Chapter 2

Facilities, observations and data reduction

2.1 Introduction

Early and deep observations of supernovae are crucial to understand the physics of these events. Because of their transient nature and relatively fast decay of flux after discovery, quick follow-up observations of such objects are required. The supernovae are randomly discovered by different groups of amateur astronomers and being reported as IAU and/or CBET circulars. The events are now also discovered and monitored in a much more systematic way through different survey programs like ‘Catalina Real-Time Transient Surveys’ (CRTS), ‘Lick Observatory Supernova Search’ (LOSS), ‘Robotic Optical Transient Search Experiment’ (ROTSE), ‘Palomar Transient Factory’ (PTF), ‘MASTER robotic Net’ and many other survey programs. Our primary-goal is to perform a systematic and extensive follow-up of different proximate and bright events reported in IAU and CBET circulars. The multiwave-
length campaign of these events along with optical spectroscopic monitoring are essential to probe the dynamics of the event and the burst mechanism, to estimate the energetics and to understand the nature of the progenitors and its environment.

For a dense and extended coverage of a transient at optical wavebands, getting an access of an optical telescope with moderate light gathering power along with a high efficiency detector is required. For the entire thesis, the optical data were primarily acquired from three Indian telescopes namely 1.04-m Sampurnanand Telescope (ST) at Aryabhatta Research Institute of Observational Sciences (ARIES), 2.0-m IUCAA Girawali Observatory telescope under Inter University Center for Astronomy and Astrophysics (IUCAA) near Pune and 2.01-m Himalayan Chandra Telescope (HCT) at Indian Astronomical Observatory (IAO), Hanle. Additionally, the optical observations were also conducted from 3m to 6m optical telescopes outside India, through different active scientific collaborations between ARIES and other observatories over the globe. The optical telescopes, outside India, that have been used for this work are: (i) 0.61-m Lowell telescope, Perth Observatory, Australia; (ii) 0.45-m ROTSE III telescope, McDonald Observatory, USA; (iii) 0.6-m REM telescope, La Silla, Chile; (iv) 3.5-m Telescopio Nazionale Galileo (TNG), Canary Islands, Spain; (v) 3.6-m New Technology Telescope (NTT), ESO, Chile; (vi) 6.0-m Big Telescope Alt-azimuthal (BTA), Special Astrophysical Observatory, Russia and (vii) 9.2-m Hobby Eberly Telescope (HET), McDonald Observatory, USA.

To understand the multi-wavelength characteristics of these objects Near-ultraviolet (NUV) X-ray and Radio data have also been acquired, reduced and analysed. For NUV and X-ray data, archives of space-based telescopes like *Swift*/UVOT and XRT were used. For radio analysis, archival data of radio telescope ‘Very Large Array (VLA)’ has been used. This facility is situated at the Mexican Desert and op-
2.2 Data acquisition at optical and near infrared wavebands

The characteristics of optical telescopes and the back-end instruments used for collecting the photometric and spectroscopic data discussed here.

2.2.1 The optical telescopes

India lies in the middle of the 180° wide longitude belt between Canary Islands (∼ 20° W) and Eastern Australia (∼ 160° E), thus filling in an important temporal...
Table 2.1 Parameters of the used Indian optical telescopes.

<table>
<thead>
<tr>
<th>Telescope</th>
<th>ST</th>
<th>HCT</th>
<th>IGO</th>
</tr>
</thead>
<tbody>
<tr>
<td>Site</td>
<td>Nainital</td>
<td>Hanle</td>
<td>Pune</td>
</tr>
<tr>
<td>Longitude</td>
<td>79° 27’ E</td>
<td>78° 57’ E</td>
<td>73° 40’ E</td>
</tr>
<tr>
<td>Latitude</td>
<td>29° 21’ N</td>
<td>32° 46’ N</td>
<td>19° 05’ N</td>
</tr>
<tr>
<td>Altitude</td>
<td>1951 m</td>
<td>4500 m</td>
<td>1000 m</td>
</tr>
<tr>
<td>System</td>
<td>Ritchey-Chrétien</td>
<td>Ritchey-Chrétien</td>
<td>Ritchey-Chrétien</td>
</tr>
<tr>
<td>Focal length</td>
<td>Cassegrain: f/13</td>
<td>Cassegrain: f/9</td>
<td>Cassegrain: f/10</td>
</tr>
<tr>
<td>Primary Mirror</td>
<td>1.04-m</td>
<td>2.01-m</td>
<td>2.0-m</td>
</tr>
<tr>
<td>Diameter</td>
<td>0.37”/pix</td>
<td>0.17”/pix</td>
<td>0.31”/pix</td>
</tr>
</tbody>
</table>

