail times between +120 d and +137 d.

Because the open CCD transmission is broad spanning the whole optical window, I establish a calibration relation for the ROTSE and KAIT unfiltered photometry with integrated $BVRI$ flux. First, the $BVRI$ magnitudes are converted to absolute flux using the relations given by Ref. [12] corresponding to an A0 star. The value of $\log_{10}(L_{\text{ROTSE}}/L_{BVRI})$ shows a strong correlation with the $B-V$ color. A linear regression model fit of $\log_{10}(L_{\text{ROTSE}}/L_{BVRI})$ versus $B-V$ is shown in the top panel of Figure 5.8 and the residuals in the bottom panel yield a precision better than 5%; while a temporal relation there yields about 8% precision. For KAIT, this analysis yields 4% and 6% precision respectively. Note that for ROTSE calibration, I used the Konkoly $BVRI$ set, and for KAIT unfiltered calibration, I used KAIT $BVRI$ set. Summary of the fit results for each case is given in Table 5.4.

As the open CCD has non-negligible quantum efficiency in the UV, I next try the same procedure for pseudo bolometric $UBVRI$ calibration for both ROTSE and KAIT unfiltered photometry. Here the Swift $u$ magnitudes are first converted to Johnson-Cousins $U$ band following Ref. [127] and integrated to $BVRI$ sets. The integration limits are [3285 Å, 8750 Å].
Figure 5.7. Left: Estimated pseudo-bolometric $BVRI$, $UBVRI$, $UBVRIJHK$, and bolometric light curves of SN 2013ej. The bolometric light curve also includes contribution from UV flux $<\sim 3200$ Å before +30d. In the plateau, $UBVRIJHK$ closely yields the bolometric light curve. Right: Fractional contribution from UV, optical, and near-IR to the bolometric flux. The UV covers flux lower to $U$ band, the near-IR (NIR) fraction is the flux beyond the $I$ band, and the optical fraction covers flux in between ($UBVRI$). The UV fraction is avoided after the plateau, addressing the marginal contribution and potential contamination from red leaks.

Table 5.4. $B - V$ Dependent Pseudo-Bolometric $BVRI$ and $UBVRI$ Calibration of ROTSE and KAIT Unfiltered Data

<table>
<thead>
<tr>
<th>Calibration</th>
<th>Intercept</th>
<th>Slope</th>
<th>$\chi^2$/dof</th>
</tr>
</thead>
<tbody>
<tr>
<td>ROTSE Unf. – Konkoly $BVRI$</td>
<td>$0.362 \pm 0.004$</td>
<td>$0.1004 \pm 0.004$</td>
<td>0.54</td>
</tr>
<tr>
<td>KAIT Unf. – KAIT $BVRI$</td>
<td>$0.401 \pm 0.007$</td>
<td>$0.082 \pm 0.006$</td>
<td>1.58</td>
</tr>
<tr>
<td>ROTSE Unf. – $UBVRI$</td>
<td>$0.299 \pm 0.009$</td>
<td>$0.161 \pm 0.007$</td>
<td>1.05</td>
</tr>
<tr>
<td>KAIT Unf. – $UBVRI$</td>
<td>$0.363 \pm 0.006$</td>
<td>$0.112 \pm 0.006$</td>
<td>2.71</td>
</tr>
</tbody>
</table>
Figure 5.8. Left: Pseudo-bolometric BVRI (Konkoly) calibration of SN 2013ej with ROTSE data. Residuals after applying a $B - V$ color-dependent correction (histogram in blue) show less than 5% scatter. The histogram (green) on the left shows the residuals by subtracting the two fluxes as a function of time. Right: Same as in the left panel, for KAIT unfiltered data to KAIT BVRI calibration. Precision in this case are 6% for temporal comparison and 4% when a $B - V$ dependence is applied.
where the lower and upper limits are chosen are extended by the half width at half-
maximum intensity (HWHM) in the $U$ and $I$ bands. Comparing the temporal evolution of
$\log_{10}(L_{\text{ROTSE}}/L_{UBVRI})$, the RMS of the residuals yield 13% precision. With a $B-V$ de-
pendent correction, the scatter reduces to 6%. Same analysis for KAIT unfiltered to $UBVRI$
(using $Swift$ $u$ and KAIT $BVRI$) yield 5% precision after a color-dependent correction.
The fit results from this analysis are also summarized in Table 5.4.

Because of several observing constraints, SNe can not be concurrently observed over all
wavelengths in UV, optical and IR bands, or even outside this range. SNe of a certain class,
that have the most complete observations provide an opportunity to estimate the missing
fluxes of other similar events within the same class. Here I gather an ensemble of very
well observed SNe-IIP from the literature and establish a bolometric calibration method,
when the data are limited to $BVRI$ or $UBVRI$ bands. Table 5.5 lists the SNe in the
calibration sample, for which each event has extensive data spanning UV through IR. For
the sake of defining terms, integrated flux is represented as $UBVRIJHK$, while the same
thing derived from calibration is defined as “$UtoK$”. The same naming scheme applies to
other permutations. The “bolometric” flux is the integration form $Swift$ $uvw2$ in the blue
to $K$ in the red. Later in the SN evolution, the IR flux past $K$ band can be significant, but
I ignore such contribution in this procedure.

The light curves in Fig. 5.7 are obtained with data in the $UBVRIJHK$ bands inte-
grating in the wavelength range 3285–23,850 Å. For SN 2013ej, the missing $K$ band flux is
estimated from the average fractional flux in $K$ with respect to the $UBVRIJH$ flux using
the calibration sample. $K$ band contributes to $\sim 2\%$ at $+10$ d, rising to $5-6\%$ at $+80$ d.
Sparse $K$ band data is published by Ref. [79]. The estimated $K$ band flux at similar epochs
yields offset of $< 1\%$ of the total bolometric flux in the plateau and agreement is better in
the tail.

The pseudo-bolometric $UBVRI$ flux is significantly lower compared to the estimated
bolometric flux (see Fig. 5.7) with increasing difference over time. The bolometric light
curve declines rapidly by 0.4 dex ($= 10^{0.4}$) in the first 30 days, and relatively slowly by
Table 5.5. SN IIP Bolometric Calibration Sample with Well-Sampled Photometry from $U$ through $K$

<table>
<thead>
<tr>
<th>Object</th>
<th>Host</th>
<th>Distance (Mpc)</th>
<th>Total $E(B-V)$ (mag)</th>
<th>$V$ Plateau Slope (mag/100 days)</th>
<th>Feature</th>
<th>References</th>
</tr>
</thead>
<tbody>
<tr>
<td>SN 1999em</td>
<td>NGC 1637</td>
<td>11.7 ± 1.0</td>
<td>0.10</td>
<td>0.31 ± 0.05</td>
<td>Normal</td>
<td>1,2,3,4</td>
</tr>
<tr>
<td>SN 2004et</td>
<td>NGC 6946</td>
<td>5.6 ± 0.3</td>
<td>0.41</td>
<td>0.72 ± 0.03</td>
<td>Over Luminous</td>
<td>5,6</td>
</tr>
<tr>
<td>SN 2005cs</td>
<td>M51</td>
<td>8.4 ± 0.7</td>
<td>0.05</td>
<td>−0.10 ± 0.05</td>
<td>Subluminous</td>
<td>7,8</td>
</tr>
<tr>
<td>SN 2013ej</td>
<td>M74</td>
<td>9.0$^{+0.4}_{-0.6}$</td>
<td>0.06</td>
<td>1.95 ± 0.06</td>
<td>Normal</td>
<td>This paper</td>
</tr>
</tbody>
</table>

(1) [46]; (2) [93]; (3) [89]; (4) [94]; (5) [136]; (6) [97]; (7) [118]; (8) [170]

another 0.4 dex in the next 50 days. $UBVRIJHK$ luminosity is different from the bolometric luminosity significantly only before about $+20$ d; but closely resembles the bolometric flux after that. As shown in the bottom panel of Fig. 5.7, the UV portion of the total flux drops from about 38% at $+8$ d to below 10% by $+22$ d. Beyond this point, the optical fraction drops slowly and remains above 60% until the plateau transitions, dropping slightly in the tail phase. The NIR fraction is about 40% during the plateau, and remains steady during the tail phase.

The ratio of $UBVRIJHK$ luminosity to $UBVRI$ luminosity in the calibration sample (see Table 5.5) shows a tight correlation with $B-V$ color. A linear fit of $\log_{10}(L_{UBVRIJHK}/L_{UBVRI})$ versus $B-V$ color is performed using this sample (Fig. 5.9) but more energetic, atypical SN IIP SN 2004et, with largely uncertain $E(B-V)$, is excluded as an outlier. The best fit model is obtained as

$$\log_{10}(\frac{L_{UtoK}}{L_{UBVRI}}) = (0.0856 \pm 0.0012) + (0.1056 \pm 0.0012) \times (B-V). \quad (5.2)$$

A set of histograms showing the residuals of measured $UBVRIJHK$ and modeled $UtoK$ for each event is shown in Fig. 5.10. This calibration yields 1–2% precision for SN 2013ej, SN 1999em, and SN 2005cs. This also yields a $\sim$6% precision for the excluded SN 2004et.
Figure 5.9. Dependence of flux ratio with $B - V$. Strong correlation of log$_{10}(UBVRIJHK/UBVRI)$ with $B - V$ is obtained for SNe 1999em, 2005cs, and 2013ej. The global fit excludes outlier SN 2004et and a $\chi^2$/dof = 1.51 is obtained.

Figure 5.10. Histograms of residuals obtained after subtracting log$_{10}(L_{UtoK}/L_{UBVRI})$ from log$_{10}(L_{UBVRIJHK}/L_{UBVRI})$. The RMS of the differences yield 1–2% scatter for SN 2013ej, SN 1999em, and SN 2005cs. The same calibration model for SN 2004et, which is excluded from the fit, yields $\sim$ 6% precision.
With this, I now perform a two-fold calibration for events with only ROTSE unfiltered photometry: (1) ROTSE unfiltered to $UBVRI$ using the fit in Table 5.4 that yields effective $UtoI$, and (2) $UtoI$, which is effective $UBVRI$ to $UtoK$ using Eq. 5.2. I estimate the relative uncertainty from this procedure to be 6% or less. Performing this to KAIT unfiltered photometry yields similar result.

5.1.4. Spectroscopy

The spectral observations obtained for SN 2013ej are logged in Table 5.1. In the following analysis, all the spectra are corrected for the recession of the host galaxy redshift $z = 0.002192$ (NED/IPAC Extragalactic Database\(^8\)). Earlier, I performed SNID templates fit of early spectra and found the best fitted redshift of 0.002.

5.1.4.1. Key spectral features and evolution

The early time spectra at +8 d, +9 d, and +11 d are blue continuum, with a few P Cygni profiles of neutral H Balmer lines and He I lines. At this stage, all the other ions have very low opacity to be conspicuously observed as spectral features. The H- Balmer lines ($\text{H}\alpha \lambda 6562.85$, $\text{H}\beta \lambda 4861.36$, and $\text{H}\gamma \lambda 4340.49$) are very broad, with subsequently decreasing emission component over time. The +11 d spectrum shows O I $\lambda 7775$. The next spectrum at +19 d shows several strong absorption signatures of typical SNe II.

As noted earlier, an unusually strong absorption line appears early blueward of $\text{H}\alpha$. This feature gets stronger until +19 d in this sample, subsequently appears as a small absorption notch in the +44 d spectrum and disappears in the +48 d spectrum. This feature was observed by Ref. [165] to be stronger than $\text{H}\alpha$ until +21 d and becoming weaker than $\text{H}\alpha$ by +23 d in their dataset. In Ref. [40], this was identified as Si II $\lambda 6355$ (see Section 5.1.4.3). Note that Si-II identification is also favored by Ref. [19, 79, 165]. In typical SN-IIP, Si II seemed to occur much later. While Si II had not been observed this early previously, it may have been marginally detected at +10 d and +12 d in SN 2006bp ([132]) but was gone by

\(^8\)https://ned.ipac.caltech.edu/
+25 d. At similar epochs, Si II in SN 2013ej is much stronger than in SN 2006bp. The velocity profile of Si-II shows faster evolution than that of Hα in SN 2006bp, while it evolves slower than Hα in SN 2013ej. This makes SN 2013ej to exhibit an unusually strong early Si II. The subsequent spectral evolution of SN 2013ej exhibits typical singly ionized lines of Ca II, Fe II, Ti II, Sc II, and Ba II like other IIP events. He I λ5876 gets weaker until +15 d and disappears by +19 d; suggesting the ejecta temperature decreasing below the critical excitation temperature. Heavy iron-group elements appear with the commencement of the plateau phase as the photosphere penetrates deeper. After +19 d, Na I D lines appear stronger than the other neutral elements, presumably due to non-LTE effects ([70]).

No evidence of narrow Na lines and high-velocity features (HVF) is seen in this sample, potentially indicating negligible ejecta interaction with the circumstellar material (CSM). The Ca II near-IR triplet can be dissociated into a doublet at 8520 Å and a singlet at 8662 Å after +19 d, but the profile is blended before +15 d.

5.1.4.2. Spectral homogeneity in the UV

In the UV, SNe -IIP show remarkable homogeneity ([60]). Early-phase UV spectra (2000–3000 Å) of SNe 1999em, 2005ay, and 2005cs show similar continuum shapes and spectral features. On contrary, Ref. [10] pointed that SNe IIb, that are believed to have thinner H-rich envelopes display stronger diversity in their UV. A limited sample of early-time UV spectra of SNe-IIP can not conclude the argument of homogeneity verses diversity at present. So SN 2013ej adds valuable contribution to this sample. Swift observation of near-UV spectra of SN 2013ej with its UVOT/UGRISM instrument are shown in Fig. 5.3. Fig 5.11 shows the +8 d and the +11 d spectra compared to those of other SNe II. Note that spectra are corrected for interstellar extinction and scaled to match the fluxes 2500–3000 Å region. Around 10-12 days after explosion, SN 2013ej UV spectra support UV spectral homogeneity. However, such similarity is not seen at ∼ 1 week spectra. SN 1987A and SN 2005cs both show some deviation from the spectrum of SN 2013ej at this phase. Careful inspection of the UVOT/UGRISM frames revealed that the SN 2005cs spectrum was contaminated by
Figure 5.11. Left: SN 2013ej UV spectrum along with the atypical SN 1987A and a subluminous SN 2005cs. SN 1987A spectrum shows a sharp cutoff. The excess flux for SN 2005cs below 2500 Å is suspected to be some contamination from a nearby source. Right: Homogeneous SN IIP UV sample at $\sim 10 - 12$ d. A SN IIb spectrum is shown for comparison and is clearly distinct from the rest.

emission from a nearby source. In addition, SN 1987A, which shows a cutoff in the UV below 3000 Å, was an atypical SN IIP from a blue supergiant progenitor. For comparison, Fig. 5.11 also shows a UV spectrum of SN-IIb event at a similar epoch.
Figure 5.12. Syn++ fits of representative SN 2013ej spectra. Data are shown in black while respective models are in blue. Observed features are mostly identified from the synthetic fits. H I line profiles are not reproduced, perhaps because of a purely scattering based model that ignores any emission from the recombination cascades and also NLTE effects.
5.1.4.3. Line identification and spectral modeling

Key spectral features explained in Section 5.1.4.1 were first identified with Ref. [70] on ion signatures of SN spectra. Furthermore, confirmations were performed by Syn++ ([157]) modeling optical spectra as shown in Figure 5.12. Although the Syn++ modeling is successful in producing most of the ionic features, the most conspicuous Hα is not reproduced; reflecting the limitation that the code does not account for the emission from recombination cascades. For Hα, both time varying and non-local thermodynamical equilibrium (NLTE) are expected to affect the observed profile. The feature blueward of Hα line in the +11 d and +19 d spectra is fitted with the Si II line, and is shown in Fig. 5.12. This feature can perhaps be argued to be the HVF of Hα, but given the absence of HVFs of other Balmer lines, this is unlikely. Moreover, a HVF input there does not reproduce a decent overall fit. Note that no HVFs were observed in the near-IR sample of Ref. [165]. Although the observed feature is favored for SiII, a higher confidence can be achieved only from more realistic modeling. HVFs have been incorporated by Ref. [19] for H I lines blended with the photospheric component resulting to broader Balmer lines beyond +42 d. Lines of intermediate-mass elements and iron-group elements are observed after +15 d; s-process products Ba II and Sc II are also identified in the +19 d and until the last plateau spectrum at +94 d.

5.1.4.4. Velocity evolution

The ionic and photospheric velocity of the IIP ejecta can be measured from the observed spectra. While it is more convenient to obtain the ion velocities from the absorption wings, the photospheric velocity is generally estimated from weak lines (e.g. [20]), or from global fitting of the spectrum (e.g [170]). For SN 2013ej, I consider the He I λ5876 line before +19 d, and Fe II λ5129 line after that as the photospheric velocity proxy. The velocity evolution of some of the strongest lines in the optical are shown in the left panel of Fig. 5.13. The black circles are the photospheric velocities obtained from Syn++ fits and are given in Table 5.6, and the dashed line represents the expansion method of Ref. [170]. This model will be sampled at photometric epochs for distance determination in the next section.
The line velocities are obtained from the minimum of the Gaussian fit of the absorption profile for each line, transformed to velocity space using relativistic Doppler relation.

The Hα line shows slower deceleration than Hβ and other metallic ions, but the evolution profile of H I lines are relatively flatter, as also pointed by Ref. [19]. Strong correlations of Fe II λ5169 line \(v_{\text{Fe II}}\) with the Hβ line \(v_{\text{H\beta}}\) have been demonstrated by Ref. [51,128] using larger ensemble of IIP events. The linear relation of the two obtained for SN 2013ej from +15–+48 days spectra yields \(v_{\text{Fe II}} = (0.85 \pm 0.03) v_{\text{H\beta}}\) in agreement with \(v_{\text{Fe II}} = (0.84 \pm 0.05) v_{\text{H\beta}}\) as obtained by Ref. [128].

5.1.4.5. Distance measurement with EPM

The theoretical description of EPM is given in Section 2.3. Here, I present the distance measurement of SN 2013ej using Eq. 2.5 adopting the usage of bolometric flux. EPM analysis for a larger ROTSE SN-IIP sample is presented in Chapter 6.

Different techniques have been used to measure the distance to the SN 2013ej host, M74, yielding a value of \(D \approx 7 \pm 2\) Mpc ([143,167,169]) to \(D \approx 9.5 \pm 0.5\) Mpc ([113,182]). EPM analysis have been applied for various samples of SNe IIP ([17,34,37,61,69,83,93,154,170]). Combining the data from two SNe hosted by the same galaxy, one can reduce the uncertainties of distance estimation and improve the reliability of the method ([170]). Using SN 2013ej and SN 2002ap ([169]) for EPM, the distance to M74 is estimated and shown to be in good agreement with other independent studies. Note that SN 2002ap is a broad-lined SN Ic and the application of EPM may be very limited, if justified. While the atmospheres of stripped-envelope (SE) CC SNe are hard to model, the result obtained here using simple approximations may potentially be extended to a larger data set.

EPM basically relies on contemporaneous spectroscopic and photometric observations. In practice, this is difficult to obtain and the method relies on extrapolations from limited measurements. EPM also requires the dilution parameter accounting for deviation of the photosphere from a black body. For SN 2013ej, dilution parameter is taken from Ref. [38] for H-rich SNe IIP. For the H-poor SN 2002ap, in Ref. [40], we set \(\zeta = 1\) as a first approximation,
Figure 5.13. Velocity evolution of strong lines of SN 2013ej. Empty black circles represent the photospheric velocity obtained from the Syn++ fits. The dashed line is the photospheric velocity model using the method from Ref. [170]. Figure is taken from Ref. [40].
Table 5.6. Photospheric velocities of SN 2013ej obtained from Syn++ fitting. Phases are rounded to the closest day.

<table>
<thead>
<tr>
<th>MJD</th>
<th>Phase (days)</th>
<th>$v_{\text{phot}}$ (km s$^{-1}$)</th>
<th>Uncertainty (km s$^{-1}$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>56505.5</td>
<td>+8</td>
<td>10200</td>
<td>1000</td>
</tr>
<tr>
<td>56506.5</td>
<td>+9</td>
<td>9700</td>
<td>1000</td>
</tr>
<tr>
<td>56508.5</td>
<td>+11</td>
<td>8800</td>
<td>800</td>
</tr>
<tr>
<td>56516.5</td>
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<td>600</td>
</tr>
<tr>
<td>56541.5</td>
<td>+44</td>
<td>4700</td>
<td>500</td>
</tr>
<tr>
<td>56545.5</td>
<td>+48</td>
<td>4900</td>
<td>400</td>
</tr>
<tr>
<td>56566.5</td>
<td>+69</td>
<td>3200</td>
<td>400</td>
</tr>
<tr>
<td>56570.5</td>
<td>+73</td>
<td>3740</td>
<td>500</td>
</tr>
</tbody>
</table>

as a lesson learned from distance estimation of the Type IIb SN 2011dh ([170]). As the ejecta of the SN 2002ap practically show no H, unlike the Type IIb SN 2011dh, the dilution of the blackbody flux due to Thompson scattering on free electrons is expected to be even less. Note that the full justification of $\zeta = 1$ would require full NLTE atmosphere modeling for SN 2002ap, which is beyond the scope of this study. An estimate of systematic error on the distance from this approximation was done by Ref. [40] as described below.

Following Eq. 2.5, the angular size $\theta$ is estimated using bolometric flux derived in Section 5.1.3.4. For SN 2002ap, $f_{\text{bol}}$ is derived in Ref. [40] by integrating optical light curves from Ref. [57], [117], and [169] and the near-IR data from Ref. [179]. The UV is estimated using a linear SED between 3000 Å and the $U$ band, and consistency is observed with the spectral shape of SN 2002ap below 4500 Å (e.g., [169]). The next things required for EPM are the temperature and photospheric velocity. For SN 2013ej, these are obtained as explained above, and velocities are interpolated through power law expansion (e.g., [156]). For SN 2002ap, extinction is evaluated from $E(B-V) = 0.09$ mag ([169]). Photospheric velocities are evaluated with Si II features as suggested by Ref. [169] and then expanded
with a power law. Temperatures are modeled following Ref. [101] with

\[ T_{\text{eff}} = -0.122(B - V) + 3.875 \]  (5.3)

Table 5.7 lists the parameters estimated for EPM before performing the final fit. The explosion epochs are set to \( t_0 = \text{MJD 56496.9} \) (2013 July 23.9 UT) and \( t_0 = \text{MJD 52302.0} \) (2002 Jan 28.0 UT) for SNe 2013ej and 2002ap, respectively.

The parameters are fitted using \( \chi^2 \) minimization of Eq. 2.8. This results in \( D = 8.86 \pm 0.21 \) Mpc (fixed \( t_0 \)), while floating \( t_0 \) gives \( D = 9.09 \pm 0.30 \) Mpc and \( t_0 = -0.59 \pm 0.47 \) which is consistent with our estimated \( t_0 \). If \( t \) is taken as independent variable, \( D = 8.93 \pm 0.10 \) Mpc is obtained with fixed \( t_0 \) and \( D = 9.25 \pm 0.30 \) Mpc and \( t_0 = 0.09 \pm 0.48 \) days with a floating \( t_0 \). The weighted average of these four values gives \( D = 8.96 \pm 0.08 \) Mpc. The systematic uncertainty possibly introduced from our \( t_0 \) in Section 5.1.3.1 is estimated taking lower and upper bounds for \( t_0 \) as \(-1.3 \) d (obtained from floating power index, see Section 5.1.3.1) and \(+0.9 \) d (Lulin detection). This adjustment yields a difference of \(+0.35 \) and \(-0.60 \) Mpc respectively. For the final estimate of total uncertainty, these are added in quadrature with the statistical uncertainty obtained above. A distance of \( 9.0_{-0.6}^{+0.4} \) Mpc is the final distance measured for M74 using two SNe. The derived distance can be obtained by inverting the slope of the line shown in Fig 5.14.