In optical astronomy filters are used mainly two reasons: (i) either to reduce the overall intensity of light or (ii) to restrict the wavelength range of the incoming light. Nowadays detectors have a wider wavelength response and an improved sensitivity in comparison to previous detectors. Hence the filters are important for the flux measurement of various astronomical objects at different wavelengths. The broad band filters used for this study are Johnson UBV and Cousins RI with ST, Nainital and Bessell UBVRI filter system with HCT, Hanle. Details about the filter system

1[http://aries.res.in](http://aries.res.in)
2[http://www.iiap.res.in/centres/iao.htm](http://www.iiap.res.in/centres/iao.htm)
3[http://iucaa.ernet.in](http://iucaa.ernet.in)
Table 2.2 Details of the filter system. The thickness of the glass in mm is indicated in brackets

<table>
<thead>
<tr>
<th>Telescope</th>
<th>Filters</th>
<th>$\lambda_{\text{eff}}$ (Å)</th>
<th>Band Width (Å)</th>
<th>Glass combination</th>
</tr>
</thead>
<tbody>
<tr>
<td>ST</td>
<td>$U$</td>
<td>3650</td>
<td>680</td>
<td>UG 1(2) + CuSO4 Solution(5)</td>
</tr>
<tr>
<td></td>
<td>$B$</td>
<td>4400</td>
<td>980</td>
<td>GG 385(2) + BG 18(1) + BG 12(1) + KG 3(2)</td>
</tr>
<tr>
<td></td>
<td>$V$</td>
<td>5500</td>
<td>890</td>
<td>GG 495(2) + BG 12(2) + KG 3(2)</td>
</tr>
<tr>
<td></td>
<td>$R$</td>
<td>6500</td>
<td>2200</td>
<td>OG 570(2) + KG 3(3)</td>
</tr>
<tr>
<td></td>
<td>$I$</td>
<td>8000</td>
<td>2400</td>
<td>RG 9(3) + WG 305(2)</td>
</tr>
<tr>
<td></td>
<td>Hα*</td>
<td>6565</td>
<td>80</td>
<td>—</td>
</tr>
</tbody>
</table>

| HCT       | $U$     | 3500            | 500           | UG 1(1) + S8 612(2) + WG 295(2) |
|           | $B$     | 4400            | 900           | BG 37(3) + BG 39(1) + GG 395(2) |
|           | $V$     | 5500            | 800           | BG 40(3) + GG 495(2) |
|           | $R$     | 6500            | 1500          | OG 570(3) + KG 3(2) |
|           | $I$     | 7900            | 1100          | RG 9(2) + WG 305(3) |

*This is a Narrow Band filter. For continuum subtraction R-band image has been taken.

are given in Table 2.2. The filters used are basically combination of coloured glasses of WG, GG (Sulpher and Cadmium sulfide) and OG (Cadmium selenide) which absorb light of quite sharply defined wavelengths. The UG (violet) and BG (blue) are made of ionic glasses.

### 2.2.3 The detectors for imaging

Detectors are the indispensable part of modern telescopes and research facilities. We have used **Charge Coupled Device** (CCD) as a detector for our observations. This device, which replaced the photographic film as well as photomultiplier tubes – is an ultra-sensitive, silicon semiconductor, ‘chip’-based optical detector is composed of an array of thousands or millions of individual, tiny picture elements that almost
Figure 2.2 Response curve of 2k×4k CCD (squares) at \(-100\,^\circ\text{C}\), 2k×2k CCD at \(-120\,^\circ\text{C}\) (triangles) and 1k×1k CCD at \(-100\,^\circ\text{C}\) (circles). The curve indicates that all the three CCDs peak near \(\lambda \sim 6700\,\text{Å}\).

touch each other and are called ‘pixels. This was able to produce a two dimensional detection as an integrating mode, good linear response, dynamic range (\(\sim 10^5\)) and quantum efficiency (\(\gtrsim 80\%\) at \(\lambda \sim 6900\,\text{Å}\)) which permit to image much fainter objects in relatively lesser time with a large dynamic range of recording over a vastly different levels of brightness. CCDs are nowadays used as detectors in most of the X-ray, NUV, optical and NIR imaging astronomical instruments. CCD operates on the principle of photoelectric effect.

The CCD camera used for astronomical purpose consists of a two dimensional array of photon detectors in a layer of semi-conducting material silicon. Its operation is based on photoelectric effect which can be described as follows. When the incident photons having energy 1.1–4 eV, hit the silicon chip within a pixel, it
generate single electron-hole pairs inside the chip, whereas those of higher energy produce multiple pairs. Each individual pixel is capable of collecting the photons and storing its energy in the form of produced electrons which can be read out from the CCD array to a computer in the form of a digital image of varying intensities of light detected by the CCD. The potential well of each pixel is capable to collect a definite number of electrons and it behaves like a bucket when it is exposed to collect the light. For a faint object, the shutter of the camera can be opened for a longer timespan, to generate and collect more number of $e^-$ (integrated) inside each pixel of the CCD and that can be readout after completion of exposure. In effect, the faint object becomes more prominent in the processed digital image. Depending on the capacity of CCD pixels (known as ‘Full-well capacity’), one can increase the time duration (known as ‘integration time’) to collect more photons.

The information from the rows of pixels move down to a single parallel row (the serial register) which is read out sequentially by Analog-to-Digital (A/D) converter where it is measured and then recorded. The measuring device is emptied and once again the process is repeated. This process continues until all of the pixels have been measured (read out). The time required to read out the information of a CCD by its associated electronic equipments is coined as readout time of that CCD.

The pixel size of the CCDs used in the ST and HCT is 24 $\mu$m. Particularly, for ST, the CCD is mounted at the f/13 Cassegrain focus of the telescope. Plate scale of the CCD chip is 0.38 arcsec per pixel, and the entire chip covers a field of 13 $\times$ 13 square arcmin on the sky. The gain and readout noise of the CCD camera are $10 \ e^-$ per analog-to-digital unit and $5.3 \ e^-$ respectively.

At room temperature and for a longer integration time (more than hundred milliseconds) thermal noise is created by the random generation of dark current.
Therefore, CCD must be cooled if the integration time is more than a few seconds to avoid the silicons self-generated dark signal filling the potential well. CCDs used for the present observations were cooled to about $-110^\circ\text{C}$ in a liquid nitrogen dewar. The typical CCD response is presented in Figure 2.2. The Quantum efficiency (QE) which is a function of wavelength seems to be an important characteristic of CCD. The figure shows that the response curves of CCDs used in this study all are peaked near $\lambda \sim 6700\text{Å}$.

Astronomical data are usually archived in a universally accepted common format, namely the ‘Flexible Image Transport System’ or FITS. However, because of different technical reasons the recorded data for different telescopes may vary. For e.g., in Sampurnanand Telescope the output comes in FTS – format. This FTS format is converted to FITS format using a routine ‘WR2FITS’.

### 2.2.4 The spectrograph

For the present study, two Indian facilities – Hanle Faint Object Spectrograph Camera (HFOSC) and IUCAA Faint Object Spectrograph & Camera (IFOSC) have been used along with other optical facilities throughout the world. Here HFOSC and IFOSC instruments will be briefly discussed.