In Ref. [40], we cross checked the validity of \( \zeta = 1 \), that was artificially set for SN 2002ap. Repeating the analysis using the dilution model from Ref. [38], which is clearly an overkill for Ic SN atmosphere yielded change less than 0.5 Mpc. Thus, we concluded that the offset is not higher than 0.5 Mpc.

An independent EPM analysis for SN 2013ej was performed by Ref. [133] and they derived \( D = 9.1 \pm 0.8 \) Mpc. Furthermore, with a motivation of performing EPM analysis for ROTSE data, I repeat the distance estimation using the bolometric calibration for ROTSE unfiltered fluxes estimated in Section 5.1.3.4. I only included data points beyond \(+15 \) d for SN 2013ej, as earlier data would be dominated by the \( U \) band which is not included in the calibration procedure. I derive the EPM distance of \( 9.7 \pm 0.6 \) Mpc from SN 2013ej alone,
Table 5.7. Parameters for EPM

<table>
<thead>
<tr>
<th>time (days)</th>
<th>$\theta$ (10$^8$ km Mpc$^{-1}$)</th>
<th>$\theta/v_{\text{phot}}$ (day Mpc$^{-1}$)</th>
<th>Uncertainty (day Mpc$^{-1}$)</th>
</tr>
</thead>
<tbody>
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<td>SN 2013ej</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>8.60</td>
<td>9.44</td>
<td>1.08</td>
<td>0.12</td>
</tr>
<tr>
<td>10.60</td>
<td>10.10</td>
<td>1.24</td>
<td>0.12</td>
</tr>
<tr>
<td>13.60</td>
<td>11.45</td>
<td>1.52</td>
<td>0.17</td>
</tr>
<tr>
<td>14.60</td>
<td>12.78</td>
<td>1.74</td>
<td>0.19</td>
</tr>
<tr>
<td>15.60</td>
<td>12.98</td>
<td>1.81</td>
<td>0.20</td>
</tr>
<tr>
<td>16.60</td>
<td>13.30</td>
<td>1.90</td>
<td>0.20</td>
</tr>
<tr>
<td>19.60</td>
<td>13.93</td>
<td>2.13</td>
<td>0.22</td>
</tr>
<tr>
<td>20.60</td>
<td>15.31</td>
<td>2.39</td>
<td>0.24</td>
</tr>
<tr>
<td>24.60</td>
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<tr>
<td>25.60</td>
<td>15.80</td>
<td>2.72</td>
<td>0.28</td>
</tr>
<tr>
<td>SN 2002ap</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>4.89</td>
<td>11.96</td>
<td>0.44</td>
<td>0.34</td>
</tr>
<tr>
<td>6.48</td>
<td>12.34</td>
<td>0.67</td>
<td>0.30</td>
</tr>
<tr>
<td>7.48</td>
<td>13.00</td>
<td>0.78</td>
<td>0.32</td>
</tr>
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<tr>
<td>12.87</td>
<td>16.15</td>
<td>1.55</td>
<td>0.40</td>
</tr>
<tr>
<td>13.47</td>
<td>16.11</td>
<td>1.51</td>
<td>0.39</td>
</tr>
<tr>
<td>13.86</td>
<td>16.10</td>
<td>1.69</td>
<td>0.39</td>
</tr>
</tbody>
</table>
in agreement with the previous estimation. Adopting the upper and lower bounds to $t_0$ as above, the final distance estimated is $9.7^{+0.9}_{-0.7}$ Mpc. Comparison of the distance derived here is shown in Table 5.8, taken from Ref. [40].

With the distance determined, in the next section, I describe the estimation of some of the physical parameters from the explosion.

5.1.4.6. Explosion properties and kinematics

SNe IIP show a diverse explosion properties such as the energy released from the explosion and the mass of synthesized material. In this section, I derive some of the kinematic parameters for SN 2013ej through direct observation or modeling.

5.1.4.6.1 Ni Mass: The plateau ends when the ionization front, and thus the photosphere, reaches the bottom of the hydrogen envelope. After this, almost the entire luminosity is powered by the radioactive decay of elements that were produced during the explosion. Thus, the light curve is expected to show a characteristic exponential decay. This suggests that the gamma-rays and positrons from radioactive decay thermalize in the ejecta. If one assumes the full trapping of gamma-rays and positrons in the ejecta, the mass of freshly synthesized Ni should be strongly constrained from the tail luminosity.

I show below that the decline rate in fact changes after $+183d$. However, before that I estimate the Ni mass using methods from the literature. A linear fit of the $UBVRI$ light curve from $+120$ d to $+183$ d is performed and the fitted line is extrapolated to $240$ d to make a direct comparison with their result for SN 2012aw ([18]). $L(240\text{ d})$ for SN 2013ej is estimated to be $1.32 \pm 0.05 \times 10^{40}$ erg s$^{-1}$, while that for SN 2012aw was found to be $4.53 \pm 0.11 \times 10^{40}$ erg s$^{-1}$. Assuming total trapping in both cases and using the observed luminosity ratio of $0.29 \pm 0.02$ and using $M_{Ni}$ for SN 2012aw $0.058 \pm 0.002 M_\odot$ ([18]), $M_{Ni} = 0.017 \pm 0.001 M_\odot$ is derived for SN 2013ej. Using the method of Ref. [67]

$$M_{Ni} = 7.866 \times 10^{-44} L_t \exp\left[\frac{(t_t - t_0)/(1 + z) - 6.1}{111.26}\right] M_\odot.$$  \hspace{1cm} (5.4)
Figure 5.14. Measurement of distance to M74 using the EPM technique on SN 2013ej and SN 2002ap data. The solid line represents the final result yielding distance of 9.0 Mpc, and the dotted lines indicate 1σ uncertainty.

Table 5.8. Recent Distance Estimates for M74 using multiple methods

<table>
<thead>
<tr>
<th>Method</th>
<th>$D$</th>
<th>Reference</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>(Mpc)</td>
<td></td>
</tr>
<tr>
<td>T-F</td>
<td>9.68 ± 1.63</td>
<td>[158]</td>
</tr>
<tr>
<td>BBSG</td>
<td>7.31 ± 1.23</td>
<td>[143]</td>
</tr>
<tr>
<td>Disk gravitational stability</td>
<td>9.4</td>
<td>[182]</td>
</tr>
<tr>
<td>Light echo</td>
<td>7.2</td>
<td>[167]</td>
</tr>
<tr>
<td>SCM (SN 2008gd)</td>
<td>9.91 ± 1.2</td>
<td>[113]</td>
</tr>
<tr>
<td>EPM (SN 2002ap)</td>
<td>6.7</td>
<td>[169]</td>
</tr>
<tr>
<td>TRGB</td>
<td>10.2 ± 0.6</td>
<td>[80]</td>
</tr>
<tr>
<td>EPM</td>
<td>$9.0^{+0.4}_{-0.6}$</td>
<td>[40]</td>
</tr>
<tr>
<td>EPM: ROTSE calibrated</td>
<td>$9.7^{+0.9}_{-0.7}$</td>
<td>[40]</td>
</tr>
</tbody>
</table>
$L_t$ is calculated at 20 epochs between +120 d and +183 d of $V$-band magnitude. Adopting the bolometric correction of 0.26 mag from Ref. [67], the weighted mean luminosity is estimated to be $5.82 \pm 0.26 \times 10^{40}$ erg s$^{-1}$, corresponding to +157 d. From this, $M_{\text{Ni}}$ is calculated to be $0.018 \pm 0.002$ M$\odot$. Repeating this for the bolometric light curve derived in Section 5.1.3.4, one gets $M_{\text{Ni}}$ to be $0.019 \pm 0.003$ M$\odot$. An average of the above three values is $M_{\text{Ni}} = 0.018 \pm 0.001$ M$\odot$, taken as my final estimate of Ni mass.

However, with the late time data of SN 2013ej from the KAIT and Nickel telescopes, the slope of the tail decline doesn’t seem unique (see Fig. 5.15). I estimate the epoch of slope break by fitting the late time bolometric flux beyond +120 d to a broken exponential law of the form

$$F(t) = S A e^{-\frac{t}{\tau_1}} \left[1 + e^{\alpha(t-t_{br})}\right]^\frac{1}{\alpha} \left(\frac{1}{\tau_1} - \frac{1}{\tau_2}\right)$$

(5.5)

where, $\tau_1$ and $\tau_2$ are characteristic times for the two exponential models, $A$ is the initial flux, $t_{br}$ is the break time, $\alpha$ is the smoothing parameter and $S$ is the scaling factor, given by,

$$S = (1 + e^{-\alpha t_{br}})^\frac{1}{\alpha} \left(\frac{1}{\tau_1} - \frac{1}{\tau_2}\right)$$

(5.6)

The Eq. 5.5 has been analogously applied to study the radial profile break of the surface brightness from the galactic disks ([106]). The best fit parameters are, $\tau_1 = 73.89 \pm 5.00$ days, $\tau_2 = 94.73 \pm 1.39$ days, $t_{br} = 183.28 \pm 15.67$ and $\alpha = 0.23 \pm 1.14$ and the $\chi^2/dof$ from the fit is 1.55. Before and after the break point, the slopes are obtained to be, $Slope_1 = 0.015 \pm 0.001$ mag day$^{-1}$ and $Slope_2 = 0.011 \pm 0.001$ mag day$^{-1}$. This reveals that $Slope_1$ is significantly steeper than $Slope_2$, which is closer to the $^{56}\text{Co} \rightarrow ^{56}\text{Fe}$ decay rate.

These distinct decline behaviors certainly make impact on the inferred initial Ni mass. In Ref. [40], I fit the late-time bolometric light curve of SN 2013ej with a model following Ref. [169] and [163]. In this model, an optically thin ejecta is heated by the partial trapping of gamma rays and positrons from radioactive decay of $^{56}\text{Ni}$ and $^{56}\text{Co}$. For the gamma-rays,
Relative flux \( \chi^2 / \text{dof} = 1.55 \).

\[ \tau_1 = 73.89 \pm 5.00 \]
\[ \tau_2 = 94.73 \pm 1.39 \]

\begin{align*}
\text{Slope}_1 &= 0.015 +/− 0.001 \text{ mag day}^{-1} \\
\text{Slope}_2 &= 0.011 +/− 0.001 \text{ mag day}^{-1} \\
t_{\text{break}} &= 183.38 +/− 15.67 \text{ d.}
\end{align*}

Figure 5.15. Derived bolometric light curve of SN 2013ej. Beyond +120 d is fitted with a broken exponential model of Eq. 5.5. The best fit model is represented by the solid line, and the dashed lines are the extrapolations of single exponential laws in either direction. The crossover point is at the fitted break time, \( t_{br} = +183.4 \pm 15.7 \text{ d.} \)
the deposition function can be expressed as

\[ D_\gamma = 1 - e^{-\tau_\gamma} = 1 - \exp\left[-\left(\frac{T_0(\gamma)}{t}\right)^2\right], \]  

(5.7)

where \( \tau_\gamma \) is the optical depth for gamma-rays in the whole ejecta. The timescale the gamma-ray optical depth decreases ([175]) is

\[ T_0(\gamma) = \sqrt{C \kappa_\gamma M_{ej}^2 / E_{kin}}, \]

(5.8)

where \( \kappa_\gamma \) is the gamma-ray opacity, \( M_{ej} \) is the ejecta mass and \( C \) is a constant that depends on the density distribution of the ejecta. For simplicity, \( C = 9/40\pi \) is taken following Ref. [175]. For positrons, the deposition function, \( D_+ \), takes the same form but with a different opacity. \( \kappa_\gamma = 0.027 \text{ cm}^2 \text{ g}^{-1} \) and \( \kappa_+ = 7 \text{ cm}^2 \text{ g}^{-1} \) were adopted in Ref. [40] following Ref. [31, 163].

With this, one can express the late-time bolometric luminosity as

\[
L_{bol} = M_{Ni}[(S_{Ni}(t) + 0.92S_{Co}(t))D_\gamma \\
+ (0.03 + 0.05 \times D_\gamma)S_{Co}(t)D_+],
\]

(5.9)

where \( M_{Ni} \) is the initial mass of the synthesized radioactive \( ^{56}\text{Ni} \) during the explosion, \( S_{Ni} \) and \( S_{Co} \) are for the total energy input from the Ni- and Co-decay (see Ref. [21, 153] for more discussion). One simplicity of this model is that it assumes instantaneous release of the deposited energy from radioactive decay, ignoring any photon diffusion unlike the model of Ref. [5]. It is thus reasonable approximation for the epochs when the ejecta is almost entirely transparent, i.e. during the nebular tail phase.

In Fig. 5.16, I show two separate fits of Eq. 5.9 restricting tail data in the pre break (+120 d < t < +183 d) and post break (> +183 d) regime. While the steeper decline in the
Figure 5.16. Radioactive decay model fits using Eq. 5.9 on the tail of the SN 2013ej bolometric light curve. Light curves are separately fitted for the pre and post break times. On the upper panel, the fit is constrained to data points between +120 d and the break point +183 d and On the lower panel, the fit is performed on data beyond the break time. Both the gamma-rays and positron leakage are considered in the model. The red line is for the full trapping of gamma-rays and positrons on both the upper and lower panels.
bolometric light curve in the pre break regime is consistent with the findings from Ref. [79] and [19]; the extended post break photometry in this sample reveals a shallower decay rate. Fit for the post break scenario yields the initial nickel mass and the gamma-opacity timescale as $M_{\text{Ni}} = 0.013 \pm 0.001 M_\odot$ and $T_0(\gamma) = 465 \pm 18$ days. This timescale is much longer than $\sim 173$ day, as obtained by Ref. [19] using their data between +100 and +200 days. A longer $T_0(\gamma)$ indicates that the light curve of SN 2013ej was probably not fully transitioned to the radioactive tail before +183 days. I derive the Ni mass of $M_{\text{Ni}} = 0.019 \pm 0.001 M_\odot$ from the pre-break fit. This result is consistent with above result, obtained using methods from Ref. [18] and [67], and also with independent estimates by Ref. [19, 79]. Note that both Ref. [79] and [19] assumed full gamma-ray and positron trapping inside the similar time window. However, these values are likely overestimated and the value derived here and in Ref. [40] from later time analysis beyond +183 d is probably correct.

5.1.4.6.2 Kinematic parameters: Hydrodynamic study (e.g., [79]) and semi-analytic approach (e.g., [19]) following Ref. [6] have been done for SN 2013ej in order to estimate the explosion properties. Here I make approximate estimates from an independent data set using the approach of Ref. [96]. While complications concerning the radioactive heating and explosion energy impacting the model exist, simple assumptions can be useful to compare the kinematic parameters with those of the more extensive studies.

Using plateau duration ($\Delta t$), the mid-plateau absolute $V$ band magnitude $M_V$, and the concurrent expansion velocity, Ref. [96] derive expressions for explosion energy, ejected mass, and pre-SN radius. For SN 2013ej, I estimate plateau mid point using $(t_{\text{peak}} + t_p)/2$; where $t_{\text{peak}}$ is the time of peak brightness in $V$ and $t_p$ is the epoch at which the plateau ends. Using Gaussian process regression (See Appendix B) of $V$ band data until +30 d, I estimate $t_{\text{peak}}$ to be +15 d, and for the end of plateau I again perform a Gaussian process regression on the bolometric lightcurve from +80 d to +130 d, from which the point of inflection is estimated as $t_p$ to be +109 d. With this, I estimate the plateau duration of $94 \pm 7$ days with mid-point on +62 d. Now $v_{\text{ph}}$ at +62 d is determined to be $3800 \pm 500$ km s$^{-1}$ using fit in
Section 5.1.4.4, while $M_V$ is estimated to be $-16.47 \pm 0.04$ mag using linear interpolation of $V$ magnitudes between +50 d and +70 d. Using these estimates, relations from Ref. [96] give explosion energy of $0.9 \pm 0.3 \times 10^{51}$ ergs, and a pre-SN stellar radius of $250 \pm 70 \, R_\odot$. Note that Ref. [19] finds an explosion energy of $2.3 \times 10^{51}$ ergs and a radius of 450 $R_\odot$. Likewise, Ref. [79] constrain energy in the $0.7 - 2.1 \times 10^{51}$ ergs range and radius in the $230 - 600 \, R_\odot$ range. As the uncertainty in Ref. [19] is not stated, and Ref. [79] give a range of values, these calculations appear to be consistent with theirs. Both energy and radius calculated here are consistent with the SNe IIP ensemble studies of Ref. [67] and [108], which also use the same Ref. [96] relations.

Similarly, the ejecta mass, $M_{ej}$, for SN 2013ej is estimated using Ref. [96] relations to be $13.8 \pm 4.2 \, M_\odot$. Ref. [67] and [108] show an ejecta mass distribution in the range 14–56 $M_\odot$ and 10–30 $M_\odot$, respectively. For SN 2013ej, Ref. [79] and [19] find ejecta masses of 10.6 $M_\odot$ and 12±3 $M_\odot$ respectively. Incorporating a dense core and an extended low mass envelope in their light curve model, Ref. [109] derive an ejecta mass of 10.6 $M_\odot$. These estimates are all consistent with the measurement based on the Ref. [96] model. Assuming a remnant mass of 1.4 $M_\odot$, the final pre-explosion progenitor mass for SN 2013ej is $15.2 \pm 4.2 \, M_\odot$. Note that when quoting the mass, care must be taken to explicitly distinguish the progenitor mass, pre-explosion mass or ZAMS mass. The mass estimated here is for the final pre-explosion mass. Using archival HST data, Ref. [59] found the ZAMS mass to be in the 8–15.5 $M_\odot$ range, which is consistent with an M-type supergiant.

Using X-ray observations, Ref. [27] estimated steady mass loss of $3 \times 10^{-6} \, M_\odot \, yr^{-1}$ off the progenitor over the last 400 years. From this, they estimated a ZAMS mass of 14 $M_\odot$. Given the uncertainties in these measurements, my estimation of the final progenitor mass of $15.2 \pm 4.2 \, M_\odot$ using LN85 is consistent. On the general population of SNe IIP, Ref. [147] obtain a progenitor ZAMS mass distribution between 8 to 17 $M_\odot$. Nebular phase spectral modeling (e.g. [82]) have shown to provide tighter constraints on the ZAMS of these events. However, results from detailed hydrodynamical modeling (e.g. [160, 161]), stellar evolutionary models (e.g. [147]) and nebular spectra modeling show conflicts in the derivation
of initial masses of these progenitor stars. Table 5.9, taken from Ref. [40] summarizes the findings of this analysis and compares to the existing literature results.

5.2. Evidence of interaction of SN 2012cg with a main sequence companion

Even though SNe Ia show remarkable similarity after a correction in their peak-shape properties, there are uncertainties associated with the explosion mechanisms, mass distribution of the progenitors and so on. As the surveys are becoming more sensitive, these events are being captured at their earliest phases, sometimes within a day if not hours after explosion (e.g., [77, 98]). Such early detections often reveal some peculiar behaviors that challenge the normally expected concept of explosion. While such early time peculiarities do not restrain their use for cosmology from observations around and after the peak, they might well put systematic constraints on our understanding of explosion scenarios.

Observation of SN 2012cg as early as -17d relative to $B_{\text{max}}$ from 5 telescopes in 7 photometric filters, including ROTSE IIIb unfiltered photometry, show an excess in the optical emission in the B band ([98]). This study showed the first convincing direct evidence of impact of an otherwise normal SN explosion with a non degenerate companion; thereby supporting a single degenerate (SD) model of Type Ia SN explosion. It does not discard however any other potential explosion scenario.

5.2.1. ROTSE observations

ROTSE IIIb detected SN 2012cg initially on May 17.178, 1.1 hrs before the discovery epoch reported in Ref. [145]. We did not observe any emission at the SN position on May 16.177 at 5σ detection threshold, limiting magnitude 16.9. The initial photometry employed a subtraction of host using the host subtraction prescription described in Section 5.1.2.1. The obtained photometry was calibrated to USNO B1.0 $R$-band magnitude. Since the CCD quantum efficiency is sensitive to as low as 3000 Å, SN 2012cg ROTSE photometry was adjusted to $B$-band zero point in the analysis, after calibrating to APASS $V$ band using stars within 3’ radius from the SN position. ROTSE photometric data is published in Ref. [98].
Figure 5.17. Early time light curve of SN2012cg in multiple bands including ROTSE IIIb from Ref. [98]. The ROTSE data are adjusted to $B$ band zero point. Excess emission is evident in all cases on -16d and -15d from $B_{max}$. Dashed black lines are the normal SN 2011fe data, while dashed green lines are $t^n$ power law model fits restricted to -14d to -8d observations.
Table 5.9. Calculated Physical Parameters of SN 2013ej. Table is taken from Ref. [40].

<table>
<thead>
<tr>
<th>Parameter/Reference</th>
<th>[19]</th>
<th>[79]</th>
<th>[59]</th>
<th>[165]</th>
<th>[40]</th>
</tr>
</thead>
<tbody>
<tr>
<td>Explosion Energy ($10^{51}$ ergs)</td>
<td>2.3</td>
<td>0.7–2.1</td>
<td>–</td>
<td>–</td>
<td>0.9 ± 0.3</td>
</tr>
<tr>
<td>Progenitor Mass (Pre-explosion) ($M_\odot$)</td>
<td>14.0 ± 3.0</td>
<td>12–13 (ZAMS)</td>
<td>8–15.5</td>
<td>–</td>
<td>15 ± 4.2</td>
</tr>
<tr>
<td>Pre-SN Radius ($R_\odot$)</td>
<td>450 ± 112</td>
<td>230–600</td>
<td>–</td>
<td>400–600</td>
<td>250 ± 70</td>
</tr>
<tr>
<td>$M_{NI}$ ($M_\odot$)</td>
<td>0.019 ± 0.002</td>
<td>0.02 ± 0.01</td>
<td>–</td>
<td>–</td>
<td>0.013 ± 0.001</td>
</tr>
<tr>
<td>Plateau Duration (Days)</td>
<td>~ 85</td>
<td>~ 50</td>
<td>–</td>
<td>–</td>
<td>94 ± 7</td>
</tr>
<tr>
<td>Distance Assumed (Mpc)</td>
<td>9.57 ± 0.7</td>
<td>9.6 ± 0.7</td>
<td>9.1 ± 1.0</td>
<td>9.1</td>
<td>–</td>
</tr>
<tr>
<td>Distance Measured (Mpc)</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>9.0^{+0.4}_{-0.6}</td>
</tr>
</tbody>
</table>

5.2.2. Photometric analysis and discussion

Fig. 5.2.1, taken from Ref. [98], shows early time light curves of SN 2012cg in different photometric bands. At -16d, the excess emission in $B$ band is 0.2 mag, which is statistically significant, when compared to a normal SN Ia evolution at similar times. Fig. 5.2.1 shows the early flux deviation from a normal SN 2011fe, and a $t^n$ power law model, constrained between -14 to -8 days. The spectral observations at similar epochs, presented by Ref. [98] also showed excess in the blue. Performing an explosion model fit, they showed that the excess was consistent with the interaction of the SN with the non generate main sequence star of mass 6 $M_\odot$ (See Fig. 6 in their paper). While ROTSE unfiltered photometry has maximum response in the mid optical regime, there is significant efficiency in the high energy region of the optical, and some fraction of the UV. The early time excess emission of SN 2012cg is seen for both $U$ and $B$ bands. High energy photon emission is expected if coming from interaction, while the NIR observation in Ref. [98] lack the emission and are on the usual trend. -16d and -15d short term excess in the light makes $B - V$ diverge from the available SN data and normal Type Ia models. Assuming no angular dependence, Ref. [98] found these observations consistent with the SN interaction with a binary main sequence companion, as predicted by Ref. [84] models but the analysis also can not rule out an off axis interaction with a higher mass companion. Type Ia have shown a wide range of ejected
mass (e.g [137]), thus the explosion mechanism is not fully understood. This study warrants more observation of SN in the earliest times after explosion to statistically study the Type Ia explosion mechanisms. Note that several other studies (e.g [63]) have findings consistent with this study, while other work using very late time data of SN 2012cg by Ref. [142] have added constraints on the explosion that are contrary to the claims made here. These analyses suggest that a larger sample of the early time observations of SN Ia is important to address the issues that are directly connected to their explosions.