IFOSC is the principal instrument currently mounted on the direct ‘Cassegrain-port’ of 2.0-m IGO telescope. This instrument was designed on the principle of very successful FOSC model [for detail see Buzzoni et al. 1984] which was developed at the Copenhagen University Observatory for the ‘European Southern Observatory’ (ESO) and ‘Nordic’ Telescopes. This technology is capable to take spectrum as well as images of the objects without changing the detector at the Cassegrain-
port. Either images or the spectra are produced on an EEV\textsuperscript{1} 2048×2048, thinned, back-illuminated CCD with 13.5\textmu m pixel size. The field of view is about 10.5′. Apart from Bessell $B,V,R,I$ imaging filters and photometric equipments, a set of Grisms\textsuperscript{2} with resolutions ranging from 190 to 3700 covering the wavelength between 3500−9000Å are also placed in the IFOSC instrument. For spectroscopy of a single object, the light coming from the object is allowed to pass through a slit and then it falls on the dispersive element like Grism (consisted with Grating and Prism), which disperses the light and produce spectrum on the CCD. The quality of produced spectrum is highly dependent on the convolved response of CCD and Grism along with the efficiency of the telescope and exposure-time. For present work a long-slit, with slit-width 1.5″ has been used along with three different grisms − IFORS5 (range: 3300−6300Å), IFOSC7 (range: 3800−6840Å) and IFOSC8 (range: 5800−8300Å).

The basic principle of HFOSC is similar to other FOSC instruments. HFOSC is a focal reducer, allowing for a larger field coverage for a given detector along with low and medium resolution grism spectroscopy by inserting dispersive elements between the collimator and camera. The CCD camera has 2048×2048 pixel elements, each of size 15\textmu m, working within the wavelength range 3500−9500Å. The spatial resolution 0.30″/pixel, whereas the spectral resolution is $\sim$ 1.6Å (medium resolution) to 50Å (low resolution) for a 1″ slit. Mainly two grisms − Gr.7 (range: 3800−8000Å) and Gr.8 (range: 5800−9200Å) have been used for spectroscopic monitoring of the supernovae.

\textsuperscript{1}The acronym stands for ‘English Electric Valve company ltd.’, who is developing these devices.
\textsuperscript{2}http://www.iucna.ernet.in/~ipt/etc/ETC/help.html#grism
2.2.5 The near infrared telescopes

For several supernovae near-Infrared (NIR) data have been acquired from different telescopes through scientific collaboration with different groups throughout the world. Mainly 0.6-m ‘Rapid Eye Mount (REM) telescope at La Silla, Chile [Zerbi et al., 2004] and 3.8-m ‘United Kingdom Infrared Telescope’ (UKIRT) located on the summit of Hawaii’s Mauna Kea [Hirst et al. 2006 and references therein] have been used to observe these catastrophe in $J$ (central wavelength $\sim 12200\,\text{Å}$, bandwidth $\sim 1620\,\text{Å}$) $H$ (central wavelength $\sim 16300\,\text{Å}$, bandwidth $\sim 2510\,\text{Å}$) and $K$ (central wavelength $\sim 21900\,\text{Å}$, bandwidth $\sim 2620\,\text{Å}$) bands. In REM telescope, ‘REMIR’ camera is used for imaging while ‘Wide Field Camera’ (WFCAM) is used as a detector for UKIRT telescope.

2.3 Optical and near infrared Photometry

For Optical and NIR image processing several softwares such as IRAF$^1$, M IDAS$^2$, DAOPHOT$^3$ etc. have been developed by different observatories to carry out different tasks. The images acquired from CCDs are contaminated by the combination of atmosphere, telescope, filter, the CCD response and the associated electronics and certainly need to be processed to upgrade the image quality before using for scientific work. Strong atmospheric turbulence, poor focusing and deformed optical system (geometric distortion) and charge diffusion in the detector can also generate

$^1$IRAF stands for Image Reduction and Analysis Facility distributed by the National Optical Astronomy Observatories which is operated by the Association of Universities for research in Astronomy, Inc. under co-operative agreement with the National Science Foundation

$^2$M IDAS stands for Munich Image and Data Analysis System designed and developed by the European Southern Observatory ESO in Munich, Germany

$^3$DAOPHOT stands for Dominion Astrophysical Observatory Photometry
a poor quality and distorted image. Image processing helps partially to compensate these deformation and to restore the original cleaned image. The softwares, mentioned above are also used to extract several physical parameters from the images. The three basic stages for image processing and optical photometric measurements are usually referred to as Pre-processing, Processing and Post-processing. Detail descriptions of these techniques are presented by Howell [2006].