5.3. Photometric analysis of SN 2013df with ROTSE IIIb

The late time physics of stripped core SNe are suggested to directly constrain the early time behavior (e.g. [175]) but this does not appear to be obvious from the photometric analysis of an ensemble of such events. With a motivation to understand the tail behaviour of a stripped core SN event, I performed a late time analysis of SN 2013df combining ROTSE data with others from the literature. The basic physics of the late time is similar to SN IIP and is presented in Section 5.1.4.6.1 for SN 2013ej. But late time connection with early time, where the H- recombination is still prominent, is not justifiable as in the case of stripped core SNe. Here I summarize the results from the analysis of SN 2013df.

Because of the lack of a good template, I limited this analysis using data set upto +150 days after explosion ([152]). A crude aperture photometry based host subtraction algorithm was adopted and this was not robust enough to get the late time low S/N photometry of the object. Host subtraction was also performed for SN 2013ej as discussed in Section 5.1.2.1.

5.3.1. Early and late time light curve analysis

Figure 5.18, taken from Ref. [152], shows the light curves at early and late times. The early time light curve shows a typical secondary peak on all bands. ROTSE data start off from the minimum after the first peak. Using early time data, Ref. [152] construct a bolometric light curve with two structural configuration: A He rich core, and a (H+He) rich shell; and a He core and a pure H-rich shell. It was found that most of the initial peak flux
comes from the shell interaction, lasting for about 10 days. After that, a smooth emission from the core collapse pertains.

In the late time, the ROTSE data is combined with other observations by Ref. [152] to constrain the explosion parameters. Following the prescription in Section 5.1.4.6.1, a $\gamma + e^+$ leakage model of Eq. 5.9 is considered. The characteristic diffusion timescale for the $\gamma$ rays ($T_0$) is found to be 95 days and that for the $e^+$ ($T_+$) is found to be about 1500 days from the fit of the late time. An ejecta mass of 3.7–4.2 $M_\odot$, and kinetic energy of $1.6–2.3 \times 10^{51}$ ergs were derived from these results in Ref. [152]. Although Ref. [152] show internal consistency of these parameters with the early time analysis in the same study, the overall result was found to be contradictory to the observed late and early time discrepancy on the analysis of Ref. [175] using an ensemble of stripped envelope core-collapse SNe.

I was able to make a better template for SN 2013df a few months after this analysis was performed. The obtained lightcurve with the new image differencing is given in Section 4.3.2. While the earlier reductions are consistent with the new image differencing at similar epochs, the new light curve yields more detections and reduced photometric errors.

With this, I conclude the astrophysical analyses of three individual SNe and in the next chapter, I present a cosmological analysis of measuring the Hubble constant using nearby ROTSE SNe-IIP sample. One of the key lessons learned in terms of photometric reduction and analysis was that the ROTSE photometry of both the SN 2012cg and SN 2013df was tricky as the existing RSVP image subtraction code was not fully efficient to model the host background accurately. SN 2013ej occurred on the outskirts of the host galaxy, therefore, an aperture photometry was sufficient. While such object specific reduction procedure on three cases could be calibrated accurately, I was convinced that a more consistent and self contained image subtraction tool was needed before performing any study involving multiple SNe. I will present new software I developed for image differencing in the first section of the next chapter.
Figure 5.18. Left: Konkoly $BVRI$ and ROTSE unfiltered lightcurve of SN2013df at early time, compared with Ref. [104] (MG14) and swift data. Right: The late time analysis of SN 2013df using ROTSE and literature data. The solid line is the best fit model assuming an incomplete trapping of gamma rays and positrons from the radioactive decay in the ejecta. The legend gives the best fitted characteristic trapping timescales of $\gamma$ and $e^+$ respectively, derived in Ref. [152]. Both figures are taken from Ref. [152].
In this chapter, I first present the mathematical framework of new image differencing code that is used to reduce the ROTSE SNe light curves. In the following section, I present a cosmological analysis in the low redshift universe using ROTSE SNe IIP sample. While making the cosmological measurements, I utilize several methods of calibration for various photometric and spectroscopic observables of each SNe in the sample and develop analysis tools that may be used by the larger community for independent data sets. This will set up a stage for photometric, spectroscopic and redshift dependent calibrations for the distance measurement of each of the events. Once the distances are estimated, I present the measurement of the Hubble parameter in the nearby universe.

6.1. Image differencing for ROTSE SNe photometry

Today is the era of precision cosmology. SNe, by nature, live in their host galaxy. Their photometric background varies from host type and position of the SN from the host nucleus. An accurate estimation of background is essential to precisely measure the intrinsic light signal from the SN alone. Although SNe can sometimes outshine the whole galaxy, it is challenging to exactly account for the only SNe light, by accurately subtracting the host light and this becomes more critical when the signal is photon limited. Furthermore, the observations are also subjected to spatial and temporal variations of the observing conditions, and the rapid variation in the nature of an event can provide additional complication. Pixel boundaries can create artifacts that can over or under estimate signal during subtraction. This problem is more significant if the pixel size is large and the host galaxy background has a steep gradient. I describe an optimal image differencing method and its implementation for ROTSE SNe photometry. This code is called ImageDiff and is public through my github
repository: https://github.com/gdhungana/ImageDiff. At present, it is specific to the ROTSE data model but the plan is to make it experiment agnostic in the future.

6.1.1. Point Spread Function modeling

Image differencing involves the subtraction of a reference image from a new science image such that any object not in the reference image can be obtained from the subtracted (difference) image. Because of the variations of various parameters, such as the point spread function (PSF), between the science and the reference images, care must be taken in the image differencing. The ImageDiff fundamentally utilizes kernel based PSF modeling, where using the reference image of presumably higher signal to noise ratio, I model a new image equivalent to science image but without the SN. A similar approach was used for the cross convolution technique adopted by RSVP ([181]), but here, I do not convolve the science image but only the reference image with the kernel. I also add flexibility in the kernel choice in this work.

A science image $S(x, y)$ is modeled from convolving a template image $T(x, y)$ with a PSF kernel $K(u, v)$, plus a noise term $\epsilon(x, y)$, such that

$$S(x, y) = (K \otimes T)(x, y) + \epsilon(x, y)$$ (6.1)

Here, $x, y$ are pixel coordinates and $u, v$ are kernel coordinates. I build a kernel basis such that, $K(u, v) = \sum_i c_i k_i(u, v)$. This now becomes a linear problem.

$$S = \sum_i A_i c_i + \epsilon$$ (6.2)

where, $A_i = k_i \otimes T$. I intend to find coefficients $c_i$ of the linear combination, corresponding to each kernel $k_i$. Assuming Gaussian errors, the maximum likelihood (minimum $\chi^2$) solution for the coefficient matrix will be

$$C = (A^T N^{-1} A)^{-1} A^T N^{-1} S$$ (6.3)
where $N$ is the pixel noise matrix, which is a diagonal matrix as the pixel errors are treated as statistically uncorrelated. The inverse of the covariance matrix $A^TN^{-1}A$ must exist. A small prior is added at the level of machine precision to ensure the matrix remains well conditioned.

The difference image is then simply given in pixel coordinates by

$$D(x, y) = S(x, y) - \left( \sum_i A_i c_i \right)(x, y) \quad (6.4)$$

6.1.1.1. Choice of Kernels

I use three different kernel bases,

1. Sum of Gaussians basis: Gaussian functions multiplied by 2-dimensional polynomials

$$k_i(u, v) = e^{-\left(u^2+v^2\right)/2\sigma^2}u^p v^q \quad (6.5)$$

$i$ runs over all permutation of $n$, $p$, $q$. The polynomial order expansion used is $0 \leq p + q \leq O_n$. The default choices for 3 Gaussians is $\sigma = [0.7, 1.5, 3.]$ with $O_n = [4, 3, 2]$. So, total number of kernels = $\sum_n (O_n + 1) \times (O_n + 2) / 2 = 31$

2. Gauss Hermite polynomial basis: A Gaussian core is multiplied by Hermite polynomials giving

$$k_i(u, v) = e^{-\left(u^2+v^2\right)/2\sigma^2}H_m(u)H_n(v) \quad (6.6)$$

where $H_n(x) = (-1)^n e^{x^2} \frac{d^n}{dx^n} e^{-x^2}$ is the $n^{th}$ order Hermite polynomial. An obvious merit of Gauss-Hermite kernel over symmetric Gaussian Kernels is that this can model asymmetry because of odd-even functions of Hermite polynomials. An asymmetric PSF can occur for many different reasons such as atmospheric conditions. Fig. 6.2 shows the 1D Gauss Hermite polynomial functions in the left panel and shapelet basis functions defined by two integers on the right panel. Both figures are adopted from Ref. [99].
Figure 6.1. Image subtraction of a $280 \times 280$ pixel subimage of ROTSE tss1246+1249 field science image (middle) using sum of Gaussian kernel convolution of template image (left). SN 2004gk is clearly visible at the center of the residual image (right).

Figure 6.2. Left: 1D Gaussian weighted Hermite polynomials. The number of peaks goes as $(n+1)$. Right: 2D projection of a 3D shapelets up to 6th order. Increase of linear extent with higher order is evident. Both figures are adopted from Ref. [99].
3. Delta function basis: This kernel constitutes only delta functions

\[ k_{i,j}(u, v) = \delta(u - i)\delta(v - j) \] (6.7)

A 11 \times 11 pixel size kernel has 121 orthonormal, single pixel bases. The benefit of the delta function is that it is shape independent, so there is no parameter to tune. However, this may need regularization to ascertain well conditioning of the solution for the model and prevent overfitting. See Ref. [9] for an application of delta function kernels.

6.1.1.2. Using Principal Component Analysis (PCA)

PCA is a technique of reducing data dimensionality without losing any significant feature of the data. The principal components are the eigenvectors of the covariance of the dataset. They are sorted by the eigenvalues in the decending order, i.e there is least variance associated with the component with highest eigenvalue. From the kernel-convolved templates, I construct an orthogonal eigen basis using PCA. This work uses the \texttt{empca} package ([7]) to compute the PCA using the expectation maximization (EM) method. Appendix A gives the mathematical formalism of the EM method.

<table>
<thead>
<tr>
<th>Kernel Type</th>
<th>$\chi^2$/dof</th>
<th>R value</th>
<th>Residual/\sigma</th>
</tr>
</thead>
<tbody>
<tr>
<td>Gaussian sum</td>
<td>0.82</td>
<td>0.97</td>
<td>$\mathcal{N}(-0.01, 0.91)$</td>
</tr>
<tr>
<td>Gauss Hermite</td>
<td>0.82</td>
<td>0.97</td>
<td>$\mathcal{N}(-0.01, 0.90)$</td>
</tr>
<tr>
<td>Delta Function</td>
<td>0.81</td>
<td>0.97</td>
<td>$\mathcal{N}(-0.01, 0.90)$</td>
</tr>
<tr>
<td>EMPCA</td>
<td>0.85</td>
<td>0.93</td>
<td>$\mathcal{N}(0.03, 0.94)$</td>
</tr>
</tbody>
</table>
Figure 6.3. Top: ROTSE rqa0137 + 1547 field subimage from a recent image (middle), higher S/N template image (left) and the difference image using \texttt{ImageDiff} (right). Bottom: A slice from the center of the subimage showing the galaxy M74 profile, the fit model and the residual (left); same residual in a 2D shown in 80 × 80 pixels for clarity, R is the measure of fraction of observation variance preserved in the difference image; pull distribution showing a Gaussian fit (red) and a zero mean, unit variance normal distribution (black) in the right. The pull shows nearly standard normal distribution.
An observation $\vec{y}$ can be expanded in the PCA eigenbasis $\{\vec{\phi}\}$ as

$$\vec{y} = \sum_k c_k \vec{\phi}_k \quad (6.8)$$

By orthogonality of the eigenvectors,

$$\vec{y} \cdot \vec{\phi}_m = \sum_k c_k \vec{\phi}_k \cdot \vec{\phi}_m = \sum_k c_k \delta_{km} = c_m \quad (6.9)$$

The aim here is to solve for the eigenvectors, while the coefficients are evaluated in a maximum likelihood sense when fitting to data. Weights are applied as the inverse of the variance for each pixel,

- **X** - Data Matrix: stack kernel convolved templates + 1 science image
- **P** - Matrix of Principal Component Eigenvectors
- **C** - Matrix of coefficient vectors to be determined
- **V** - Covariance between the variables of all observations.

We want to minimize,

$$\chi^2 = (X - PC)^T V^{-1} (X - PC) \quad (6.10)$$

Then, the $\chi^2$ minimization of Eq. 6.3 given an observation $\vec{x}$ becomes

$$\vec{c} = (P^T N^{-1} P)^{-1} P^T N^{-1} \vec{x} \quad (6.11)$$

An example showing the performance of the `empca` method on the same field as in Fig. 6.3 but at a later time with SN is shown in 6.4.
Figure 6.4. *Top:* Image differencing using PCA method, of SN 2013ej, in the same field as in Figure 6.3. *Bottom:* A ROTSE follow up SN 2013df, where the SN event is very near to the core of a huge background galaxy. Images are luminance normalized to 90% confidence interval.
6.1.2. Simulation

A PSF model profile with a Gaussian core and a wing component, allowing ellipticity variation is considered from Ref. [16].

\[ I(x, y) = \frac{(1 - b) e^{-r^2/2 \sigma^2}}{\sqrt{2\pi} \sigma} + \frac{be^{-r/r_0}}{2\pi rr_0}; r_{\text{ell}} = \sqrt{qx^2 + y^2/q} \] (6.12)

- \( b \): controls the wing contribution
- \( r \): radial offset from the PSF spot
- \( r_0 \): characteristic size of the wing
- \( q \): ellipticity
- \( x \) & \( y \) are related to CCD coordinates by rotation/translation.

Monte Carlo (MC) simulation is performed by injecting objects of random magnitudes at random locations within a subimage of a ROTSE data image. To disallow tight blending with the point sources in the image, scikit ([119]) \( k \) – \( d \) tree query is performed taking a radius of 1 FWHM of the PSFs of the data image, derived by the sextractor. For the injected sources, the image subtraction is performed iteratively one by one and the final photometry is performed on the difference image. The extracted magnitudes are compared with the input magnitudes. An example simulation for the ROTSE field rqa0137+1546 is shown in Fig. 6.5. The root mean square (RMS) of the photometry residuals is at 0.05 magnitude level, and a pull distribution shows nearly standard normal \( \sim N(0.1, 0.94) \) distribution.

6.1.3. Performance

In most of the SNe analysed, the new image subtraction yields 10-20% more detections, and the scatter of the residuals also is remarkably narrower. Fig. 6.6 shows an example of old versus new image subtraction for SN 2004gk. Each light curve is fitted with Gaussian Process (GP) regression using scikit. On the rightmost plot is shown the residuals of new and old light curves obtained after subtracting the GP best fit models. It is observed that
Figure 6.5. Left: 1000 MC objects on top of a data image. Points show injected sources but do not represent the true PSF shape of the simulated objects. Right: Performance of ImageDiff. Overall residual mean is 0 yielding no bias, RMS $\sim$0.05 magnitude; while close to the limiting magnitude, RMS $\sim$0.1 mag. Pull distribution $\sim \mathcal{N}(0.1, 0.94)$.
the new image differencing not only has higher detection efficiency but also has over 3.5 times improvement in the residuals. The pull distributions are found to be $\mathcal{N}(0.01, 1.03)$ for the new and $\mathcal{N}(0.39, 2.82)$ for the old method. The typical pulls on the other SN light curves obtained with ImageDiff also follows within 10% of a standard normal distribution.

This now sets up a self consistent data reduction and analysis to proceed with the cosmological analysis using an ensemble of SNe.

6.2. SNe IIP distance measurement: Methodology

As discussed in Chapter 2, cosmological analyses using SNe IIP have been performed using multiple techniques. The main idea is to measure distances of the events at various redshifts and perform a cosmological model fit. The most commonly used methods are the standard candle method (SCM) and the expanding photosphere method (EPM). In Chapter 2, I described the basic assumptions for both of these methods and gave the mathematical framework for EPM in Section 2.3. Here I derive the distances of nearby ROTSE SNe IIP sample by adopting the EPM method. This method exhibits rich literature through several independent studies for IIP SNe (e.g., [17, 34, 37, 61, 69, 83, 93, 154, 170] etc.). Using EPM method, the distance of the nearby SN 2013ej was also measured using bolometric flux in Ref. [40], as presented in Chapter 5. In this analysis, I will only use the unfiltered photometric observations and the optical spectroscopy.

6.2.1. EPM requirements

As shown in Section 2.3, the requirements for estimating distance with the EPM involve a set of concurrent photometric and spectroscopic observations. But there are observational limitations to this as it is very rare to obtain a time series concurrent photometric and spectroscopic observations. Therefore, the motivation here is to establish empirical relations of the physical parameters when the photometric and spectroscopic data are limited to as little as one observation. Even though the ROTSE sample, as you will see, has substantial photometric observations, spectroscopic observations are limited, sometimes to just one
Figure 6.6. Performance of new and old image differencing methods. The reduced data points are normalized to Ref. [47] V band magnitude (shown in blue points), on MJD 53389.0. The solid line is a Gaussian Process regression fit, with the filled region 95% confidence posterior prediction. The rightmost panel shows the residuals of old (green) and new (blue) light curves after subtracting the respective best fits. The rms scatter for the green and blue histograms are respectively 0.36 and 0.10 mag, and the pulls on the new image differencing yield a dispersion of 1.03, which is substantially improved compared to the old differencing, where the pull yields a dispersion of 2.82.
observation. Such constraint will impact bigger surveys as well where the telescope time is limited. Thus it is important to establish generic analytic and empirical models to characterize these events for cosmological studies, which otherwise would require a thorough time series data.

The photometric observations yield the flux output, from which the angular size can be related to the distance. The SN is treated as an expanding diluted black body (BB), and the size inferred from the observed flux is compared to the size obtained using the expansion velocity that can be obtained from spectra. Following Eq. 2.8, given the observed flux in a particular wave band and thus the angular size $\theta$, distance can be obtained if the expansion velocities ($v_{\text{phot}}$) are known. An obvious need here is the interpolation/extrapolation modeling as the photometric and spectroscopic observations are not concurrent. The following Sections describe the methods to model the evolution of $T$ and $v_{\text{phot}}$ based on few observations, which can then be sampled at the photometric epochs for each event.

6.2.1.1. Determination of $v_{\text{phot}}$

The photospheric velocity can be estimated for a given epoch using the velocities of weak lines in the SN spectra, as these are expected to form at the photosphere (e.g., [156]). During the plateau phase, the absorption minimum of P-Cygni profiles of FeII $\lambda$4924, $\lambda$5018 and $\lambda$5169 appear to be the best estimator of $v_{\text{phot}}$. In the earlier times, where these lines are usually absent, He I $\lambda$5876 is a good estimator ([155]). Modeling of the spectroscopic observations can provide better estimate as that would account for not just the local spectral features such as a single line properties, but also global features such as the photospheric velocity, ejecta density profile and so on. In the study of SN 2013ej described in Chapter 5, I modeled the spectra using Syn++, from where the photospheric velocities were obtained directly as a model parameter. It was, however, also shown that the ionic velocities from FeII were consistent with the evolution of global photospheric velocity. Therefore, when observed, I will use Fe II 5169 absorption minimum as a proxy of photospheric velocity. For early times, I use He I $\lambda$5876 line or H Balmer lines and transform to photospheric velocities.
using observed correlations as you will see below.

Estimating the minimum of the absorption profile via a Gaussian model fitting requires clean, unblended absorption feature. This also requires the accurate estimation of the continuum. To avoid potentially large uncertainty associated with isolating a single line, I use 1D Gaussian mixture model (GMM) approach to fit a fairly wide section of the spectrum (typically several hundred Angstroms) about the line of interest. The idea behind this is that the GMM method can accurately model even the blending of lines as separate components and also correct for the bias that can be introduced from inaccurate continuum subtraction.

First, a continuum is estimated using side bins (masking the central line profile) with a smooth cubic spline fit of a section of a spectrum in the rest frame. On the absorption profile after subtracting the continuum, a mixture of 1D Gaussians is fitted iteratively varying the number of Gaussian components to model the residual. The best model is selected as the one that yields the minimum Bayesian information criterion (BIC). BIC is similar to the Akaike Information Criterion (AIC) and the idea is to find a maximum likelihood solution given the model. It also considers both the unknown parameters in the model, in this case, the means and standard deviations of the Gaussian components, and the total degrees of freedom, which helps to avoid overfitting. Both BIC and AIC only tell the relative likelihood from a set of models, and not that the model is the correct model in the absolute sense.

An example GMM fit, on a SN 2004gy spectrum observed on Jan 10, 2005, is shown in Fig. 6.7. The top left plot shows the region of interest chosen as ± 600 Å about the Hβ λ4861 line. A continuum is estimated using cubic spline fit on the side bins after masking the P-Cygni profile window 4500 – 5000 Å. The residual on the absorption is fitted with GMM varying the number of components from 1 to 9. A mixture of 6 Gaussian components is favored as suggested by minimum BIC (bottom right) and the best fit is shown on the top right along with the individual components. From the best fit, the velocity of Hβ λ4861 line is calculated to be 10,762 km/s using the component overlaid with the Hβ absorption profile and is shown on the bottom left plot. The uncertainty is calculated from this Gaussian component using the standard error on the mean.
Figure 6.7. Estimation of the ionic velocity of Hβ line using a mixture of Gaussian model. Top Left: Spectrum of SN 2004gy taken on Jan 10, 2005. Continuum is estimated using spline fit masking the central window 4500 – 5000 Å, while the full region is ± 600 Å. Top Right: The normalized residuals after subtracting the continuum in the absorption portion. Histograms show the density distribution in the region. The best fit mixture model and the respective 6 components are shown. Bottom Left: Estimation of Velocity from the best fit. The dashed line shows true rest frame position of Hβ while the solid line shows the position of the absorption minimum. Bottom Right: BIC estimation for mixture models with multiple number of components (n) evaluated iteratively for n=0 to 9.
The direct measurements of the photospheric velocities from spectra have to be extrapolated to the photometric epochs for EPM analysis. Several studies have shown strong correlations on the evolution of velocities for different lines (e.g., [51, 128]). Correlation of Fe II λ5169 velocity with Hα and Hβ velocities from Ref. [51] are shown in Fig. 6.8. The left plot shows the evolution of FeII line calibrated against 50 day value. The middle and the right plot show the observed correlations of Fe II line velocities with H Balmer line velocities for an ensemble of IIP SNe. I will use these correlations for sampling velocities in Section 6.3.2.