2.3.1 Pre-processing

The patch of sky observed for scientific work is known as the raw frame. Apart from the raw frames it is also necessary to have bias frames (frames taken with zero second exposure to determine the noise level of the CCD) and flat frames (frames taken after exposing CCD for few seconds to correct for pixel-to-pixel variations in the CCD response as well as any non-uniform illumination of the detector itself). During pre-processing, bias-subtraction and flat-fielding of the object frames are done. A median of several bias images taken during observation is first subtracted from the raw object and flat frames. A median flat image in each filter is produced after taking several bias-subtracted flat frames. Since for median of a sample at least 3 parent elements are required, the acquired number of bias and flat frames in each filter should be \( \geq 3 \). The bias corrected object frame is flat fielded using the median flat field image. These two step correct the object frame for bias level and non-uniformity within pixel response due to atmosphere and observing system. Pre-processing has been done using IRAF task and briefly described below.

Bias frame subtraction
These frames are acquired after taking exposure for zero seconds when the shutter remains closed and the CCD is simply readout. These frames are used to determine the underlying noise level of the CCD during observing span. The bias value in a CCD image is usually a low spatial frequency variation throughout the array, caused by the CCD on-chip amplifiers. Bias frame contains both the DC offset level and the variations on that level. For a better statistical estimate, a median bias frame is generated from 3 or more raw bias frames. The IRAF-routine ‘ZEROCOMBINE’ is used to obtain a median bias frame.

**Flat fielding**

In practice all the pixels in a CCD do not show same response. With due course of time, dust and other physical distortion can appear on the exposed side of the CCD. Moreover, during opening and closing of the shutter, the central portion of the CCD gets more light in comparison to the edge. It is thus necessary to bring all the pixels to an uniform response level after considering these artifacts. For this, flat frames are obtained by illuminating the CCD with a uniform source of light in different filters. Flat field calibration frames are needed for each passband and different instrumental setup used for which object frames are taken. For a better statistical estimate, a median flat frame is generated from 3 or more bias subtracted flat frames and then further normalized after dividing the median flat by its median response. This is called normalized flat. Bias subtracted object frames are divided by normalised flat frame to correct for pixel-to-pixel variations in the CCD response. ‘FLATCOMBINE’ task in IRAF is used to get a median flat frame, ‘IMSTAT’ and ‘IMARITH’ routines are used to get median response and generation of normalized flat and flat fielded object frames. After FLATCOMBINE, the remaining
part of preprocessing can also be done through ‘CCDPROC’ task to get a cleaned object frame free from CCD noise and non-uniform pixel variations.

**Removing the cosmic hits**

Even after getting object frames free from CCD noise and non-uniform pixel variations usually bright specks in intensity, not spread over more than one pixel, are noticed in addition to stars in an object frame. These are because of cosmic rays. Stellar and cosmic ray intensity profiles are different from each other which is the underlying property used for the removal of cosmic rays from the object frames. ‘COSMICRAYS’ task is used for the removal of cosmic rays from the object frames free from CCD noise and non-uniform pixel variations.

**Producing dethard image for NIR imaging**

Since the earth radiates in infrared band, the near-IR sky is highly fluctuating. To remove this random fluctuation from near-IR images, always a number of images (≥3) are acquired after shifting the telescope little bit around the object. For a particular band, the median of the bias subtracted images gives a median profile of the sky fluctuation, which can be subtracted from every individual image to get clean NIR images. This process is called ‘detharing’. The other processes of data reduction are similar to optical photometry. Thus, for NIR imaging FLAT frames are not required, the median combined image itself behaves like a flat response. ‘IMCOMBINE’ task is used to get a median image and ‘IMARITH’ is used for individual subtraction.

**Alignment and stacking**
To improve the signal-to-noise ratio for faint objects, sometimes several frames are acquired with different exposures and then co-added. To do this, individual frames in a particular filter are first aligned with respect to each other and then either added or averaged (stacking), depending on the science objective. The alignment is done using the task ‘IMALIGN’ or ‘GEOMAP’ and ‘GEOTRAN’ tasks in IRAF. The stacking is done using task ‘IMCOMBINE’.