6.2.1.2. Determination of $T$

Typically, the effective temperature is estimated as color temperature ($T$); which is evaluated using a BB fit of the SED constructed from the broadband photometric measurements. As I am limited to only ROTSE observations, this procedure is not possible. Therefore, generic analytic relation for the evolution of temperature is modeled and used for calibration of the data I have. From the observation of the temperature evolution using multiband photometry of the well studied SNe in the literature, an exponentially decaying behavior appears to accurately model the temperature evolution for the epochs considered. The left panel of Fig. 6.9 shows the temperature evolution of well studied IIP SNe from the literature. When both the epochs and temperatures of each object are calibrated against 50 day values and the same exponential model fit is performed on the combined distribution, the rms scatter reduces down significantly and the fit yields a reasonable $\chi^2$. A 50 day choice is made to match the velocity evolution model, and for IIP, this choice has been a common use case as it is about the midpoint of the typical plateau length, where the evolution is relatively stable. The exponential model used is given by

$$\frac{T(t)}{T(50)} = a + b e^{c(t/50)} \quad (6.13)$$

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Figure 6.8. Left: Velocity evolution of Fe II λ5169 line from Ref. [51], and Middle and Right: Correlations of Fe II velocity with the Balmer Hα and Hβ line.

The best fit ($\chi^2/dof = 1.04$) model parameters for the full sample shown in Fig. 6.9 are obtained to be $a = 0.908 \pm 0.012$, $b = 2.662 \pm 0.091$; and $c = 3.492 \pm 0.143$. A power law model with declining index of $-0.44 \pm 0.01$ also gives a reasonable fit but with a poorer $\chi^2/dof = 1.6$, compared to my favored exponentially declining model.

6.2.1.3. Determination of epochs for EPM analysis

The epochs considered for EPM, as mentioned in Chapter 2, should satisfy the basic assumptions that the ejecta expansion is homologous and radiation is isotropic. Hydrodynamical simulations of SNe IIP (e.g., [159] have shown that the initial acceleration from the shock settles within 1-2 days, and this brief accelerated phase transitions to a homologous expansion phase. For the analysis presented here, I conservatively consider epochs between $\sim 7$ and 35 days after adopted explosion epoch. Although, in principle, the later times could also be used, this cut sets an advantage that the photometric calibration is less affected from potential systematics due to emerging spectral features in the $V$ band that I will calibrate ROTSE measurement against in the next Section.
Figure 6.9. Left: Evolution of color temperature of four well sampled IIP SNe, derived from fitting $BVI$ fluxes to Planck function. Color coded dashed line represent respective exponential decay law fits using Eq. 6.13. Right: Decay law global fit after calibrating the epochs and temperature with 50 day values for each of the SNe in the sample. The best fit model, shown in line yields $\chi^2/dof = 1.04$. 

Model: $\frac{T(t)}{T(50)} = a + b e^{-c(t/50)}$
6.3. ROTSE SNe IIP sample

The ROTSE IIP sample is constructed from all SNe from 2004 onward. With the requirement that there are several photometric measurements in the 1 week-5 weeks window and at least one spectrum, the final sample constitutes 12 SNe IIP given in Table 6.2.

6.3.1. Photometric calibration

For the EPM study, I perform relative photometry for each observation to $V$ band measurement by calibrating field stars with the APASS\textsuperscript{1} DR9 catalog data. Several calculations of EPM have avoided R band measurements, so as to bypass potential systematics from a strong Hα signature in a IIP spectrum. Note that the ROTSE CCD response is significantly different than $V$ band filter response as shown in Fig. 4.1. While the peak of ROTSE response falls near 5500 Å, a $V$ band calibration has been performed in the past and equally justifiable as $R$ band calibration. The use of $V$ band in this case is more sensible: firstly to avoid the strong Hα feature, and secondly to build consistency with the dilution parameter model that also avoids $R$ band. However, the accuracy of the $V$ band calibration must be established for a rapidly evolving SED. To quantify any potential offset associated with the calibration of ROTSE flux with the $V$ band flux of the field stars, I perform a Monte Carlo simulation of blackbody continuum spectra of varying temperature 2 kK-17 kK, randomly normalized to $V$ band magnitude in the range 12-18. The temperature and magnitude ranges are chosen so as to agree with the SN IIP temperatures and magnitudes in the similar epochs for the sample. The respective blackbody spectrum are convolved with the ROTSE response function and $V$ band filter response function, and magnitudes are estimated. As shown in the left plot of Fig. 6.10, it is clear that there is over and undercorrection. The offset is analytically modeled with an exponentially growing function of temperature ($T$). This correction yields no fundamental bias, and the residual rms is only 0.01 mag, which is much smaller than the typical statistical uncertainty of the ROTSE magnitudes. I propagate this as un-

\textsuperscript{1}https://www.aavso.org/apass
Table 6.2. ROTSE IIP sample for the EPM study

<table>
<thead>
<tr>
<th>SN</th>
<th>Program</th>
<th>ROTSE Field</th>
<th>Host Galaxy</th>
<th>Spectra Used</th>
<th>$z$</th>
<th>$E(B-V)_\text{tot}$</th>
<th>Adopted $t_0$ (MJD)</th>
</tr>
</thead>
<tbody>
<tr>
<td>SN 2004gy</td>
<td>TSS</td>
<td>skc1307+2626</td>
<td>NGP9</td>
<td>1</td>
<td>0.02600 ± 0.00100</td>
<td>0.0100 ± 0.0007</td>
<td>53362.9 ± 2.5</td>
</tr>
<tr>
<td>SN 2005ay</td>
<td>TSS</td>
<td>tss1152+4327</td>
<td>NGC 3938</td>
<td>3</td>
<td>0.00270 ± 0.00001</td>
<td>0.0183 ± 0.0002</td>
<td>53452.5 ± 4.0</td>
</tr>
<tr>
<td>SN 2006bp</td>
<td>TSS</td>
<td>tss1159+5136</td>
<td>NGC 3953</td>
<td>4</td>
<td>0.00351 ± 0.00001</td>
<td>0.4000 ± 0.0100</td>
<td>53833.7 ± 2.0</td>
</tr>
<tr>
<td>SN 2006bj</td>
<td>TSS</td>
<td>tss1220+0756</td>
<td>Anon.</td>
<td>1</td>
<td>0.03770 ± 0.00100</td>
<td>0.2000 ± 0.0004</td>
<td>53815.3 ± 3.0</td>
</tr>
<tr>
<td>SN 2008bj</td>
<td>RSVP</td>
<td>sks1155+4643</td>
<td>MCG +08-22-20</td>
<td>1</td>
<td>0.01806 ± 0.00001</td>
<td>0.2060 ± 0.1000</td>
<td>54534.0 ± 2.5</td>
</tr>
<tr>
<td>SN 2008gy</td>
<td>RSVP</td>
<td>sks0117+1352</td>
<td>SDSS J012044.48+144139.6</td>
<td>1</td>
<td>0.05906 ± 0.000053</td>
<td>0.2823 ± 0.0582</td>
<td>54726.9 ± 3.5</td>
</tr>
<tr>
<td>SN 2008in</td>
<td>RSVP</td>
<td>tss1224+0440</td>
<td>NGC 4303</td>
<td>3</td>
<td>0.00522 ± 0.00001</td>
<td>0.1000 ± 0.1000</td>
<td>54825.1 ± 2.1</td>
</tr>
<tr>
<td>SN 2008bj</td>
<td>RSVP</td>
<td>tss1209+4958</td>
<td>NGC 4088</td>
<td>3</td>
<td>0.00252 ± 0.00001</td>
<td>0.3670 ± 0.0070</td>
<td>54925.0 ± 5.0</td>
</tr>
<tr>
<td>PTF10gva</td>
<td>RSVP</td>
<td>tss1225+1112</td>
<td>SDSS J122355.39+103448.9</td>
<td>1</td>
<td>0.02753 ± 0.00012</td>
<td>0.263 ± 0.0008</td>
<td>55329.3 ± 0.9</td>
</tr>
<tr>
<td>SN 2013ab</td>
<td>TS’3</td>
<td>vsp1443+0953</td>
<td>NGC 5669</td>
<td>2</td>
<td>0.00456 ± 0.00001</td>
<td>0.044 ± 0.0066</td>
<td>56339.5 ± 1.0</td>
</tr>
<tr>
<td>SN 2013bu</td>
<td>TS’3</td>
<td>skt2237+3425</td>
<td>NGC 7331</td>
<td>1</td>
<td>0.002722 ± 0.000004</td>
<td>0.078 ± 0.0006</td>
<td>56399.3 ± 1.0</td>
</tr>
<tr>
<td>SN 2013ej</td>
<td>TS’3</td>
<td>rspt0137+1547</td>
<td>NGC 0628/M74</td>
<td>5</td>
<td>0.00219 ± 0.000003</td>
<td>0.0610 ± 0.0010</td>
<td>56496.9 ± 0.3</td>
</tr>
</tbody>
</table>

correlated systematic uncertainty in the final photometry. Note that, one could equivalently derive a color dependent correction from synthetic broadband magnitudes. As broadband observations are not available for the sample, and I am explicitly measuring temperatures, I have used a temperature dependent correction model. The final corrected photometry during the plateau phase is compared with $V$ band data of several IIP SNe, and are found to be statistically consistent.

$$m_{\text{ROTSE, } V} - V = a + b(1 - e^{cT}) \equiv \text{corr}$$ (6.14)

where $m_{\text{ROTSE, } V}$ is the magnitude obtained from calibrating to $V$ data before correction. The final calibrated magnitude is then $m_{\text{ROTSE, } V - \text{corr}}$. Best fit model parameters obtained using the simulation of 100 random blackbody spectra yield, $a = -9.46 \pm 0.11; b = 9.52 \pm 0.11$ and $c = (8.06 \pm 0.048) \times 10^{-4}$. The state before correction, the correction model and state after correction for the 100 monte carlo sample is shown in Fig. 6.10. The rms of the residuals is about 0.01 mag and is propagated to the final uncertainty.

6.3.2. Sampling of $v_{\text{phot}}$ and $T$
Figure 6.10. Correction of systematic effects of calibrating ROTSE magnitudes to catalog $V$ band data. Shown is the offset for 100 Monte Carlo blackbody spectra spanning 2000K to 17000K temperature, randomly normalized to $V$ band mag range 12-18 (left), offsets varying with temperature and the correction model given by Eq. 6.14 (middle) and offsets after applying correction (right).
Figure 6.11. Lightcurve of ROTSE IIP sample. The magnitudes are calibrated to APASS V band and corrected using 6.14, extinction corrected and $K$-correction applied for events with $z > 0.01.$
For each SN in the sample, the photospheric velocity at the spectral epoch is measured using the Fe II $\lambda$5169 line (if available), HeI $\lambda$5876 line or by using correlations with the Balmer lines presented above in Section 6.2.1.1. For sampling velocity at photometric epochs, each measured velocity is extrapolated using the power law model from Ref. [51] as shown on the left plot of Fig. 6.8. For objects with multiple spectra, the final photospheric velocity evolution is statistically derived from the weighted mean of the derived models, and uncertainty is estimated using the square root of weighted variance of the model values at each epoch.

To estimate the effective color or blackbody temperature, each spectrum is first corrected for redshift. Further, unreddening of the spectrum is performed by applying the reddening curve using parametrization of Ref. [56] for the galactic extinction. For photometry, color extinction $E(B - V)$ are obtained from literature for most of the SNe, while Ref. [138] is used for the Milky-way extinction when no information is available.

If the observed spectral coverage is wide enough, a set of synthetic $BVI$ magnitudes are derived at the rest frame, and the resulting SED is fitted to Planck function to estimate the effective temperature. If the spectral coverage is not enough, the corrected spectrum is fitted to Planck function after masking H$\alpha$ and other telluric lines from the atmosphere when present. This measured temperature is used to generate the expansion model using Eq. 6.13, which is used to sample temperatures at photometric epochs between +7 and +35d. Like in the photospheric velocity, for objects with multiple spectra, the final temperature is estimated from the weighted mean of derived model values at each photometric epoch. Table 6.3 tabulates the derived +50 day values of photospheric velocity and temperature for all SNe in the sample.

6.3.3. $K$-correction

As discussed in Chapter 1, for the SNe with cosmologically significant redshift, the photometric magnitudes have to be accounted with a $K$-correction. While the SNe IIP sample is limited with ROTSE data and mostly single spectrum, I derive a $K$-correction model by
Figure 6.12. Photospheric velocity and the derived model sampled at photospheric epochs using Ref. [51]. Blue points is measured from spectrum using Mixture of Gaussians method. The respective lines used are labelled. Dashed lines represent the extrapolation of photospheric velocity and red points indicate sampling at photometric epochs. When multiple spectra are used, error are evaluated using weighting scheme.
Figure 6.13. Evolution of effective color or BB temperature($T$). Blue points are the measured temperatures obtained from spectral or SED fit to Planck function. Red points are the sampled model temperature at the photometric epochs.
transforming the spectra of the nearby sample to the respective higher redshifts, assuming the color evolution for the SNe IIP remain uniform. For each SN that needs a $K$-correction, spectroscopic data in plateau of all the nearby SNe are redshifted and both the observer frame and the SN rest frame magnitudes are evaluated for $V$ band. The difference in the magnitudes of the rest frame from the observer frame gives the $K$-correction. A GP regression is performed on the obtained $K$-correction values, to make a prediction of $K$-correction for the desired photometric epochs of that particular SN. An example GP regression fit and 95% confidence level posterior prediction is shown in Fig. 6.14 for SN 2004gy and SN 2008bj.

6.3.4. Systematic uncertainty

EPM exhibits several systematic uncertainties. The uncertainty due to the dilution parameter, $\zeta$, is hard to model. I used the polynomial model from Ref. [38] for $\zeta$, and do not attribute any systematic uncertainty from $\zeta$ in this EPM calculation. Now that temperature and velocity are modeled, I account the systematic uncertainties from the models on the adopted explosion epoch. The uncertainty in the galactic (based on Ref. [138]) and host extinction $E(B-V)$ when available are also accounted for in the final measurements. Additional systematic uncertainty from the $K$-correction modeling is also propagated. In addition, the dominant systematic uncertainty comes from the adopted $t_0$, the time of explosion. So better knowledge of explosion times can reduce the total error significantly.

EPM distances are derived for the 12 ROTSE SNe as shown in the Fig. 6.16. Five of the SNe, which have $z > 0.01$ are accounted for $K$-correction before the final EPM distance measurement. For the nearby objects, this is negligible and thus is ignored. Thus, the derived distance should be treated as the Luminosity distance. Distance modulus are calculated from the derived distances and a Hubble Diagram is constructed for the ROTSE IIP sample as shown in the Fig. 6.16.
Figure 6.14. $K$-corrections derived spectroscopically from the nearby SNe sample at $z = 0.0269$, corresponding to SN 2004gy, and at $z = 0.019$, corresponding to SN 2008bj. The shaded region represents the 95% confidence region of the posterior prediction of the GP regression.

Table 6.3. Summary of EPM parameters and derived distance for the IIP sample

<table>
<thead>
<tr>
<th>SN</th>
<th>$v_{\text{phot}}(50)$ ($10^3$ km s$^{-1}$)</th>
<th>$T(50)$ ($10^3$ K)</th>
<th>Distance (Mpc)</th>
<th>$t_0$ (days)</th>
</tr>
</thead>
<tbody>
<tr>
<td>SN 2004gy</td>
<td>4.9 ± 0.2</td>
<td>5.2 ± 0.2</td>
<td>115.5 ± 7.9 ± 14.9</td>
<td>−2.3 ± 1.6 ± 2.5</td>
</tr>
<tr>
<td>SN 2005ay</td>
<td>3.8 ± 0.1</td>
<td>4.8 ± 0.1</td>
<td>21.6 ± 0.4 ± 4.1</td>
<td>0.8 ± 0.4 ± 4.0</td>
</tr>
<tr>
<td>SN 2006bj</td>
<td>4.5 ± 0.2</td>
<td>8.5 ± 0.4</td>
<td>136.9 ± 8.2 ± 38.0</td>
<td>0.4 ± 1.1 ± 4.0</td>
</tr>
<tr>
<td>SN 2006bp</td>
<td>4.3 ± 0.3</td>
<td>4.6 ± 0.2</td>
<td>19.4 ± 0.4 ± 3.4</td>
<td>0.5 ± 0.6 ± 2.0</td>
</tr>
<tr>
<td>SN 2008bj</td>
<td>5.3 ± 0.3</td>
<td>6.2 ± 0.4</td>
<td>89.3 ± 3.6 ± 15.8</td>
<td>3.6 ± 0.8 ± 2.5</td>
</tr>
<tr>
<td>SN 2008gd</td>
<td>4.9 ± 0.3</td>
<td>6.4 ± 0.3</td>
<td>200.5 ± 20.5 ± 50.3</td>
<td>−2.8 ± 2.5 ± 3.5</td>
</tr>
<tr>
<td>SN 2008in</td>
<td>2.6 ± 0.3</td>
<td>5.5 ± 0.2</td>
<td>15.3 ± 0.2 ± 2.9</td>
<td>0.2 ± 0.4 ± 2.1</td>
</tr>
<tr>
<td>SN 2009dd</td>
<td>3.7 ± 0.4</td>
<td>4.6 ± 0.3</td>
<td>14.3 ± 0.3 ± 1.4</td>
<td>−1.4 ± 0.5 ± 1.3</td>
</tr>
<tr>
<td>PTF10gva</td>
<td>5.4 ± 1.0</td>
<td>6.5 ± 0.2</td>
<td>148.6 ± 8.5 ± 22.5</td>
<td>−1.2 ± 1.5 ± 0.9</td>
</tr>
<tr>
<td>SN 2013ab</td>
<td>4.2 ± 0.2</td>
<td>4.9 ± 0.2</td>
<td>24.2 ± 0.5 ± 2.9</td>
<td>−0.4 ± 0.5 ± 1.0</td>
</tr>
<tr>
<td>SN 2013bu</td>
<td>3.1 ± 0.2</td>
<td>3.7 ± 0.2</td>
<td>17.0 ± 0.7 ± 2.0</td>
<td>1.8 ± 0.8 ± 1.0</td>
</tr>
<tr>
<td>SN 2013ej</td>
<td>4.4 ± 0.4</td>
<td>5.5 ± 0.3</td>
<td>8.9 ± 0.2 ± 0.2</td>
<td>0.1 ± 0.4 ± 0.3</td>
</tr>
</tbody>
</table>
Figure 6.15. Distance estimates of the ROTSE IIP sample using EPM.
Figure 6.16. Hubble Diagram constructed from ROTSE IIP sample. Solid and dashed lines are for the flat cosmology with $\Omega_m = 0.3, \Omega_\Lambda = 0.7$, and $H_0 = 70 \pm 5$ km s$^{-1}$ Mpc$^{-1}$. The values are not corrected for local peculiar motions, which can dominate the Hubble expansion significantly in the very nearby universe.
### 6.4. Measurement of Hubble parameter

With the derived distance (or distance modulus), one can perform parameter estimation using a cosmological model. Since this sample is limited to low $z$, I only try to estimate the Hubble parameter, and not try to further constrain any other parameters of the cosmological model. I now define a model for distance modulus following 1.23 as

$$\mu_{\text{model}} = 5 \log_{10}(D_L(H_0|z_{CMB}, \Omega_M, \Omega_{\Lambda})) + 25 \quad (6.15)$$

For parametric inference, it is common to perform a Markov Chain Monte Carlo (MCMC) simulation. Using the publicly available package EMCEE²([58]), the MCMC can be performed minimizing the negative Likelihood function,

$$-2 \ln \mathcal{L} = \sum_{\text{SN}} \left\{ \frac{[\mu_{i,\text{meas}} - \mu_{i,\text{model}}]^2}{\sigma_{\text{tot}}^2} + \ln(\sigma_{\text{tot}}^2) \right\} \quad (6.16)$$

Note that sum is implied for all the SNe in the sample. The total uncertainty includes the measurement errors ($\sigma_{\mu}$), errors from the model and an additional systematic uncertainty ($\sigma_{\text{int}}$) that accomodates an integration of all the factors that are not accounted in the model; all added in quadrature. In this case, $\sigma_{\text{int}}$ is the minimum uncertainty of using EPM for cosmological distance measurements and accounts for the intrinsic scatter in the Hubble diagram.

$$\sigma_{\text{tot}}^2 = \sigma_{\mu}^2 + \left( \frac{5(1 + z)}{z(1 + z/2) \ln(10)} \right)^2 + \sigma_{\text{int}}^2 \quad (6.17)$$

With this set up, I use a flat universe ($\Omega_M + \Omega_{\Lambda} = 1; \Omega_M = 0.3$), and the only free parameters are Hubble parameter ($H'0$) and the intrinsic uncertainty $\sigma_{\text{int}}$. The problem now is to evaluate a posterior probability distribution of the free parameters. In Bayesian framework, the joint posterior probability function for the parameters is

$$p(H_0, \sigma_{\text{int}}|z_{CMB}, \mu, \sigma_{\mu}, \Omega_M, \Omega_{\Lambda}) \propto p(H_0, \sigma_{\text{int}})p(\mu|z_{CMB}, \sigma_{\mu}, H_0, \sigma_{\text{int}}) \quad (6.18)$$

²http://dfm.io/emcee/current/
The likelihood function \( p(\mu|z_{CMB}, \sigma_\mu, H_0, \sigma_{int}) \) is given by Eq. 6.16; while for the prior \( p(H_0, \sigma_{int}) \), I chose a relatively less informative flat prior given as

\[
50 < H_0 < 90; \ -10 < \ln \sigma_{int} < 2
\]

(6.19)

MCMC sampler is constructed picking 500 random inital points about the maximum likelihood estimation. Then MCMC is performed on another 500 iterations yielding 250000 points for the joint posterior distribution. The final 2D and 1D marginalized posterior distribution for the parameters are shown in the corner plot in the left of Fig. 6.17. Contours represent 1, 2 and 3\( \sigma \) confidence regions. The vertical and horizontal blue lines represent the maximum likelihood estimate from the data. The corresponding Hubble diagram is shown in the right plot, with solid black line representing the best fit model, while the gray lines represent the MCMC samples. The bottom inset is the Hubble residual with blue dashed lines showing the best estimated \( \sigma_{int} = 0.8 \) mag scatter.