**Template subtraction – for supernova photometry**

When the supernova is embedded inside its host, the SN flux is expected to have a substantial contribution from the host-galaxy background. At early phases the SN flux dominates the total flux, thus with a PSF-fitting method (discussed below) gives more or less correct estimation of flux. At later epochs, the galaxy flux may brighten the SN light curves depending on its location in the galaxy. To remove the galaxy contribution, pre-SN image of the host or very late post-SN image of the host (when SN contribution becomes negligible), can be utilized as a template to map and estimate the background flux due to host. This technique is also used to discover the new transients and known as ‘Template subtraction technique’ [Alard, 2000; Alard and Lupton, 1998]. Self-written scripts employing IRAF tasks which included alignment, PSF and intensity matching of the galaxy template and SN images and subtraction of the template from the SN images were used to do this task and also verified with widely used ISIS tool. The effect of galaxy subtraction to get the true SN flux has been shown pictorially in figure 2.3. This has been applied on the field of galaxy NGC 2770 to determine the true fluxes (and hence the magnitudes) of two SNe 2007uy and 2008D. The detail discussion on the behaviour of these two SNe has been presented in chapter 5. The different kinds of host galaxy
subtraction will be discussed in different chapters according to their applicability.

2.3.2 Processing

Now, the cleaned and aligned images can be used for the main scientific work — to measure the observed flux of different objects from the images. The main processing is termed as photometry and involves the extraction of stellar positions and magnitudes from a given CCD image. *DAOPHOT II [Stetson, 1987]* is used to perform the photometry of individual stellar images recorded on the object frames. The different steps involved in processing are as follows.

**Object detection**

The automatic detection algorithm ‘*FIND*’ of DAOPHOT II package is used to detect star like objects above a certain user defined detection level, rejecting bad pixels, rows and columns in the object frame. The detection level is normally constrained as 3 times of the standard deviation of sky fluctuation. This routine detects and locates small, positive brightness enhancement within an image and able to distinguish stellar profile from galaxy or any other extended objects profiles, cosmic rays or any other high energy particle profile. The image is convolved with analytical stellar profile — either Gaussian/Lorntzian/Moffat/Penny with a given FWHM, provided by user to locate a star in a pixel by fitting the profile to the brightness values of surrounding sub-array of pixels. The stars are located by performing a pixel by pixel search and selecting locations where the Gaussian profile fit is good i.e. where the central height of the best fitted Gaussian model achieves a large, positive value possibly lying near the center of star image. The output of *FIND*
Figure 2.3 The necessity of Template subtraction. The **upper-left panel** shows the field of galaxy NGC 2770, hosting the SNe 2007uy and 2008D (see chapter 5). The **upper-right panel** shows the field after template subtraction. The effect of galaxy is subtracted. **Lower panel** shows the visual light curve of SN 2008D. The green points are found without doing the galaxy back-ground subtraction while the red dots represent the light curve after galaxy subtraction. The effect of galaxy is less when the SN in bright (less magnitude) and more when the SN is faint (high magnitude).
catalogues the stars of the frame by mentioning their X and Y positions.

**Photometry**

The next step is to get the observed flux in the form of instrumental magnitude of the stars detected by *FIND*. By photometry, a quantitative (numerical) values for the brightness of objects are obtained. Two basic steps for obtaining photometric data from CCD are: ‘Aperture Photometry’ and ‘Profile Fitting Photometry’. For bright objects and less-crowded field Aperture Photometry is preferred, while for faint object and crowded field Profile Fitting Photometry is essential.

*Aperture Photometry*

Aperture photometry is used to measure the brightness of an object without considering the possible contamination due to sources such as sky, pixel defects or other nearby stars and galaxies. If a CCD frame contains few stars, it is possible to perform photometry of individual stars by simply defining a circular geometric region as an ‘aperture’. Total photons detected within an aperture centered around a star is calculated after summing-up the counts within the aperture. Similarly the sky photons detected within a star free region of the same area elsewhere in the frame is counted. The difference between these two numbers corresponding to a particular star are converted into magnitude scale. For this ‘PHOT’ routine in DAOPHOT II is used. This routine calculates the star counts within a series of increasing concentric apertures and also in an annulus to evaluate the count of star free sky region that are defined by user. Similar routine PHOT in IRAF can also be used to do aperture photometry with equivalent results. The output lists the magnitude of the star in different apertures. For the faint stars and in crowded regions, aperture
photometry does not work properly.