The intrinsic scatter is quite high, reflecting the fact that low \( z \) events have not been corrected for local peculiar motions. Next, I pick a subsample \( cz_{CMB} > 3000 \) km s\(^{-1}\) and perform the same analysis. This subsample is dominated by the Hubble expansion and galactic peculiar velocity contribution to \( z \) is at \( \sim 10\% \) level (e.g., Ref. [86], who used 300 km s\(^{-1}\) for the peculiar velocity contribution). A low redshift cutoff of \( z = 0.02 \) was shown to be an appropriate lower bound by Ref. [33] for SN Ia cosmology as the observational magnitude uncertainties above this redshift were found significantly larger than the redshift uncertainties. With the redshift cutoff at 0.01 for this sample, only one event SN 2008bj occurs below \( z \) of 0.02 at \( z = 0.019 \), but has large (>20\%) uncertainty in the distance modulus. Hence the cutoff choice of \( z = 0.01 \) should not significantly impact the final result. Therefore, I repeat the above analysis in the subsample, from which the obtained posterior distributions of the parameters and the Hubble diagram along with the residuals are shown in Fig. 6.18. The best fit parameters from this analysis are found to be \( H_0 = 73.5^{+5.6}_{-5.5} \) km s\(^{-1}\) Mpc\(^{-1}\) and \( \ln \sigma_{int} = -5.63^{+3.03}_{-2.99} \). \( \sigma_{int} \) is very small compared to the full
sample, implying that the measurement uncertainties, if Gaussian, are reasonably estimated. The above analysis suggests that EPM method can be a powerful tool for IIP SNe cosmology at higher redshifts. In the analysis above, I showed that the distances could be estimated with as little as one spectroscopic observation, unlike the SCM or other photometric methods, where the observations need to be much denser and over multiple bands concurrently. Denser sampling will obviously increase the statistics and reduces the uncertainty of the EPM results as well but the low scatter in the observed evolution of physical parameters suggest that concurrent observations are not absolutely necessary. With only 5 SNe, the intrinsic scatter was obtained to be less than 0.1 mag. The accurate systematics from the dilution and potential bias from the survey can affect this result slightly but the method itself is promising to pursue on a larger data set, potentially to higher redshifts, as they become available.
Figure 6.18. Same as Fig. 6.17 but for a subsample with recession velocity $cz_{\text{CMB}} > 3000$ km s$^{-1}$. The intrinsic uncertainty $\sigma_{\text{int}}$ is much smaller as this sub sample has peculiar velocities dominated by the Hubble expansion.
Chapter 7

DESI: TELESCOPE, SOFTWARE AND SURVEY

In the last chapter, I presented analyses on cosmological distance measurements in the nearby universe using SNe, where expansion rate was constrained for low redshift universe. Now I move on to BAO method described in Chapter 3, geared towards measuring cosmological distances in the much wider redshift range, focussing on the upcoming Dark Energy Spectroscopic Instrument (DESI) experiment. DESI will create a 3D map of the universe up to high redshift using the largest spectroscopic sample by obtaining the spectra of galaxy targets $\sim 10 \times$ the existing sample from the baseline survey spanning 14000 deg$^2$ area of the sky for about five years. To achieve this, DESI pushes the sensitivity to the noise level employing a high degree of precision spectroscopy of photon limited spectral features of target galaxies. This is supported by sophisticated hardware and equally advanced software components. Below, I briefly give an overview of the hardware, the software systems and the survey description of the DESI experiment.

7.1. Telescope overview

DESI will be installed on the 4-meter Mayall telescope$^1$, located at 6875 feet altitude at the Kitt Peak, Arizona. The Mayall telescope currently uses MOSAIC corrector optics ([35]) and is being used for a precursor imaging survey for DESI. It is a Cassegrain reflector with a hyperbolic secondary while the primary is a parabolic mirror. The telescope is currently equipped with a wide field camera consisting 8K $\times$ 8K MOSAIC CCD. The telescope is a National Optical Astronomy Observatory facility. Kitt peak observes very good observing conditions. In the current $z$ band imaging, the median seeing observed with the Mayall is

$^1$https://www.noao.edu/outreach/kptour/mayall.html
1.1" and the overall transparency is also high. The telescope dome is designed to withstand wind over 120 mph and the facility experiences dry Arizona climate and observes very low light pollution as the nearest significant pollutant Tucson, Arizona is more than 50 miles away.

The existing system of the Mayall telescope will be decommissioned for DESI installation in 2018. The DESI instrument design is shown in Fig. 7.1. The instrument system mainly constitutes of

1. Prime focus corrector optics:- This constitutes the hardware for the optical assembly including 6 large corrector lenses, before the light hits the focal plane. The components of the corrector assembly are designed to meet all the DESI science requirements.

2. Focal plane assembly:- The focal plane system receives the light after passing through the corrector system. It is supplied with 5000 robotic fiber positioners, each of which place a fiber on a specific target.

3. Optical fiber system:- A set of 5000 optical fibers collect light from the focal plane, where each fiber is programmed to position targets with robotic fiber positioners. These fibers carry the optical light from the focal plane to the spectrograph.

4. Spectrograph systems:- The light received from the fibers are dispersed into spectra using 10 spectrographs, each divided into three arms covering the spectral range $\sim 3500$ - $10000$ Å. For each spectrograph, 500 spectra are projected into a $4k \times 4k$ CCD.

Below I give more detailed description for each component. The design of each component is motivated by the broader science requirements for DESI.

7.1.1. Prime focus corrector optics

While DESI utilizes the existing 4-meter primary mirror at the Mayall, the prime focus corrector system is new, facilitating a surface area of 8 deg$^2$ in the prime focus. This corrector assembly constitutes of four large fixed fused silica lenses that provide the main optical focussing and two rotating borosilicate lenses that provide the correction for the atmospheric
Figure 7.1. A layout of DESI block diagram, taken from instrument design section of the final design report in Ref. [36]. It outlines the 5 major subsystems of the DESI hardware.
dispersion. The correction assembly design provide a $3.2^\circ$ field of view, which suffices a 14000 deg$^2$ footprint to be observed 5 times over the 5 year survey period. This design meets the requirement to obtain spectra of over 30 million targets. The whole corrector system is enclosed inside a barrel at the end of which lies the focal plane, interfaced by a Focal Plane Adaptor. Outside of the barrel lies the Prime Focus Cage that provides the hexapod support to the barrel. The hexapod system provides mechanical support as the telescope rotates, by changing the positions with respect to gravity. A sketch of the corrector system is shown in Fig. 7.1 and further details can be found in the DESI final design report in Ref. [36].

7.1.2. Focal plane assembly

A total of 5000 fiber positioners span the focal plane, at a fiber density of 667/deg$^2$. A set of fiducials and Guide, Focus and Adjustment (GFA) imagers (cameras) are situated on the periphery of the focal plane. The field fiducials provide light sources for the fiber view camera that enables monitoring of fiber positions while the GFA measures telescope pointing and focus. The focal plane is divided into 10 pie-slice-shaped wedges, called petals. Each petal contains 500 fiber holes which will be filled in with the robotic fiber positioners holding the fiber tips. This makes the DESI focal plane system very unique and advanced over previous surveys.

Each fiber positioner encompasses two rotational degrees of freedom, and can reach to any point inside a 6 mm radius. A sketch of a positioner’s rotational kinematics is shown on the left of Fig. 7.2. The right plot shows coverage regions of the neighboring positioners. Each positioner carries a fiber, uniquely assigned to a science target during observation. Each reconfiguration of fiber positioner takes $< 24$ sec, which is much smaller than the requirement of 45 sec. After each positioning, the fiber view camera takes the image of the focal plane and the reconfiguration is performed iteratively correcting the fiber offsets until the positional precision $< 5\mu m$ is achieved within the time budget. The lifetime moves required for each positioner is 372,000; and tests show no degradation in over 400,000 moves.
The GFA system constitutes 10 identical cameras that can operate standalone at ambient temperature. A fraction of these cameras perform the guiding of the telescope with a precision better than 0.03”, while the remaining will be used to monitor focussing and allignment. The focussing requirement is that 50 defocussed stars be processed within 5 seconds. The pointing accuracy is expected within 20 seconds after the telescope slew.

7.1.3. Optical fiber system

The light from the full 5000 targets received at the fiber tips on the focal plane gets transmitted to the spectrograph via the optical fibers. Each fiber has a 107 µm core, about 49.5m long, and mounted in a positioner arm. The fibers are bundled in a group of 500 and carried inside a fiber cable. The fiber optics suffers the bulk transmission losses from glass, end-imperfections and surface reflections. Additionally, the focal ratio degradation diffuses the output light beam in larger solid angle compared to the light input. These effects overall affect the throughput of the fiber-optics. For DESI, the total fiber throughput is ∼40% at 3600Å and ∼85% at 9800Å.

7.1.4. Spectrograph system

The light coming out of the optical fibers are collimated with a collimator mirror and are beam-split using two dichroics into three channels (arms): Blue, Red and near-infrared (NIR). The dichroic efficiency in both reflection and transmission bands is above 95% in the no-crossover region, while the combined efficiency is above 90% in the crossover. All 5000 fibers light are dispersed into the three channels by holographic gratings in 360-590 nm, 566-772 nm and 747-980 nm; and are recorded into the 30 (3 × 10) spectrograph CCDs. The corresponding resolutions required for Blue, Red and NIR spectrographs are 2000, 3200 and 4100 respectively ([36]).

Each CCD is 4k × 4k with 15 µm pixels. A schematic optical design of the spectrograph system is shown in Fig. 7.1.4. The estimated throughput for each channel overwhelmingly surpasses the requirements as shown on the top panel of Fig. 7.1.4. An example section of
Figure 7.2. The kinematics of a fiber positioner showing rotational freedom in the “θ − φ” plane (left), and a sketch showing the coverage regions for neighbouring positioners (right). The regions labeled with ‘2x’ indicate a coverage by two positioners.
a simulated CCD image with 500 spectra corresponding to one camera is shown in Fig. 7.5. On the CCD, the fibers are bundled in 25, and fiber separation is about 7 pixels.

Each camera in the three channels of a spectrograph comprises 5 lenses and two aspheric surfaces that provide optical correction for the dispersed light. The camera is housed in its own cryostat, that is required to cool CCDs to 145K (LBNL CCDs for the red and NIR channels) and 170K (ITL CCD for the blue channel) at precision of 1K. The bottom panel of Fig. 7.1.4 shows the Quantum efficiency of the three CCD types.

7.2. DESI data systems

The DESI data systems covers software infrastructure relevant to scheduling, simulation and data reduction, including target selection, survey planning and strategy, spectroscopic data reduction and processing pipeline, data transfer, archive and distribution and data and software management.

A pipeline for generating a catalog of targets using the multi-band photometry from the imaging surveys comprises the target selection pipeline. A selection scheme is supplied for selecting the targets for observations. These targets include the main galaxy targets Emission Line Galaxy (ELG), Luminous Red Galaxy (LRG), Quasars (QSO), Bright Galaxies (BGS), Milky-way stars (MWS), standard stars and targetless sky locations.

The DESI observing planning includes assignment of particular fibers to particular targets (fiber assignment (FA)), the selection of observing field and the online data reduction that facilitates real time decisions for the following exposures. FA utilizes the constraints of fiber positioners on the focal plane and optimally assigns which targets get observed by which fibers. This is particularly critical since the dimmer targets are expected to be observed multiple times. After FA, the decision is made of which field is observed at what time. Optimum selection of fields depends on real time constraints. A priority order is initially set and the real time conditions will be used to finally decide the next optimal field using an algorithm called Next Field Selector. Once the observing starts and raw CCD images are obtained, the online data reduction (QuickLook) reduces data in quasi-real time for moni-
Figure 7.3. The spectrograph schematics for DESI, taken from [36]. The optical path for each of the three arms of the spectrograph is illustrated.
Figure 7.4. Top: Estimated throughput of DESI spectrographs for three channels. Green line is the cumulative response, and the dashed line is the requirement. Bottom: Quantum efficiency for DESI ITL and LBNL CCDs. Both figures are taken from Ref. [36].
Figure 7.5. A simulation of a DESI image on a spectrograph CCD. The fibers span horizontally (cross-dispersion) and dispersion is on the vertical direction. Fibers are bundled in a set of 25, each separated by about 7 pixels. Curvature in fiber trace is evident from the position of bright spots from spectral lines.
toring of data quality assurance (QA) and feedback. The QuickLook outputs are interfaced through a framework for databasing and real time monitoring by the observer using the online system. Exposure times for the successive exposures will be optimized in quasi-real time based on the QA outputs from the previous exposures. The dynamic exposure length is a feature of DESI that will help to maximize the survey efficiency. Previous experiences from surveys like BOSS showed significant loss of survey time due to more stable exposure lengths.

The offline spectroscopic pipeline constitutes a set of steps to convert the raw DESI telescope 2D images to a set of 1D flux-calibrated spectra that are finally passed to redshift fitting code for redshift measurement and classification. Catalogs incorporating the 3D positions of the targets are then constructed to facilitate the cosmological large scale structure and other science analyses.

Data transfer off of the mountain, back up, and distribution is managed by the data system. Full documentation of the software, maintainence of software repositories\(^2\), and releasing of data and software is also facilitated by the DESI data system. A schematic block diagram of the DESI data system is shown in Fig. 7.2. The blocks represent software, data and external subsystems to gather the relevant information. The arrows show the flow of information and cross talks among blocks.

Below I give a brief description of the offline spectroscopic data reduction pipeline, and the main stages involved.

### 7.2.1. Spectroscopic pipeline

Spectroscopic data reduction pipeline is one of the central pieces of the DESI experiment. DESI employs both offline and online processing pipelines. The software is managed and maintained in the open source DESI github repository\(^3\). It is worth summarizing here the

\(^2\)https://github.com/desihub  

\(^3\)https://github.com/desihub/desispec
Figure 7.6. A schematic block diagram of DESI data system data flow. Figure is taken from [36].
steps of the pipeline. The online specific detailed processing will be discussed in Chapter 8.

Before a science exposure raw data can be processed, a set of calibration inputs must be generated. This includes both 2D image level and 1D spectrum level calibration data to remove systematic bias from the detector, the corrector optical assembly, fibers and so on. As the target light are dispersed and projected onto the CCD, spatial locations or “traces” of fibers in the CCD pixel x-y positions need to be measured. This is obtained with continuum lamp exposures. The mapping of wavelength to the respective y (dispersion direction) position, also called wavelength calibration, in a spectrum are obtained by analysis of an arc lamp image, for which the line positions are known. To measure the smearing of light in the detector, the offline adopts a robust PSF modeling of the arc lines using “spectroperfectionism” ([16]), which uses computationally expensive 2D PSF modeling of the arc lines from where the PSF resolution at the line position is computed. Once PSF models are fitted at the arc line positions, a polynomial expansion of the measured PSFs is performed for each fiber so as to model an interpolation of PSF at any spot along the fiber. The basic idea of resolution can be understood in the following way. If $R$ represent the resolution matrix for a particular fiber in a defined wavelength grid, then the observed spectrum ($f$) for that fiber can be understood as the convolution of the $R$ matrix with the natural resolution spectrum,

$$f = R.t$$

where $t$ represent the true natural resolution target spectrum. This also suggests that deconvolution of $f$ with $R$ yields back the original spectrum $t$. In practice, the deconvolution operations are performed globally as one solution for all the fibers considering relevant covariances in a multivariate problem. See Ref. [16] for the mathematical framework and discussion. In the DESI spectroscopic pipeline, convolution and deconvolution operations will be performed in several instances as you will see below. Thus information regarding the resolutions are needed ahead of science data processing. I’ll describe the algorithms to calculate effective resolutions in the context of QuickLook in the next chapter.
DESI will have to perform cross-fiber calibrations as some fractions of the fibers will be particularly assigned to blank sky positions and standard stars. So any fiber specific systematics has to be corrected to perform calibrations across fibers. The fiber specific non uniformities across the fibers are corrected using smooth featureless continuum spectra. The continuum lamp 1D spectra are deconvolved by the respective resolutions in one global solution to yield a mean native resolution continuum spectrum. The reconvolution of the mean spectrum with respective fiber resolution matrices yield a new set of continuum spectra. Dividing the original pre-deconvolution spectra by the new set bin by bin for all fibers yield a set of correction vectors called “fiber-flat” vectors. The science spectra will be corrected for such systematics from fibers using these vectors.

7.2.1.1. Preprocessing:

DESI maintains a specific well-defined data model for the form of data after each stage of the data processing. The raw science data are first subjected to a preprocessing stage to remove 2D level systematics. This includes the 2D pixel level calibration of the images to account for the bias and dark current subtraction and flat correction that corrects for disproportionality of light in the CCD plane due to vignetting etc. Preprocessing also includes the removal of cosmic rays as much as possible. 2D calibration data are taken before the sky observing begins.

7.2.1.2. Spectral extraction:

Spectral extraction converts the 2D image level counts to 1D spectra for each fiber. It thus requires the CCD pixel level x-y mapping to fiber-wavelength mapping and also the PSF resolution model. This is achieved from previously analyzed continuum and arc lamp exposures. Based on the PSF resolution at any position along a fiber projection on the CCD, integration of the flux from the pixels and sub-pixels under the model PSF at a given spot yield the final extracted flux value for that position. This is repeated at desired wavelength grids for all fibers to obtain full 500 1D spectra from the image.
7.2.1.3. Fiber flattening:

The extracted 1D science spectra can still be impacted by any disproportionality from individual fibers. To correct for such systematics, each science spectrum is divided out by the respective fiber-flat vector obtained earlier from the continuum lamp analysis. One key feature of spectroperfectionism is that the extracted 1D spectra have no covariances whatsoever among the wavelength bins or across fibers. Same grid is chosen for the data processing throughout, allowing a natural bin by bin division for the fiber-flat correction. This also has a merit of easy error propagation after the operation.

7.2.1.4. Sky Subtraction:

A lot of the DESI primary targets will have very low S/N of the spectral features. Spectra will be fully dominated of high sky background lines and photometric noise. Extracting galactic features at the noise level requires a robust and sophisticated background subtraction algorithm. For this, DESI uses the spectroperfectionistic approach, as in the previous cases. First, a deconvolution of the fiber-flattened spectra is performed on the sky fibers using their respective resolution matrices. This results in a deconvolved mean spectrum, an effective sky spectrum at native resolution of sky lines. This mean sky spectrum is then reconvolved with resolution matrices of all fibers, to produce the sky model spectra for all fibers. As before, the sky model spectra are then subtracted from the science spectra bin by bin.

7.2.1.5. Flux Calibration:

The photons from the target after the sky subtraction have to be converted to physical units. For this the best matching stellar spectra for each standard star is chosen from a sea of templates, which is then scaled spectro-photometrically to the imaging magnitudes. In 2015, I wrote the early version of best match template fitting code and spectro-photometric normalization of the template model to the observed magnitudes from imaging data. Consider a sky subtracted 1D spectrum of a standard star in a DESI fiber, that has an apparent magnitude \((m_\lambda)\) obtained from imaging survey in some photometric band with filter response
function $R_\lambda$. Subscript $\lambda$ is the index corresponding to a specific photometric band. Before performing the template fitting, all three channels are first normalized by dividing out by the continuum and combined to a common wavelength grid. Continuum normalization is performed for the theoretical templates as well, which is followed by resampling in the same wavelength grid as data. A $\chi^2$ based statistical analysis is performed against hundreds of stellar templates and the model that yields the minimum $\chi^2$ is selected as the best fit template model $t(\lambda)$. The best model rest-frame magnitude ($m_{\text{rest}}$) is calculated by convolving with $R_\lambda$ using

$$m_{\text{rest},\lambda} = \int R_\lambda(\lambda)t(\lambda)d\lambda$$

This rest-frame template magnitude is then normalized to $m_\lambda$ of the star to obtain a scale factor,

$$\text{scale factor} = 10^{(m_{\text{rest},\lambda} - m_\lambda)/2.5}$$

This scale factor is the ratio of fluxes in the rest frame physical units divided by counts from the observation for the star. Therefore multiplying the observed spectral photons with this scale factor converts the observed photons to physical fluxes, i.e., the flux calibrated spectra for the standard stars. A standard star flux calibrated spectrum and the best fit template model after normalization is shown in Fig. 7.7. The DESI $brz$ channels are shown in blue, red and green colors and the high resolution best fit template model is shown underneath in black. Note that all the templates were sampled to wavelength grid as data before fitting. Many absorption lines match and the continuum also matches quite well. Few features at the channel edges are observed in the DESI spectra potentially due to edge binning.

This software has been updated to account for the bias from telluric atmospheric line subtraction artifacts, the peculiar motions of the stars, and the extinctions from the galactic dust in the most recent release. Current update also includes an interpolation in multi-parameter space of astrophysical parameters such as effective temperature, specific gravity, [Fe/H] ratio etc. to tune the final best fit model. As you may note, for each DESI exposure,
Figure 7.7. Flux calibrated standard star DESI spectrum showing all three channels in blue, red and green. Underneath lies the best matching high resolution template spectrum in black.
there will be dedicated fibers for multiple standard stars for flux calibration of all spectra in the exposure. These will be used to construct as many calibration scales which will have dispersion, mostly affected by systematics from fiber resolutions and photometric uncertainties. For the optimal solution, final flux calibration is obtained following spectroperfectionism like in the previous steps. First, a mean calibration vector is obtained by deconvolving each standard star spectra with respective fiber resolution in one global solution, and then further convolved with all fiber resolution to obtain the calibration vectors for each fiber. The extracted 1D photon level spectra are multiplied by these scaling vectors to yield the final flux calibrated spectra.

7.2.1.6. Redshift and classification

After the flux calibration, the redshift of all the targets are to be measured to build the final target catalogs for science analyses. A PCA based template fitting redshift code\(^4\) is used for redshift measurement of the flux calibrated spectra. For each target, a best fit redshift is returned, with the respective $\chi^2$. In addition, a warning flag $ZWARN$ is also returned as a metric that gives the confidence of a successful redshift and target classification is obtained from the model templates. A $ZWARN$ value of 0 is assigned if the confidence level is high. The confidence level is based on the $\Delta \chi^2$ value of the best match and the second best match and so on. From the tests based on the data challenge simulations, the redshift success rate surpasses the requirements for most targets. Occasionally, the redshift returned is not correct even though the $ZWARN$ flag is 0. In the simulations, these are identified as catastrophic failures and their fraction is at the subpercent to percent level. In real data, there is no way of knowing these failures unless a detailed individual spectral analysis is performed. So it is very important to test and improve the redshift fitting to keep these failure rates as small as possible. PCA templates and the fitting software is under continuous development to improve the success rate with increased realism in the simulations.

\(^4\)https://github.com/desihub/redrock
After the redshift fitting and target classification are performed with the spectroscopic pipeline, the targets are assembled to make a large scale structure catalog for science analyses.

With this description of the hardware and software components, I now give a brief description of the DESI survey and the precursor imaging surveys that are accumulating the DESI targets.

7.3. Survey overview

To produce an unprecedented 3D map of the universe, DESI will trace the matter distribution in the universe with a set of tracers comprised of the emission line galaxies (ELG), luminous red galaxies (LRG), quasars (QSO) Lyman $\alpha$ QSO, and bright galaxies (BGS). It will also observe about $\sim$10 million milky way stars (MWS) during the bright sky times. This targeting will amass more than 30 million galaxy spectra, many of which, particularly faint ones, will be observed multiple times. Representative rest-frame spectra of the DESI target galaxies and their corresponding representative wedge map in the redshift/distance space is shown in Fig. 7.8. A summary of the target densities of each target class is given in Table 7.1, taken from Table 3.1 of DESI final design report ([35]).

<table>
<thead>
<tr>
<th>Galaxy type</th>
<th>Redshift range</th>
<th>Bands used</th>
<th>Targets per deg$^2$</th>
<th>Exposures per deg$^2$</th>
<th>Good z’s per deg$^2$</th>
<th>Baseline Sample</th>
</tr>
</thead>
<tbody>
<tr>
<td>LRG</td>
<td>0.4-1.0</td>
<td>$r, z, W1$</td>
<td>350</td>
<td>580</td>
<td>285</td>
<td>4.0 M</td>
</tr>
<tr>
<td>ELG</td>
<td>0.6-1.6</td>
<td>$g, r, z$</td>
<td>2400</td>
<td>1870</td>
<td>1220</td>
<td>17.1 M</td>
</tr>
<tr>
<td>QSO (tracers)</td>
<td>$&lt; 2.1$</td>
<td>$g, r, z, W1, W2$</td>
<td>170</td>
<td>170</td>
<td>120</td>
<td>1.7 M</td>
</tr>
<tr>
<td>QSO (Ly-α)</td>
<td>$&gt; 2.1$</td>
<td>$g, r, z, W1, W2$</td>
<td>90</td>
<td>250</td>
<td>50</td>
<td>0.7 M</td>
</tr>
<tr>
<td><strong>Total in dark time</strong></td>
<td></td>
<td></td>
<td>3010</td>
<td>2870</td>
<td>1675</td>
<td>23.6 M</td>
</tr>
<tr>
<td>BGS</td>
<td>0.05-0.4</td>
<td>$r$</td>
<td>700</td>
<td>700</td>
<td>700</td>
<td>9.8 M</td>
</tr>
<tr>
<td><strong>Total in bright time</strong></td>
<td></td>
<td></td>
<td>700</td>
<td>700</td>
<td>700</td>
<td>9.8 M</td>
</tr>
</tbody>
</table>

Table 7.1. Properties of different DESI target class.
Figure 7.8. Top: Representative target galaxy rest frame spectra of DESI target galaxies. Bottom: A representative wedge map of DESI targets with redshift (distance).
7.3.1. DESI targeting and imaging Surveys

DESI targets come from the precursor imaging surveys, some of which are completed, while others are close to finishing. The targets will be selected based on optical $grz$-band photometry from ground and near-infrared photometry using $WISE W1$, and $W2$ from space. DESI imaging are expected to provide a $5\sigma$ lower limit of depths at AB magnitudes of 24.0, 23.4, and 22.5 in $g$, $r$, $z$ bands and 20.0 and 19.3 in $WISE W1$ and $W2$ bands.

The optical imaging is carried out with three telescopes. The Dark Energy Camera Legacy Survey (DECaLS), using the Blanco 4-m telescope in Cerro Tololo, Chile; will provide $grz$ photometry for over 9000 deg$^2$ at Dec. $\leq +34$ deg in the DESI footprint. The Bok 2.3-m telescope at Kitt Peak will provide $gr$ imaging over 5000 deg$^2$ at Dec. $\geq +34$ deg with the Beijing-Arizona Sky Survey (BASS), while the Mayall 4-m telescope on the same site provides $z$ band imaging in the same footprint using Mosaic $z$-band Legacy Survey (MzLS). I performed 15-full nights observing of the MzLS survey in early 2017. This observing involved obtaining calibration frames for the night, and running of the scheduler for the night and image reduction pipelines and visually monitoring the data quality with frequent intervention on the exposure times and control given the observing conditions. After completing the whole night imaging, and resting the telescope to its park position in the morning, several scripts are run to provide the summary of data taking, and progress of imaging survey to the collaboration. MzLS imaging is soon completing as the telescope should be prepared for DESI installation. The legacy survey recently released their data release DR5 involving the BASS, DECaLS, and MzLS data in October, 2017. Above surveys will be complemented both in the north galactic cap (NGC) and south galactic cap (SGC) with extended programs such as DECaLS+ and DES. Fig. 7.9 shows the sky map of the imaging surveys which will provide the targets for DESI. $WISE$ all sky survey provides infrared imaging that will be critical for target selection of LRGs and QSOs. A summary of imaging surveys for DESI targets is shown in Table 7.2, taken from Table 3.2 from Ref. [35].

Table 7.2. Summary of precursor imaging surveys for DESI, taken from Ref. [36]

<table>
<thead>
<tr>
<th>Telescope/survey</th>
<th>Bands</th>
<th>Area deg²</th>
<th>Location</th>
<th>Status</th>
</tr>
</thead>
<tbody>
<tr>
<td>Blanco DECam (DECaLS)</td>
<td>$g, r, z$</td>
<td>9K</td>
<td>NGC+SGC (Dec. $\leq +34$ deg)</td>
<td>Begun 2014</td>
</tr>
<tr>
<td>Bok 90Prime (BASS)</td>
<td>$g, r$</td>
<td>5K</td>
<td>NGC (Dec. $\geq +34$ deg)</td>
<td>Begun 2015</td>
</tr>
<tr>
<td>Mayall MOSAIC-3 (MzLS)</td>
<td>$z$</td>
<td>5K</td>
<td>NGC (Dec. $\geq +34$ deg)</td>
<td>Begun 2016</td>
</tr>
<tr>
<td>WISE-W1</td>
<td>3.4 $\mu$m</td>
<td>all sky</td>
<td>all sky</td>
<td>Completed</td>
</tr>
<tr>
<td>WISE-W2</td>
<td>4.6 $\mu$m</td>
<td>all sky</td>
<td>all sky</td>
<td>Completed</td>
</tr>
</tbody>
</table>

Figure 7.9. Primary imaging surveys providing the targets for the DESI spectroscopic survey. NGC targets at Dec. $\leq +34^\circ$ will be obtained with DECaLS, while NGC targets at DEC $\geq +34^\circ$ will be obtained from BASS ($g$ and $r$-bands) and MzLS($z$-band). SGC targets will be obtained with the DECaLS, DES and DECaLS+. Figure is taken from Ref. [35].
7.3.1.1. Emission line galaxies

ELGs provide the largest target sample for DESI. DESI will obtain \( \sim 17 \) million ELG spectra over the survey time in the redshift range 0.6–1.6. ELGs constitute spiral and irregular galaxies that possess high star formation rate, due to which they typically show bluer rest-frame colors compared to galaxies comprising of old stellar populations such as LRGs. Spectral properties of ELGs show strong nebular emission lines. The observation that star formation rate at \( z \sim 1 \) was higher than today is an advantage for DESI ELG redshift range. A typical rest frame spectrum of an ELG is shown in Fig. 7.10. The spectrum constitutes a blue continuum and is dominated by numerous nebular lines including Balmer lines such as \( \text{H} \alpha \lambda 6563 \) and \( \text{H} \beta \lambda 4861 \), and forbidden lines of \([\text{OII}] \lambda 3727,3729\) doublet. DESI instrument design is expected to resolve the [OII] doublet for a line width of 70 km s\(^{-1}\). This line is critical to DESI ELG sample and a successful redshift is expected for [OII] luminosity higher than \( 8 \times 10^{-17} \) erg s\(^{-1}\) cm\(^{-2}\).

Current ELG target selection algorithm employs a \( g - r \) vs \( r - z \) color-color space cuts from previous deeper surveys normalized to legacy survey sensitivity. As shown in Fig. 7.11, the selection scheme, represented by region bounded by black lines, does good isolation of the ELG population in the redshift range 0.6 \(< z < 1.6\) for strong [OII] emitting galaxies. Selection of galaxies to a depth such as \( r \) magnitude of 23.4 further minimizes the contamination from stars and other galaxies, while meeting the DESI science requirements. Given the selection scheme using \( grz \) photometry, the expected DESI detection of [OII] flux is at \( > 7 \sigma \), which secures a successful redshift. However, the target selection can be affected by several factors such as galaxies at \( z > 1.63 \), where no emission lines can be detected, or low redshift galaxies \( (z < 0.6) \) or stars or QSOs. Regardless of these factors, target selection yields about 80% success rate for galaxies in the correct redshift range.
Figure 7.10. A representative rest-frame ELG spectrum showing the Balmer emission lines and other nebular emission lines from [OII] and [OIII]. In the inset is zoomed [OII] doublet. This is one of the most reliable lines for DESI ELG sample. Overplotted are the filter response functions for DECaLS $grz$ bands. Figure is taken from Ref. [35].
Figure 7.11. A $g - r$ vs $r - z$ color based selection scheme for ELG targets. The objects shown are obtained from DEEP2 survey, but normalized to DESI depth and sensitivity. The region bounded by black line yields the targets that meet the DESI requirements.
7.4. DESI projections

The current state of art and upcoming BAO experiments, both spectroscopic (e.g., Baryon Oscillation Spectroscopic Survey (BOSS), extended BOSS (eBOSS), Hobby Eberly Telescope Dark Energy Experiment (HETDEX), Wide Field Infra-Red Space Telescope (WFIRST)), and imaging (e.g., Dark Energy Survey (DES), Large Synoptic Survey Telescope (LSST)), will provide synergies for different science cases. However, DESI alone is expected to achieve an unprecedented cosmological precision over a wide range of redshifts. Experiments such as Euclid and WFIRST can achieve competitive precision of the BAO distance scale but they will be limited to a narrower redshift range. In contrast, DESI will map the full expansion history of the universe through redshift of 3.5. A Fisher matrix forecast of the BAO scale error after normalizing the survey dependencies for various surveys is shown on the left plot of Fig. 7.12, taken from Ref. [36]. The right plot shows the projected DESI measurement sensitivity and the expansion rate of the universe. A representative current plot from recent surveys was shown in Fig. 3.4.

In this chapter, I set up a stage showing how the raw data from spectrographs will be used to create a galaxy catalog for the final large scale structure analysis. In the next chapter, I will present my work on spectroscopic simulations and the online data reduction pipeline for DESI. These simulations and pipeline performance will be used in the final galaxy clustering analysis in Chapter 9.
Figure 7.12. Left: Forecast of BAO scale error for DESI and other surveys. DESI will obtain the precise BAO measurements up to the redshift of $\sim 3.5$. Right: Projected DESI measurement of the expansion rate. Plots are taken from Ref. [36].
Raw data level simulations are critical both for the data processing software development and to perform realistic predictions to prepare as much as possible before the real observation begins. DESI online raw data processing will be critical to guarantee science analysis level data quality and to dynamically maximize the overall survey efficiency. In this chapter, I describe my contribution to the online data reduction pipeline development and analyses involving spectroscopic simulations. The simulations described here will facilitate my final large scale structure analysis in Chapter 9.

8.1. Spectroscopic simulations

DESI is a spectroscopic instrument. The spectroscopic simulation tools developed by the data systems team facilitates development of the softwares of various kinds, that allows one to understand the hardware and software responses and efficiencies to optimize various elements of the experiment. Within the DESI data systems, both raw pixel level 2D image simulation and a less extensive spectrum level 1D simulation have been developed. 2D simulation serves as the most accurate representation of raw data exploiting the full dimensionality of CCD space and can be processed with the full spectroscopic pipeline. The 1D simulation, however, facilitates faster outcomes bypassing the CCD pixel level imaging and further extracting 1D spectra. 2D image simulations are important to study the variations in the dispersion and cross-dispersion directions, fiber tracing, cross-fiber responses and the extraction efficiencies etc across the CCD plane, but post extraction simulations are confined to only 1D spectral processing need. An order of magnitude faster speed of 1D simulator merits over 2D pixel level simulations for several analyses in the data challenges where the need of pixel level data is less important. So the DESI data system has made extensive use of both the simulators,
which I describe below.

A spectroscopic simulation employs a native resolution template galaxy spectrum at a particular redshift, and pre-defined astrophysical priors, such as magnitude, [OII] luminosity or different line ratios, SN contamination etc. Simulations of an ensemble of such targets requires selection of template spectra accounting realistic distribution of astrophysical parameters. This is facilitated by Monte Carlo sampling off of realistic distributions of such properties within DESI spectroscopic simulation software\(^1\).

Once a template spectrum is selected with desired astrophysical properties, it is convolved with a high resolution sky spectrum model. This is then subjected to convolution of atmospheric losses and observational conditions followed by the instrumental throughputs and loss models of the DESI fibers.

For the 2D image simulator\(^2\), the resulting spectrum is projected on to the CCD using a model PSF defining the detector resolution and the fiber tracing model that specifies a particular wavelength for a particular fiber landing on a particular pixel on the CCD. The outcome is a DESI 2D image with the spectra superimposed on top. In the CCD image, the spectral photons are converted to the ADC counts using model amplifier gains, also accounting the read-out noise. Such a simulated image is effectively equivalent to raw DESI image as if obtained from the spectrograph during observation. This image is expected to be processed through the pipeline to yield the final 1D spectra, from which redshifts could be measured. Additional realism such as random cosmic ray hits etc are also incorporated at earlier stages within the simulation.

For the 1D spectral simulation\(^3\), the inputs before the light hits the spectrograph are same as in the 2D case. The 1D simulator uses a model PSF resolution but unlike 2D simulator, it does not employ projection of light on to the CCD. The simulation is completely carried out

\(^1\)https://github.com/desihub/desisim

\(^2\)https://github.com/desihub/specter

\(^3\)https://github.com/desihub/specsim
in 1D, i.e., a high resolution template spectrum is converted to DESI output like spectrum, bypassing the 2D level image simulation and extraction. The use of the fast 1D spectral simulation for science analysis warrants the accuracy validation of 1D simulation with the more accurate 2D simulations. In 2015, I performed a study for cross validating the 1D simulation output with the 2D extracted spectrum. Few issues were identified and corrected such that the resulting outputs were on the same footing. A high level of agreement was observed on the final outputs from 2D and 1D simulations.

While a preliminary data reduction pipeline for 2D raw pixel data processing existed as offline pipeline, there was no such a reduction pipeline integrated for 1D simulated output files following the DESI standard data model. The need for such a pipeline was highly demanded as 2D simulation became expensive for several analyses in the data challenge and focussed studies elsewhere. I wrote an integrated pipeline⁴ for the 1D simulation in 2015. This would take the same inputs as 2D simulator for the templates, observing conditions, redshift distribution etc, but would wrap the 1D simulator module and write out the DESI standard post extraction files as if they were created by offline data reduction pipeline using the raw DESI images. This 1D simulation pipeline was used for several data challenges and systematic studies of various kinds. The pipeline was further improved to add more flexibility in the input options from the command line. Based on this pipeline, I performed an analysis studying the systematic effects on the redshift efficiencies of galaxies due to varying observing conditions. This study will be presented in detail in Chapter 9.

DESI also uses simulations at the large scale structure catalog level to perform galaxy clustering analysis. This uses a model of spectroscopic pipeline performance of the final redshift measurement of target spectra. End to end processing of individual galaxy in a catalog of millions through the projection of spectra on to raw images followed by the data reduction to yield the final redshift is very expensive and time-consuming. Thus, based on the 1D spectroscopic simulations of few thousands of galaxies by varying observational conditions

in the input, I wrote an initial redshift efficiency model for the spectroscopic performance. Quick analyses of large scale structure using millions of targets without the need of image level 2D or spectrum level 1D simulations followed by the data processing was highly desirable for many data challenge studies. An accurate modeling of the spectroscopic efficiency would suffice several DESI sensitive cosmological analyses from the galaxy distribution directly. Given a target object’s sky positions, redshift and the observational conditions it is likely going to be observed under, this efficiency model assigned a redshift and uncertainties that would be similar to the outcome of the full data reduction pipeline if the object was really observed. I will present detail analysis of galaxy clustering using such a model in Chapter 9. I now move to the quicklook data reduction pipeline that I engaged thoroughly during my PhD.

8.2. The QuickLook pipeline

To ensure that the spectroscopic data obtained bears enough S/N for DESI science analyses, the raw data coming out of the spectrograph from each exposure needs proper monitoring for data quality in quasi-real time. This is done with the online data processing QuickLook (QL) pipeline. As the offline processing is costly for real time data quality checks, QL has implemented faster yet robust processing algorithms, defining several data quality (QA) metrics that the user can monitor from the online system. With the output QA metrics, QL also provides feedback to the online system so that the exposure length for the next observation can be calibrated.

QL includes a similar set of reduction steps as the offline data reduction pipeline but adopts a self-contained package to maintain consistency of fast reduction algorithms. Although QL is modular to switch between online and offline algorithms, offline algorithms are mostly used there for consistency checks of the results from the faster QL algorithms. QL interfaces with the DESI online system and the web and database framework, called Quicklook Framework (QLF). I extensively contributed to the integrated pipeline, configuration, pipeline algorithms (PA), data quality QA metrics, data model and documentation,
and interfaces to offline and QLF tools. A graphical scheme of QL architecture is shown in Fig. 8.2.

QL pipeline integrates multiple independent components facilitating maximum modularity of the code. While the objective of each PA is specific, the code is designed such that the pipeline steps are managed by a Python base class. The same is true for the QAs. The integration of these independent modules are done within the pipeline by direct call of the functions or through API interfaces. Below I describe the main components of the QL data processing, as raw data is obtained for a given exposure.

8.2.1. Data processing pipeline algorithms

A pipeline algorithm (PA) refers to a particular processing step of the pipeline, that does a specific task on the data obtained. Depending on the nature of the data: calibration or science, a data processing pipeline is set up with relevant PA tasks. After each PA, a set of QAs run that analyze the PA output and produce metrics of various forms for monitoring the quality of the data obtained. In this section I’ll describe the PAs adopted in the QL.

8.2.1.1. Preprocessing calibration

A single exposure of DESI science data is structured in a compact raw data FITS file, that exhibits data of all 10 spectrographs and all channels. Each raw file bears unique exposure ID for the given observation. The raw DESI file for science exposures also includes the fibermap, a dictionary that maps a fiber to a target and contains photometric and other information of the targets. The preprocessing step is performed on the raw images. It constitutes of bias subtraction, that corrects for the detector zero point; followed by dark subtraction, which accounts for the dark current in the readout electronics. The last processing does the 2D pixel flattening, which is the pixel by pixel flattening to correct for any vignetting in the image. Each of the bias, dark and pixel flat is a separate image taken earlier before on sky science observation begins. The preprocessing is done at individual camera level, i.e a total of 30 images are resulted from one exposure. The preprocessing PA module adopts the offline
Figure 8.1. Design of DESI QL pipeline. QL processing is done for one camera at a time. A QL pipeline for a given camera for a given exposure exhibits a layer of pipeline processing PA algorithms, each of which can have several QA metrics.
preprocessing functions and interfaces with DESI Input/Output (I/O) modules to produce
the preprocessed image for each camera. This image also consists of the 2D array of inverse
variance values for the pixels accounting for the Poisson error and the read noise.

8.2.1.2. Spectral extraction

One of the key differences of QL with offline is the extraction, as QL needs a faster, yet
robust algorithm to extract photon level spectral features as much as possible. Currently,
QL uses a row by row box car extraction of the spectrum from the 2D CCD pixels. A PSF
calibration input obtained from arc lamp processing (see Section 8.2.2) is supplied to map
the CCD $x - y$ pixel space to fiber-wavelength space. Each fiber is extracted by placing a
box of specified window at the trace centroid aligning in the cross dispersion direction in the
preprocessed CCD image, and integrating the flux within. For a given wavelength, mapping
to a particular pixel, the integrated flux inside a box is given by

$$ f_\lambda = \sum_{i=-hw}^{+hw} w_i p_i $$

(8.1)

where $hw$ is the half-width of the box chosen and is measured from the PSF’s centroid, $w_i$
is the weight defined as the fraction of the overlap of the box the horizontal pixels, and $p_i$
is the corresponding pixel count. The variance corresponding to the integrated fluxes are
calculated from the variance values in the image accounting the fractional overlap of the box
with the pixels and treating uncorrelated errors of the pixels.

$$ \sigma_\lambda^2 = \sum_{i=-hw}^{+hw} w_i \sigma_i^2 $$

(8.2)

where $\sigma_i^2$ is the variance of the $i^{th}$ pixel, obtained by inverting the inverse variance values
given in the preprocessed image. The variance $\sigma_\lambda^2$ is then inverted to yield the inverse variance
 corresponing to the particular box.
Once the integrated flux and inverse variance are evaluated for each row along a fiber, this spectrum is resampled to a common wavelength grid as desired for the full image extraction and same is done for all fibers iteratively. Note that DESI uses a single 1D extraction uniform wavelength grid for all fibers. A representative 1D spectrum using boxcar extraction and that with the offline extraction are shown in the top panel of Fig. 8.2. From the wider spectral features shown in the inset on the top panel, the boxcar extraction exhibits slightly poorer resolution. On the bottom panel are shown the errors obtained by inverting and taking square root of the inverse variance. Here also, the boxcar extraction show higher errors than offline extractions. This is the limit of the boxcar extraction that the resulting S/N will be lower than the more robust extraction methods. Since there is no modeling of the spectral features, all the pixels fully enclosed get the same weight regardless of the counts. This propagates maximum readout noise to the output spectra. Boxcar errors can also be dominated by the readout noise from the neighboring pixels when the central spectral features have very low flux level. In spite of these disadvantages, the level of variations observed is small, and the fast reduction approach is expected to fulfil the online data processing needs.

The final result of this process are recorded as 500 1D flux and 500 1D inverse variance vectors for the 500 fibers in the output “frame” Python object. A frame object is a DESI data model convention for 1D level spectra for a given camera, as they are extracted from a 2D image. Finally a resolution matrix is constructed using the model coefficients obtained from arc lamp analysis (see below for arc processing and derivation of these coefficients). The extracted 1D spectra, their inverse variances and the respective resolution data are stored in the frame object.

8.2.1.3. Fiber flattening

Fiber flattening is the 1D level post extraction calibration to correct for fiber to fiber variations. Given the variation in fiber to fiber resolution, QL uses the offline technique of deconvolution of the continuum 1D spectra over all fibers, computes mean, and reconvolves the mean with the fiber resolution. Dividing out the extracted 1D continuum spectra with
Figure 8.2. Top: Comparison of the boxcar extraction with the offline extraction of a representative spectrum. Boxcar extracted features are wider than the offline extractions as shown by the zoomed in region in the inset. Bottom: Same as top panel but for the errors, where boxcar errors are larger than the offline errors.
the respective reconvolved spectra gives the correction vector, to be used for calibrating the science 1D spectra. Unlike the offline fiber flat, QL does not perform any outlier rejection and smoothing of spectra. Also, it ignores the covariances between fibers and deconvolution is done in individual fiber level, as opposed to one global solution. A mean vector is calculated after the deconvolved fiber spectra are obtained. Convolution with respective fiber resolutions give the respective fiberflat vectors.

8.2.1.4. Sky subtraction

QL adopts offline algorithms for the sky subtraction of science spectra. As in the offline case described in Section 7.2.1, the extracted, fiber-flat corrected sky fibers are deconvolved to yield an effective fiber-resolution independent mean sky spectrum. This mean spectrum is then reconvolved with resolution for all fibers to create a sky model for each fiber. The spectra in the frame are then sky-subtracted bin by bin. A simple inverse variance weighted average sky model computation is also available within QL.

8.2.2. Arc processing

Arc lamp exposures are required for determining the detector resolution and the wavelength calibration for each fiber. These serve as calibration inputs for the extraction process. In QL, arc processing involves Gaussian modeling of the selected arc line PSF. Offline uses more extensive modeling of core and wings in 2D space, which is computationally too expensive for QL. With the Gaussian modeling, symmetry is assumed in the PSF shape, and non-Gaussian tail components are assumed to be negligible. After the PSF sigmas ($\sigma_{\text{meas}}$) and the respective means ($\mu_{\text{meas}}$) are measured for the arc lines, a low order polynomial fit is performed to interpolate these $\sigma_{\text{meas}}$ for each fiber along the dispersion direction using the expansion

$$\sigma(\lambda) = \sum_{l=0}^{n} c_l P_l(\lambda)$$

(8.3)
where $P_l(\lambda)$ is the Legendre polynomial of $\lambda$, $n$ determines the order of expansion, and $c$’s are the expansion coefficients. A least square fitting of $\sigma_{\text{meas}}$ at $\mu_{\text{meas}}$ yields the best fit solution of these expansion coefficients. The domain for the fit is defined to be the wavelength values corresponding to extreme pixel edges for each fiber. The interpolated sigmas obtained from the expansion models are used to fill in the resolution matrices for the entire frame. What this means is that given the fitted sigma ($\sigma_{\text{fit}}$) at a given wavelength for a given fiber, a normalized Gaussian with $\sigma_{\text{fit}}$ is sampled to yield the probabilities at the neighboring wave bins. This yields an array of values for a given fiber at a given wavelength as a measure of detector resolution. These arrays make the band diagonal values of what is called a resolution matrix. The band width is pre-defined from the expected PSF extension from the centroid and chosen symmetric. Early testing of the spectrographs have shown that at 4 pixels away from the PSF center, the contribution from the feature is at 0.4% level. The dimension of the resolution matrix is defined by the extraction scheme of 1D spectra, i.e., a resolution matrix for a fiber is $n \times n$ if the spectrum is of dimension $n$, with the corresponding band diagonal values coming from the Gaussian probabilities corresponding to a particular fiber at a particular extraction wavelength bin. The values beyond the pre-defined band diagonal range are all zeros reflecting the fact that far away pixels have no contribution to the PSF. Fig. 8.2.2 shows the fitted Gaussians of four of the lines from an arc lamp exposure on the left panel. The right panel shows the respective polynomial fits (vertically shifted for clarity) for several fibers. Overlaid are the respective offline effective Gaussian sigmas across those fibers from the simulated PSF model. I also performed tests with spectrograph test stand data available and the Gaussian extracted PSF sigmas were consistent with the offline effective Gaussian sigmas.

The final outcome of the arc processing is a Python object (“PSF” object) that stores the previous trace and wavelength calibration information along with the best fit expansion coefficients for all the 500 fibers for a given camera. I wrote a “PSF” Python class to load the information such as wavelength-fiber to CCD $x - y$ map etc. The use of this module is pervasive within the QL software.
With the above approach, the extracted 1D spectrum \( f_i \) for the \( i^{th} \) fiber can be modeled as matrix dot product,

\[
    f_i = R_i \cdot t_i
\]

(8.4)

where \( R_i \) is the resolution matrix for the \( i^{th} \) fiber and \( t_i \) is the effective fiber-resolution unconvolved spectrum vector.

Following the DESI data model, the frame outputs from extraction and subsequent PAs store the full resolution data obtained from the given PSF file output from arc processing. Running 30 QL processes in parallel via QLF identified a bottleneck due to the size of the 30 respective frame files for a single exposure, which was using the offline data model. The frame data model was modified in late 2017 to incorporate that coefficients be saved to the files rather than the full resolution data array, while resolution matrices are internally created and consistently handled as before. This was a major improvement towards the requirement of QL processes to run through QLF. For a 0.8Å grid extraction, the file size was reduced from about 160 MB to 15 MB.

8.2.3. Data quality monitoring algorithms

QL has implemented several data quality QA metrics as the data are processed by the pipeline. These metrics are calculated typically after a data reduction step. For each PA, a set of specific QAs are developed to monitor the data quality and algorithm performance of that PA. The quality algorithms are at both the CCD image (2D) and fiber spectrum (1D) levels, depending on the respective data processing step. Example of 2D pixel level metrics include the bias and noise measurement across different amplifiers in the CCD, or the imaging region of the CCD. A metric from 1D spectra can be the photon counts for sky fibers or a particular spectral line strength or its PSF width. I wrote several QAs, and prepared their relevant static plots for display. Each algorithm outputs a dictionary that can have several metrics or layer of metrics calculated after the processing. An example result is a S/N plot after the sky subtraction on a dark time exposure as shown in Fig. 8.4. For each target type, the S/N is plotted against magnitude and a polynomial fitting is
Figure 8.3. Arc processing with QL. Top: Gaussian fits of four arc lines on native CCD grid boxcar extraction for fiber 250. Bottom: Legendre polynomial fits of the Gaussian fitted sigmas of arc lines for several fibers. Dashed lines are the effective Gaussian PSF sigmas from the simulation. The PSF sigmas from the polynomials for each fiber are used to sample Gaussians to construct the resolution matrix.
performed. QL also allows cross-exposure data quality checks. Such relative data quality will be monitored by cross checking the current metric value with a reference fiducial value for the metric. The QA results will be posted to the database for higher level time series data quality monitoring, and also to provide important feedback to the dynamic exposure time calculator for re-configuring the exposure time for the next observation. Below I list the major QAs that I developed or contributed to and give brief description of their roles.

**Bias From Overscan**: This is a 2D image level QA and is performed before preprocessing step on the raw CCD image. It measures the bias in the overscan region for the whole CCD and separately on each of the four amplifier zones of the overscan region.

**Get RMS**: This QA is performed after preprocessing and measures the rms and noise in the CCD image, both in the imaging region and the overscan region. For the science, arcs and continuum lamp exposures, these values will be higher for the imaging region and lower in the overscan region. The values are also calculated from each of the four amplifier regions of the CCD to monitor the overall amplifier performance.

**Count Pixels**: This also runs after the preprocessing step. It computes the number of CCD pixels above configured threshold for each amplifier zone and the whole CCD. Any substantial variation on the pixel counts should correlate with the number of dead columns in the arcs and continuum lamp exposures. Based on the threshold, the pixel count can vary significantly in a science exposure because of the wider variation in brightness of the target spectra.

**Calc_XWSigma**: This QA is also performed on the 2D image after the preprocessing step. Using the fiber traces to $x - y$ position on the CCD, it looks at the pre-specified strong sky lines and performs a 1D Gaussian fits on the $x$ (row or cross-dispersion) and $y$ (column or dispersion) direction of the line. The resulting outputs are the Gaussian sigmas and can be
Figure 8.4. QL QA static plot of median S/N vs magnitude on a science exposure after the sky subtraction. Top: 1D histogram showing median S/N of the targets. Tallest spikes correspond to the bright standard stars, and the gaps between the fibers correspond to the sky fibers. Bottom: S/N vs magnitude for all target types for the exposure. Lines overplotted are the polynomial fits. The coefficients of the polynomial expansion and the errors are propagated in the QA output dictionary.
checked for PSF variations across the CCD plane.

**Count_Spectral_Bins:** This runs after the extraction of spectra from the 2D image. For each fiber, it counts the number of wavelength bins that are higher than configured threshold. If the amplifier gains are similar, this number will be higher for brighter targets and sky fibers are expected to have similar number of wavelength bins in a science exposure. Significant variation across sky fibers can provide clue on the detector response across CCD, unexpected loss in fibers etc.

**Sky_Continuum:** This runs on the spectra after the fiber-flat correction. It measures the median counts from configured sky continuum region tens of Å wide in the sky fibers. This monitors the detector responses and loss in fibers.

**Sky_Peaks:** Sky Peaks also runs on the sky fibers after flat correction. It measures the integrated counts around specified sky lines. Line selection is done such that at least 1 line each land on both top and bottom half of the CCD, corresponding to separate amplifier regions.

**Sky_Residual:** Sky subtraction step involves subtraction of sky model corresponding to each fiber. This QA measures the distribution of residuals from the sky spectra and the corresponding models after the sky subtraction step. In general, in a good run and if the sky models are correct, the residuals should show Gaussian distribution. Deviation from Gaussianity can signify improper modeling or some underlying issue in some fibers.

**Integrate_Spec:** This QA also runs on the sky subtracted spectra. A section or the whole spectrum is integrated, so that each target can be cross-checked against the photometric magnitudes obtained from imaging data.
Calculate SNR: This QA runs on the sky subtracted science spectra. For each target, median S/N is measured. A polynomial fit of the distribution of the S/N with respect to the photometric magnitude is performed so that the data sensitivity can be directly checked. Based on this output, the ETC can determine the optimum exposure length for the next observation. A static plot output from this QA is shown in Fig. 8.4.

8.2.4. Configuration and pipeline set up

The raw data frames are processed with the QL pipeline following a specific configuration. Each exposure type, science or calibration, requires a separate pipeline set up, the PA and QA list and their respective parameters. This is needed because the purpose of such exposures is different and the metrics and their reference parameters will also be different. For example, for an arc exposure, the metric to monitor can be the PSF stability while for the science exposure, it can be S/N of a particular galaxy target. Also the reference fiducial values for the same QA metric for a dark time program exposure will be different from that of a bright time program. Such a configuration is supplied as an input static configuration file from the command line or a data model directory destination and specifies the respective exposure type (science, arc, flat, dark etc), their program (dark, gray or bright time), the respective calibration exposures, and the chain of the processing algorithms and parameters and respective QA chain and their parameters. A configuration module is written to load the configuration file to be used by the pipeline processing. The output file naming and their respective destinations at each processing step utilizes DESI standard data model and is handled internally within QL. The top level directory tree is set by environment variables by default but can be over-written from command line for stand-alone processing, development, debugging etc.

The QL pipeline infrastructure is general to all exposure types. It includes a series of data reduction steps, each of which exhibit a layer of data quality monitoring tasks, that produce various QA metrics. During the execution of the pipeline, the output from each
step is a standard DESI model object and is passed to next step in memory. If output from each step needs to be written to files, a True/False boolean in the configuration suffices that. During a data reduction step, each QA outputs are written as dictionaries to a static YAML\textsuperscript{5} output file and a static plot is generated with another boolean if configured properly. The metrics are also stored in memory for merging of all QA outputs after the pipeline execution, but production of a merged file is optional from the command line parsing.

The QL project development is in rapid progress. Testing, integration and performance benchmarking have been carried out substantially. Several analytic testing and validity for all types of data using realistic simulations can further optimize the software, its performance and help tune the fiducial parameters.

With this description on my contribution and research activity on simulation and QL development, I now move to cosmological analysis using simulated galaxy sample, specific for DESI, in the next chapter.

\textsuperscript{5}https://pypi.python.org/pypi/PyYAML
Chapter 9

OBSERVATIONAL SYSTEMATICS MEASUREMENT FOR ELG CLUSTERING

BAO measurement is a powerful tool to measure the expansion of the universe. I discussed the BAO physics and measurement tools from the galaxy distributions obtained with redshift surveys. In the configuration space, 2-point correlation function (2PCF) shows a peak at the BAO scale as shown in the left plot of Fig. 3.2. Measurement of such a feature at various redshift bins gives the expansion history of the universe, and thus constrain the parameters in the cosmological models. BAO technique has important advantage that it can simultaneously measure the angular diameter distance and the Hubble expansion from the observations transverse and along the line of sight. If $s \equiv 150 \text{ Mpc}$ represent the comoving distance, then the transverse $s$ ruler subtends an angle, $\theta$, giving angular diameter distance $D_A(z)$ via

$$s = (1 + z)D_A(z)\theta = \theta \int_0^z \frac{c dz'}{H(z')} \quad (9.1)$$

where spatial curvature is considered 0. Such a 2D correlation can be measured from the CMB anisotropy maps as well. But the spectroscopic surveys measure the positions of galaxies in 3D, thus allowing direct measurement of correlations along the line of sight as well. i.e., if $\Delta z$ is a separation of galaxy pairs along the line of sight, the correlation directly constrain Hubble expansion by,

$$s = \frac{c \Delta z}{H(z)} \quad (9.2)$$

Thus the cosmological measurements using BAO strongly rely on the accurate position, thus redshift, measurements of the galaxies. Translating the observed correlations from galaxy distribution to dark matter distribution is however a non-trivial process. This, in general, is parametrized in the form of “bias” relating the galaxy density field to underlying dark matter density field. Large cosmological simulations using models to connect galaxy
distribution to matter distribution (e.g., halo occupation distribution or hydrodynamic simulations) can fine tune the bias (e.g., [102]) . At the largest scales, the biases are suggested to be constant (e.g., [30]) and increasingly become non-linear on smaller scales. Also systematic effects from the survey such as tiling, fiber assignment etc. will significantly impact the measured galaxy distribution and thus the clustering measurement of observed galaxies. These effects need to propagated into cosmological simulations using the properties of the survey, e.g., tiling and fiber assignment strategies, radial selection functions of the targets, spectroscopic measurement efficiencies etc. These simulations before the survey effects are applied yield galaxy catalogs called “mock” catalogs that are envisaged to be affected in the same way as the raw data. The final analyses will involve generation of many realizations of such mocks combining the galaxy formation/evolution models and the statistical variances from the survey.

In this chapter, I present an analysis of the spectroscopic redshift measurement efficiencies considering realistic distributions of each of the observing condition and study the systematic effects on the BAO measurement for the dark time ELG simulated sample for DESI. The dynamic exposure time calculator attempts to accommodate for the non-ideal observing conditions across the exposures. However, the redshift measurement of individual target may still be impacted by these conditions and such an effect needs to be understood. This is the objective of the following analysis. Before I measure the ELG clustering on a mock catalog, I perform several pre-requisite analyses, whose results are needed to be propagated to the final analysis.

9.1. Parameters of observing conditions

The parameters considered in this study include airmass, seeing, and transparency during the observations. Although ELGs will be observed during the gray time survey as well, I am not considering the moon brightness for this study as I only focus on the dark time survey. However, galactic extinction is considered as an additional parameter, as it directly affects the exposure time and thus the redshift efficiency. Note that parameters such as
seeing, extinction, stellar density etc. will also have direct systematic effects on the imaging surveys, thereby the target selection. Those imaging systematics have to be separately calculated and propagated independently to the final systematic uncertainties. Here, the analysis is performed only for the redshift measurement efficiencies and thus only comprise spectroscopic observations. Several simulations are carried out to parametrize the redshift efficiencies for varying observing conditions as described below.

9.1.1. Survey simulation

Survey simulation imply a full simulation of the DESI survey designed to optimize the survey progress, schedule and prioritize fields each afternoon and model which tiles get observed under what observing conditions. To obtain a realistic distribution of each of the observational parameters, I ran a full 5 year baseline survey simulation. This included over 26000 exposures for 16000 tiles across the DESI footprint. The survey simulation relies on algorithms such as next field selector and exposure time calculator (ETC) to optimize the full survey efficiency. The observing time ephemerides, seeing, airmass, and transparency are propagated into the output files. Ephemerides are observing time information and thus calculated directly using the open source package astropy. Seeing and transparency are based on the time series models using the distribution from the imaging survey at the Mayall telescope. Airmass is obtained from the position of the target from zenith at the observing time. For this study, observational parameters are sampled from dark time exposures, which consist about 8000 tiles. Extinction values are coordinate dependent and the values are obtained from the median extinction values $E(B - V)$ provided for the tiles in the desimodel footprint.

The 1D histograms of the observing parameters sampled from the dark time of the survey simulation are shown in Fig. 9.1. The distributions of the parameters shown in Fig. 9.1 provide input to two steps in the following analyses. One, these observations will be inputs to the spectroscopic simulations where the redshift efficiencies will be trained-modeled as a function of these parameters. Two, these observations will be sampled and assigned to each
Figure 9.1. 1D histograms showing distributions of observational parameters: Airmass (a), seeing (b) and transparency (c), obtained from a full 5 year baseline DESI survey simulation.
object on a mock catalog, on which the final cosmological analysis will be performed. The efficiency model established from first case will be used to modify the redshifts of the target objects in the catalog depending on the observing conditions associated with each target.

9.1.2. Spectroscopic simulation

Observing conditions are randomly sampled from the observed distributions and passed into the 1D spectroscopic simulation pipeline described in Section 8.1. Five distinct configurations are separately simulated. The first case is where all the parameters are kept nominal, i.e. airmass=1 (corresponding to zenith pointing); seeing =1.1” (median seeing); transparency = 1 (100%), and $E(B-V)=0$ and a fixed exposure length of 1000 seconds. This is the nominal dark time exposure anticipated for gaining enough S/N for an average ELG spectrum for DESI. The remaining four cases involved random sampling of each of the parameters one at a time from the distribution obtained with the survey simulation above, while keeping the others nominal. In each case, the varied parameter will affect the ETC, and thus exposure lengths and finally the S/N in the spectrum. A total of 5000 ELG spectra are simulated for each case. The same random generator is used throughout to guarantee that the same truth template ELG spectra are selected each time and the astrophysical distribution of the templates remain same.

9.2. Redshift measurement

The set of 5000 spectra for each of the five cases are then passed into the redshift fitter and the output redshift efficiencies are analyzed absolutely and relative to the nominal case. Error on the measured redshift, catastrophic failures and errors on the statistical estimators are calculated using,
Redshift error, \( \Delta v = c \times \frac{z_{\text{measured}} - z_{\text{true}}}{1 + z_{\text{true}}} \)

Catastrophic failure, \( |\Delta v| > 1000 \text{ km/s} \)

Error in Median = \( 1.253 \times \frac{\text{RMS}(\Delta v)}{\sqrt{n}} \)

\( n = \) no. of points in a given bin

Error in RMS = \( \frac{\text{RMS}(\Delta v)}{\sqrt{2 \times (n - 1)}} \) for \( n \geq 10 \)

Fig. 9.2 show the redshift errors for the nominal scenario. On the top panel are shown the redshift errors with respect to true redshift, \( r \)-band magnitude and true [OII] flux for the ELG sample. Shaded regions indicate the 68% and 95% confidence regions for each bins. The bottom panel show the corresponding RMS distribution. The total efficiency for the nominal case was found to be 86.4%, with a catastrophic failure rate of 0.5%. I repeat the same analysis for each of the other four cases taking the same ELG templates, and varying observing parameters one at a time, and the result obtained is summarized in the Table 9.1. Note that the cumulative efficiency on all cases of parameter variation is slightly higher than the nominal case, possibly resulting due to somewhat over-adjustment of exposure lengths.

<table>
<thead>
<tr>
<th>Case</th>
<th>Spectra</th>
<th>ZWARN=0</th>
<th>( 1\sigma(\Delta v) ) [km/s]</th>
<th>Cat. Failure</th>
<th>Cum. Efficiency (ZWARN=0)</th>
<th>Cum. Efficiency (ZWARN=0 and No. Cat. failures)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Nominal</td>
<td>5000</td>
<td>4338</td>
<td>15</td>
<td>20 (0.5%)</td>
<td>86.8%</td>
<td>86.4%</td>
</tr>
<tr>
<td>Airmass</td>
<td>5000</td>
<td>4509</td>
<td>14</td>
<td>21 (0.5%)</td>
<td>90.2%</td>
<td>89.8%</td>
</tr>
<tr>
<td>Seeing</td>
<td>5000</td>
<td>4426</td>
<td>14</td>
<td>34 (0.8%)</td>
<td>88.5%</td>
<td>87.8%</td>
</tr>
<tr>
<td>Transparency</td>
<td>5000</td>
<td>4418</td>
<td>14</td>
<td>23 (0.5%)</td>
<td>88.4%</td>
<td>87.9%</td>
</tr>
<tr>
<td>E(B-V)</td>
<td>5000</td>
<td>4485</td>
<td>14</td>
<td>15 (0.3%)</td>
<td>89.7%</td>
<td>89.4%</td>
</tr>
</tbody>
</table>
Figure 9.2. Performance of Redshift measurement for the nominal case of observing conditions. Top panel: Redshift errors shown with respect to true $z$ (a), $r$-band magnitude (b), and [OII] flux (c). Shaded regions indicate 68% and 95% confidence regions for each slice. Bottom panel: RMS for the distributions with respect to $z$ (d), $r$ magnitude (e) and [OII] flux (f). Cumulative efficiency is 86.4% with 0.5% catastrophic failure rate.
9.3. Observational effects and efficiency

To measure the systematic effects of observing conditions on the BAO measurement, I apply the redshift efficiency effects on a mock galaxy distribution and measure the 2PCF before and after applying those effects. The derivation of the efficiency model and its use on the mock catalog is described below.

9.3.1. Parametrization of redshift efficiency

The parametrization scheme for the final efficiency model encompasses two steps. First, redshift efficiency is analyzed for the nominal case where all observational parameters are fixed, but the $\text{[OII]}$ flux for the ELGs templates vary. Note that the same $\text{[OII]}$ flux distribution prior is set for all five cases of the spectroscopic simulations by asserting the same random number generator while selecting galaxy templates. Second, efficiency as a function of varying observing conditions is established relative to nominal case. The product of two give the final efficiency model.

9.3.1.1. Dependence on $\text{[OII]}$ flux

For each target, a probability is assigned for obtaining a good redshift. This probability comes from the cumulative probability obtained for the $\text{[OII]}$ flux level compared to a threshold optimized to DESI sensitivity. If $\text{[OII]}_{\text{true}}$ is the true $\text{[OII]}$ flux for a target at a certain redshift, and $\text{[OII]}_{\text{thresh}}$ is the minimum threshold to get a correct redshift at that redshift, then the cumulative density function (CDF) of it being correct can be obtained using error function. For a standard normal distribution, this is given by,

\[
P_{\text{[OII]}} = \frac{1}{2}[1 + erf\left(\frac{x - 1}{\sqrt{2}}\right)]
\]

\[
x = \frac{\text{[OII]}_{\text{true}}}{\text{[OII]}_{\text{thresh}}}
\]

(9.3)

(9.4)
The $[\text{OII}]_{\text{thresh}}$ obtained for the redshift range 0.6–1.6 is shown in Fig. 9.3. The efficiency model ultimately should give a boolean of whether a target gets a correct redshift or not. Thus if a random probability is higher than the CDF probability obtained above, the redshift will fail and vice versa. Before assigning the efficiency as a boolean for a good redshift or not for the individual targets based on the [OII] flux, I present the model affecting this efficiency due to observing conditions. The final outcome from the efficiency model is thus the convolution of observational effect and the CDF probability obtained here for the [OII] flux values.

### 9.3.1.2. Dependence on observational parameters

Let $N$ be the total number of targets in a given bin of observing parameter $y$. Let $N_1$ be the number of correct redshift ($ZWARN = 0$) out of $N$ when $y$ is varied and $N_2$ be the number of successful redshifts for the nominal case. Then, the fractional change in efficiency for that bin is given by

$$\Delta \text{Eff}_{\text{bin}} = \frac{N_1 - N_2}{N_2}$$

(9.5)

This fractional change in redshift efficiency when $y$ is varied versus when it is kept nominal, is modeled as a polynomial expansion about the nominal value,

$$v = y - y_{\text{nominal}}$$

$$\Delta \text{Eff} = c_0 v + c_1 (v^2 - <v^2>)$$

(9.6)

where $c$’s are the expansion coefficients which can be determined by fitting Eq. 9.6 to the observed binned array of $\Delta \text{Eff}$. An example model for varying seeing case is shown in Fig. 9.4. On the left is shown the $\Delta \text{Eff}$ with respect to seeing values. The red solid lines represent redshift with $ZWARN = 0$, while the blue line excludes the catastrophic failures. The respective dashed lines are the fit models.
Figure 9.3. [OII] flux threshold reflecting DESI sensitivity for a correct redshift as a function of $z$. Flux below the threshold are more likely to yield an incorrect redshift.
Figure 9.4. Fractional change in redshift efficiency due to a variation of seeing compared to nominal seeing of 1.1". The solid lines represent the successful redshifts while the dashed lines are the respective derived models.
The efficiency model is obtained as

\[
E_y = <\text{Eff}_{\text{nominal}}>(1 + \Delta\text{Eff}_y)
\]  

(9.7)

where \(<\text{Eff}_{\text{nominal}}>\) is the expected nominal efficiency. The cumulative efficiency when all observing conditions are considered, is then given by,

\[
P_{\text{obs}} = \prod_y E_y + \epsilon
\]  

(9.8)

where \(y = \{\text{airmass, transparency, seeing}\}\) for this study. The \(\epsilon\) is a noise term calculated using the random normal distribution of the effective RMS obtained after adding in quadrature the RMS values of \(\Delta\text{Eff}\) residuals in each case. The effective RMS is obtained to be 7.2%. Model extrapolation towards good observing conditions can yield efficiency values higher than 1. In that case, the values are clipped to maximum of 1.

The poor observational conditions will degrade the CDF probability obtained above from [OII] flux distribution for the nominal case. The final likelihood of obtaining a correct redshift for a target is then given by multiplying the two probabilities, i.e,

\[
P_{\text{good}} = P_{\text{[OII]}} \times P_{\text{obs}}
\]  

(9.9)

Considering a random uniform probability distribution, a target is assigned a ZWARN flag of 0 if \(P_{\text{good}}\) for the target is larger than the random probability. Otherwise the ZWARN flag is set to 4 implying an unsuccessful redshift measurement. Since the efficiency model only provides the information whether a given target has a correct redshift or not, this is not sufficient to perform the clustering analysis for which a redshift measurement is necessary for each target of the given catalog. In the next section, I describe the galaxy catalog used in this study and how the redshifts of the targets are modified to account for observational effects.
9.3.2. Mock Galaxy catalog

I use a mock ELG Galaxy catalog, used by the DESI data systems for several data challenges in 2016 and 2017. The mock is build using Gaussian random field using CoLoRe\(^1\). It covers the DESI footprint but without any depth variations. The targets have unique (RA, DEC, \(z\)) set. The cosmology is modeled using power spectrum generated by CosmoEmu v2.0\(^2\) ([72]) with input parameters from Planck 2015 results from Ref. [125]. A HEALPix\(^3\) map for the ELG mock covering the DESI footprint is shown in Fig. 9.5.

Because the goal of this study is to observe the effect of redshift efficiencies based on spectroscopic observing conditions alone on the BAO signal, I develop an analysis pipeline focusing only on the spectroscopic effects. From the obtained distribution of the observing parameters from survey simulation, I assign each target of the mock with a random value of observing parameters. Furthermore, efficiency model derived earlier also need [OII] flux for each target. I use the distribution obtained with the survey data challenge simulations performed by the DESI data team in 2016, which is shown in Fig. 9.6. To assign [OII] flux to each target in the mock, I first divide the data challenge [OII] flux distribution in the redshift bins of size 0.01 to create 1D histograms of [OII] for each redshift bin. Next, from the probability distribution of the obtained [OII] 1D histogram for that bin, a scikit Gaussian kernel density sampling of the [OII] flux is performed for the targets of the mock in that redshift bin.

Now that the mock has both the [OII] flux and observing condition for each target, passing the redshift efficiency model obtained above will assign a ZWARN flag to each target. Since the efficiency model does not affect the redshift, an error model is used to normally fluctuate the true redshifts of the targets to yield the effective measured redshifts. Errors are calculated using the random normal fluctuations about the observed RMS of [OII]

\(^1\)github.com/damonge/CoLoRe

\(^2\)http://www.hep.anl.gov/cosmology/CosmicEmu/emu.html

\(^3\)http://healpix.sourceforge.net/documentation.php
Figure 9.5. A HEALPix map of ELG Gaussian random mock covering the full DESI footprint.
Figure 9.6. Distribution of $[\text{OII}]$ flux as a function of redshift. This distribution is obtained from target selection on the mock for the DESI survey data challenge in 2016.
dependent redshift efficiency in the nominal spectroscopic simulations, shown in the bottom right plot of Fig. 9.2. Finally, for all the $ZWARN = 0$ cases, a catastrophic failure rate is applied based on the fraction of such failures observed in the nominal simulation. The catastrophically failed fraction is randomly selected and redshifts are assigned in the range 0.6 to 1.6 randomly, uniformly. With these simulations, it is not clear if the catastrophic failed fraction shows a conclusive dependence on redshift. So I take this as a conservative approach.

The final outcome from this is a new catalog with measured redshift, an error on the measured redshift and a warning flag for each input target. From here, I refer to the original mock as the true catalog and the new catalog after updating the redshift efficiency as the modified catalog. Fig. 9.7 shows the redshift distribution on a subsample of the modified catalog. All the targets shown have $ZWARN$ of 0, and those not lining up with the unit slope straight line are the catastrophic failures.

9.4. ELG Clustering analysis

I perform the clustering analysis on the true catalog and the modified catalog separately in the redshift range 0.6–0.7. The objective here is to observe the effects on the 2PCF for the modified catalog compared to that for true catalog.

9.4.1. Measuring 2 point correlation function

Measurement of the 2PCF using the Landay-Szalay estimator given in Eq. 3.16 requires computation of auto-correlation and cross-correlation between the data and the random catalogs. For this, a random catalog has to be generated that would not have any cosmological signal but have same selection effects as the data catalog. Since I am using the mock that is devoid of multi layer systematics from fiber assignment and target selection, random creation for the true mock can be obtained by shuffling the targets conserving the number density with respect to $z$, or by randomly generating a sample using the probability distribution of the data catalog. A demonstration of the downsampled (for viewing purpose) true catalog
Figure 9.7. A plot showing the modified redshift versus true input redshift. Only first 50000 objects are shown for better viewing. Most of the targets line up on the unit slope line, thus suggesting the correct redshifts; while others scattering away from this line represent the catastrophic failures.
and a random catalog in a cropped RA versus DEC map is shown in Fig. 9.8.

Just as the modified catalog was built from the true catalog using the redshift efficiency model, a modified random catalog is build from the true random catalog using the same procedure. This way of generating randoms is a forward modeling approach where both the data catalog and randoms accomodate the same effects from spectroscopic observations on a forward process. On the backward modeling, randoms are constructed after the fact. One would construct the modified random from the modified catalog, and somehow disentangle the spectroscopic effects to create an effective random catalog that would correspond to the true catalog.

The 2PCF on the true catalog is first calculated using the publicly available code treecorr\(^4\) release version 3.3 ([81]). The algorithm uses \(k - d\) trees to make the counting problem highly efficient, and the code is highly parallelized to run on multiple cores. Given the target positions in RA, DEC and distance (derived from redshift using fiducial cosmology), the calculation involves the auto correlation of distances in a predefined binning. For two catalogs (e.g., data and random), a cross correlation is computed. Finally, the 2PCF is obtained using Landay and Szalay estimator for the true catalog and the corresponding random catalog. The same procedure is repeated for the modified mock and respective modified random catalog. Fig. 9.9 shows the computed 2PCF \(\xi(r)\) for the mock and the modified catalog. A BAO bump is evident about the BAO scale of 100 Mpc/h in both cases. The modified 2PCF is noisier for obvious reason due to poorer statistics. Suppression about 20 Mpc is seen on the modified 2PCF compared to true case. While this lower scale suppression can potentially be addressed by marginalizing over the broadband shape appropriately (e.g., [135]), my concern for this analysis is the degradation of the BAO feature precision and potential shift due to redshift measurements. Therefore, in this analysis, I attempt to quantify the shift on the BAO position on the modified case relative to the true mock using the following fitting prescriptions.

\(^4\)https : //github.com/rmjarvis/TreeCorr
9.4.2. BAO Fitting model

To see any systematic effect on the BAO position due to redshift performance and quantify the subsequent uncertainty, I perform fitting of the 2PCFs measured from the true catalog and the modified catalog analyses.

In practice, analyses of the 2PCF from survey data involves creation of many mocks and similarly large set of randoms to model the full covariance (e.g., [111]). Because there is only one mock in this case, I will use the power of Gaussian process (GP) regression to model covariance from the measured 2PCF in a fully data-driven fashion. I have used GP regression in multiple context earlier in this thesis. The advantage with this fitting is that both the maximum likelihood estimate (MLE) and maximum a posterior (MAP) estimate of the regression are naturally obtained in a Bayesian framework. See Appendix B for the mathematical framework of GP.
Figure 9.9. 2PCF obtained for the mock in the 0.6-0.7 redshift range (blue) and for the ZWARN=0 subsample after applying the efficiency model (red). BAO bump is evident on both cases, but the 2PCF is noisier for the modified case.
9.4.2.1. Gaussian process regression model

As before, I use scikit sklearn package to perform the GP fits of the true and the modified 2PCFs. The GP fits for both the true and modified 2PCFs are shown in Fig. 9.10. The solid black line is the MLE model and the shaded region represent the 95% confidence MAP estimate. I add the uncertainty from the MAP estimate in quadrature to the measured uncertainty to achieve the final uncertainty for the measured 2PCF. An alternative approach would be to perform a MCMC analysis using an empirical/anlytic model (see below) and derive the parameter and intrinsic variance and covariances using the marginalized posterior distributions. I performed similar analysis for the measurement of Hubble constant using supernovae in Chapter 6.

To measure the shift on the BAO feature on the GP fit of the modified 2PCF relative to an assumed fiducial model, I use the parametrization following Ref. [4] as

\[ \xi(r) = B^2 \xi^{fid}(\alpha r) + \frac{a_2}{r^2} + \frac{a_1}{r} + a_0 \]  

where \( \xi^{fid}(\alpha r) \) is the fiducial 2PCF model, which I take the GP MLE fit of the true 2PCF obtained above. For the fiducial model, \( \alpha = 1 \) by definition. The multiplicative factor \( B^2 \) adjusts the amplitude of the model to fit the data and the additive polynomial parameters marginalize over the broad-band shape. Any shift in the BAO scale in the modified case is obtained from the best fit value of \( \alpha \). After performing the fit for the modified 2PCF, the best fit model yield a \( \chi^2/dof = 1.11 \) and the best fit value of \( \alpha \) is obtained to be \( 0.998 \pm 0.008 \). The best fit model is shown in Fig. 9.11. This value of \( \alpha \) suggests no significant bias introduced from redshift failures and the adopted efficiency model on the BAO scale.

9.4.2.2. Empirical model

Following Ref. [168] (see Eq. 9 on their paper), the observed correlation can be modeled with the Gaussian for the BAO feature over a smooth power law decay.
Figure 9.10. Left: GP regression fit of the 2PCF calculated for the true catalog. Right: Same as left plot but for the modified 2PCF. The black solid line is the MLE best fit model and the shaded region indicate the 95% MAP estimate.

Figure 9.11. Fitting of modified 2PCF (red points) to Eq. 9.10 using fiducial model as GP fit of true 2PCF (blue curve). The green line is the best fit model, for which the $\chi^2$ and best fit $\alpha$ parameter are shown.
\[ \xi(r) = B + \left( \frac{r}{r_0} \right)^{-\gamma} + \frac{N}{\sqrt{2\pi\sigma^2}} \exp\left( -\frac{(r - r_m)^2}{2\sigma^2} \right) \] (9.11)

where first term \( B \) sets up the zero-point, the second term is the power law term with index \( \gamma \) and the third term is the Gaussian function with parameters \( N, r_m \) and \( \sigma \), that is used to model the observed BAO feature. Here also, the motivation is to observe any significant shift of the BAO position from the fit on the modified 2PCF relative to the true case. This can be quantified from the best fit mean of the Gaussian \( (r_m) \) given the broadband shape of the 2PCF is marginalized by the first two terms of Eq. 9.11. Fig. 9.12 shows the fits on the 2PCFs for the true case (left) and the modified case (right).

9.4.3. Result: Observational systematics on BAO scale

From the above fitting procedures of the measured 2PCFs for the true and the modified 2PCFs, the best fit measurement of the BAO scale in both cases are found to be consistent. From the GP fit, I found a systematic uncertainty on the BAO scale about 1%. While the precision of the measured scale has worsened by a factor of \( \sim 3 \) for the modified 2PCF as seen on the empirical model fit, no statistically significant shift of BAO scale is observed. Uncertainty obtained for \( r_m \) (\( \Delta r_m \)) for the modified 2PCF constitutes both the enlarged statistical (due to reduction in the number of galaxies compared to true catalog) and potential systematic uncertainty from the observational effects. The final uncertainty from the empirical fit on the BAO shift for the given sample is calculated using Poisson approximation. Thus the BAO scale shift from the fitted \( r_m \) values for the true and the modified catalog is \( 2.1 \pm 2.7 \, \text{h}^{-1}\text{Mpc} \).

Therefore, with the available Gaussian random mock and spectroscopic efficiency model established from the observed redshift rates in the samples of spectroscopic simulations, I find no statistically significant bias in the BAO scale measurement due to redshift efficiency, where this statistical limitation is still substantial due to low Monte Carlo statistics. A systematic uncertainty at the level of 1% is estimated on the \( \alpha \) parameter when the true input mock is taken as the fiducial model. With this finding, I conclude this chapter and
Figure 9.12. Left: Fitting of the true 2PCF using the empirical model of Eq. 9.11. The respective best fit mean of the Gaussian peak, representing the BAO position, is also shown. Right: Same as left plot but for the modified case.

move towards the final conclusion in the next chapter.
Chapter 10

CONCLUSION

From the analyses presented in this thesis, I conclude this thesis with the following outcomes. ROTSE III is very useful for transient searches. The wide field areal coverage, augmented by slew speed on the order of a few seconds makes ROTSE a very useful facility to probe transients that require rapid response. Also, the fully robotic nature and full control over the observing time allows much longer baseline to design a strategic SN follow up program. I made use of this by adding targets of interest in the scheduler such that better sampled light curves could be obtained and astrophysics could be better constrained.

The use of new image differencing software that I wrote was one of the turning point to pursue a cosmological study using a sample of ROTSE SNe data. This software has shown much better performance particularly when the host background couldn’t be accurately subtracted with the old software. The new image differencing technique yielded up to 20% more detection points, and the photometric residual scatter reduced up to a factor of 3, yielding statistically reasonable pull distributions. With the flexibility of kernel models, this technique allows much larger degree of freedom compared to pre-defined PSF models used in the previous cross-convolution technique. It would be interesting to see the performance of this tool on the online system.

A bolometric correction model was established for IIP SNe. \(UBVRI\) pseudo-bolometric calibration was obtained at 12% precision with ROTSE data alone. The scatter reduced down to 5% when a \(B-V\) correction was possible. Full \(UBVRIJHK\) bolometric calibration model was established using literature data of 4 well studied SNe. There, a \(B-V\) dependent correction yielded 2% level scatter. With this, a composite correction model was also established for ROTSE data.
The EPM analysis on ROTSE SNe IIP sample yield a very encouraging measurement of Hubble constant in the low-z universe. Events as few as five give $H_0 = 73.5 \pm 5.6$ km/s/Mpc with EPM method’s intrinsic systematics below 0.1 mag. I am excited that this may be indicative that SN IIP cosmology will be a serious business in the larger SN surveys. There is room for better modeling of dilution parameters and extrapolation models of the astrophysical observables. This will also improve the calibration by reducing the systematics even more.

The upcoming DESI experiment is commissioning in 2019. Preparations for the first light are in rapid progress. The Quicklook data reduction system is nearing to the production level processing stage. Interfacing with the online operation and database framework is also developing rapidly. By analyzing the spectroscopic performance of the pipeline on a set of simulated spectra, an efficiency model was established. By sampling the observing conditions on the targets of a mock galaxy catalog and varying the redshifts based on the efficiency model, I created a modified mock catalog. Performing the clustering statistics with the two point correlation functions on the true and the modified mock catalogs, no statistically significant bias was observed on the BAO scale of the modified catalog relative to the true catalog. This was also true for the empirical model fit of the correlation function that I demonstrated. The systematic uncertainty however is still large compared to the design sensitivity at the same redshift. While one can study a mitigation strategy for the observational effects, more simulations would be needed to accurately model the covariance. This will require the generation of large number of realistic mock realizations through cosmological simulations and that is exactly where DESI is proceeding forward.

Finally, this study can be a model for approaching to quantify the systematics from observations on all galaxy types and grey and bright programs, for which there will be additional constraints from moon brightness. However in the final cosmological simulations, for a forward modeling approach, the systematics from the fiber assignment and target selection have to be correctly understood and isolated before any impact from the spectroscopy alone can be modeled.
Appendix A

Expectation maximization

Expectation maximization (EM) is a widely used technique in the statistical and machine learning analyses of various forms of data. It exhibits an iterative procedure to calculate the maximum likelihood estimate (MLE), particularly useful when the data is missing or hidden. The MLE involves finding the optimum estimate of the model parameters given the data. I have used packages following this approach significantly in two cases earlier in the image differencing algorithm in Section 6.1.1.2 and photospheric velocity estimation of SN ejecta using Gaussian mixture model in Section 6.2.1.1. Below I describe the EM algorithm and present the application in the context mentioned above.

A.1. Algorithm

Each iteration of EM involves two steps: the expectation E-step and the maximization M-step. Let $X$ be the observed data, $Z$ be the latent or missing data, and $\theta$ define the model parameters. Then, following https://en.wikipedia.org/wiki/Expectation–maximization_algorithm, the MLE for $\theta$ can be obtained using

$$L(\theta; X) = p(X|\theta) = \int p(X, Z|\theta) dZ$$  \hspace{1cm} (A.1)

Thus estimate of theta involves marginalization of the latent variables, but these are also unknown. This is where the E-step comes in.

A.1.1. E-step

E-step basically computes the expectation of the likelihood function, given the data and the estimate of parameters $\theta^{(i)}$ corresponding to values for the $i^{th}$ iteration. The first iteration
involves assigning a set of random values for $\theta$, while the subsequent iteration uses the values obtained from earlier M-steps. Thus given values of $\theta$, E-step measures the expectation of $Z$.

$$Q(\theta|\theta^{(i)}) = E_{Z|X,\theta^{(i)}}[p(X, Z|\theta)]$$  \hspace{1cm} (A.2)

A.1.2. M-step

The expectation value obtained for $Z$ from the E-step can now be used to better constrain, i.e., maximize $\theta$. These values for $\theta$ will serve the E-step for the next iteration and the whole procedure is repeated until converged.

$$\theta^{(i+1)} = \arg \max_{\theta} Q(\theta|\theta^{(i)})$$  \hspace{1cm} (A.3)

A.2. Expectation Maximization for image differencing

This is the use case I adopted from empirca as described in Section 6.1.1.2. Here I show the estimate of parameters to construct the model image such that it can be subtracted from science image can be subtracted pixel by pixel to yield the difference image. Given a set of kernel convolved template image, a PCA eigenbasis is constructed.

An observation $\vec{y}$ can be expanded in the PCA eigen basis $\{\vec{\phi}\}$ as

$$\vec{y} = \sum_{k} c_k \vec{\phi}_k$$  \hspace{1cm} (A.4)

By orthogonality of the eigen vectors,

$$\vec{y} \cdot \vec{\phi}_m = \sum_{k} c_k \vec{\phi}_k \cdot \vec{\phi}_m = \sum_{k} c_k \delta_{km} = c_m$$  \hspace{1cm} (A.5)

The aim is to solve for the PCA eigenvectors, while the coefficients are evaluated in a maximum likelihood sense when fitting to data. The PCA eigenvectors are the model parameters
to be solved and the coefficients serve as the latent variables. The EM algorithm for the
most significant eigenvector is presented in Ref. [7] with the following pseudocode.

\[ \vec{\phi} \leftarrow a \text{ random vector of same dimension as an observation} \]

repeat until converged :

For each observation \( \vec{x}_i \) :

\[ c_i \leftarrow \vec{x}_i.\vec{\phi} \quad \text{“E - step”} \]
\[ \vec{\phi} \leftarrow \sum_i c_i \vec{x}_i / \sum_i c_i^2 \quad \text{“M - step”} \]
\[ \vec{\phi} \leftarrow \vec{\phi} / |\vec{\phi}| \quad \text{Renormalize} \]

return \( \vec{\phi} \)

For the image differencing, observation set \( X \) constitutes the set of kernel convolved
templates and 1 science image. The linear model for the \( j^{th} \) column (observation) becomes

\[ X_j = PC_j + \text{noise} \quad (A.6) \]

where \( P \) is the matrix of principal components and \( C_j \) is the 1D coefficient matrix for the \( j^{th} \) component. Following [7], the coefficient vector \( \vec{c} = C_j \) for the \( j^{th} \) observation \( \vec{x} = \vec{X}_j \) can be solved using linear least square solution assuming the pixel covariance \( V_j \) to be Gaussian. Mathematically this can be written as,

\[ \vec{x} = P\vec{c} \quad (A.7) \]
\[ V^{-1}\vec{x} = V^{-1}P\vec{c} \quad (A.8) \]
\[ P^T V^{-1}\vec{x} = P^T V^{-1}P\vec{c} \quad (A.9) \]
\[ (P^T V^{-1}P)^{-1} P^T V^{-1}\vec{x} = \vec{c} \quad (A.10) \]

The final solution obtained above for \( \vec{c} \) is the same MLE solution given in Eq. 6.11.

Once an eigenvector is determined, the next eigenvector is solved using the same steps after
subtracting the projection of the earlier eigenvector from the data. i.e, for the \((j + 1)^{th}\)
eigenvector to be solved,

\[ \mathbf{X} \leftarrow \mathbf{X} - \mathbf{P}_j \otimes \mathbf{C}_j \quad (A.11) \]

where \( \otimes \) represents the outer product. Note that if there were previous eigenvectors, they are assumed to be already subtracted in the same way. This is an orthogonalization technique (e.g., Gram-Schmidt Orthogonalization\(^1\)) such that PCA eigenbasis remains orthogonal. The subtraction of the projection along \( \mathbf{P}_j \) removes the data variation in that direction. The next significant PCA eigenvector will be orthogonal to this along which the remaining data variance is minimum. This is the merit of PCA that potentially correlated data sets are decomposed into linearly uncorrelated principal components. See Ref. \([7]\) for the full solution of the PCA eigenvectors using EM method.

\(^1\)http://mathworld.wolfram.com/Gram-SchmidtOrthonormalization.html
Appendix B

Gaussian process regression

Gaussian process (GP) regression is a powerful regression technique to obtain a non-parametric or more correctly said, an infinite-parametric regression modeling of the data considered. This statistical technique is very powerful to yield full posterior distribution of the regression model and useful when no analytic solution is known.

I have used GP regression from scikit in multiple instances in this thesis, such as, modeling the K-correction in Section 6.3.3, light curve fitting for the performance analysis of image differencing in Section 6.1.3, determining plateau length of the lightcurve in Section 5.1.4.6.2 and fitting the BAO scale in Section 9.4.2. Below I describe the basic mathematical formalism of GP.

B.1. Algorithm

Let \( f \) define a set of random functions that fit a set of data \( f \) at some values \( x \), i.e, the data points are observed in a set of points \( x \) and we want to perform a GP regression as a continuous function that we wish to sample in our test positions \( x_\ast \). Assuming a normal distribution of the fluctuations of the data, the most likely function for the regression model can be obtained using the joint probability distribution (e.g., [107]),

\[
\begin{pmatrix}
  f \\
  f_\ast
\end{pmatrix}
\sim \mathcal{N}
\begin{pmatrix}
  \mu \\
  \mu_\ast
\end{pmatrix},
\begin{pmatrix}
  K & K_\ast \\
  K_\ast^T & K_{\ast\ast}
\end{pmatrix}
\]  

(B.1)
Using standard procedure of conditioning normal distributions ([107]), one obtains the posterior probability for \( f^* \), given as

\[
p(f^*|f, x, x_*) \sim \mathcal{N}(f^*|\mu^*, \Sigma^*)
\]  

(B.2)

where \( \mu^* \) is the mean function, which is the most likely solution, and \( \Sigma^* \) is the covariance of the model. Thus the GP regression gives the full MAP estimate for the regression problem. In the above procedure, one question remains about how the covariance is defined. Covariance can be obtained using some kernel function that defines how the points fluctuate. In all my analysis in this thesis, I have used the squared exponential kernel, given by

\[
k(x, x') = \sigma_v^2 \exp\left(-\frac{1}{2l^2} (x - x')^2\right)
\]  

(B.3)

The hyper-parameters \( \sigma_v \) adjusts the vertical scaling while \( l \) adjusts the horizontal length scale. With this kernel, the points that are close will covary more than those that are further away.
BIBLIOGRAPHY


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