Profile Fitting Photometry

This technique is essential for faint objects and crowded field, where outer sky is not well defined or contamination due to nearby sources is highly possible. This technique offers the best possible recovery of photometric information for stellar objects. In an optical image all the stars are affected due to identical observing conditions and have the same profile and would differ from one another only by an intensity scaling ratio. This property is the basic principle of the profile fitting photometry. In a perfect telescope having no instrumental defect and in the absence of atmospheric turbulence the image can be represented by the profile governed by the diffraction pattern of the telescope optics. In practice, the stellar image is affected by the irregularities of the instrument (telescope observations, tracking errors, optics etc.) and the seeing during ground based observations. So, a model of ‘point spread function’ (PSF) is determined from several stars in a given frame. PSFs can be modeled by a number of mathematical functions to describe the stellar brightness profile. Mostly Gaussian, modified Lorentzian or Moffat functions are used in this regard [Stetson et al., 1990].

Thus, the analytical functional forms of PSF can be used to model the PSF for a star within an image, assuming that they provide a good representation of the data itself. Sometimes a combination of one or more of these functions is used to obtain a better fit. However, we see that for CCD frames having good image quality observed under good seeing, the $\chi^2$ value in fitting Gaussian analytical function is better than that obtained from other two functions. The analytical method has the advantage of integrating numerically over the entire stellar image but the disadvantage being
that the function used to fit the profile can only be approximated. Other method of PSF determination is empirical, where it is possible to simply store the observed profile of several bright stars as data array $O(i,j)$. The empirical method has the advantage that it uses the observed profile, but the drawback in this case is that near the central region of the stellar image, the brightness varies too rapidly between the adjacent pixels to be interpolated correctly. Better results can be obtained by combining these two methods. In DAOPHOT II model PSF is generated using the routines ‘PEAK’ and ‘PSF’.

To get magnitude of all the stars in a field, the modeled PSF is applied to all the stars of the frame with the help of ‘ALLSTAR’ routine. It is a standalone programme which performs a multiple-profile fit to all the stars in a frame. This routine returns the positions, magnitudes, error in magnitudes, sky brightness, no. of iterations, $\chi^2$ values and sharpness values of all stars present in the CCD frame. PSF and ALLSTAR tasks in IRAF are also used sometimes to do PSF photometry with equivalent results.

**Conversion of counts to magnitude**

During both aperture and profile fitting photometry the flux are at first measured in terms of counts (in ADU) and then converted to magnitudes. As stated above, in aperture photometry, the magnitude of the star corresponding to any aperture is determined by summing the data inside that aperture and from the relation, $m = zpt - 2.5 \log_{10} I$, where $zpt$ is the arbitrary zero-point and $I$ is the true counts of the stars, after subtracting the sky contribution and expressed as $I = (\Sigma I_{ij} - n_{pix} i_{sky})$, where $n_{pix}$ is the number of sky pixel, $i_{sky}$ is the count of individual sky pixel and $I_{ij}$ is the total count of a particular aperture. The similar
approach is also adopted for profile fitting, where pixels within the PSF profile are considered. Similar algorithm is also applied in IRAF tasks. In DAOPHOT II assumes $zpt = 25.0$ mag which corresponds to 1 ADU.

2.3.3 Post-processing

The output of the processed frame provides the instrumental magnitude of the star like objects of that frame. To convert them to the standard magnitude indices, Post-processing is done. The steps are as follows.

Aperture correction

Aperture photometry is performed for several selected, bright, isolated stars in each observed frame using a series of apertures of increasing size. In PSF photometry, the instrumental magnitude of a star comes from the height of a model point-spread function for that frame scaled to the intensity values recorded inside the stars image. This magnitude is restricted to the aperture chosen and has to be corrected for the counts left out in the wings of the stellar profile. The correction from profile fitting magnitudes to aperture magnitudes is carried out by the process of determining aperture growth curve. This correction is known as ‘aperture correction’ and applied to the PSF magnitudes to get the aperture instrumental magnitudes. This is done by ‘DAOGROW’ routine [Stetson et al., 1990].

Cross Identification

To determine the standard magnitudes and colours, cross identification of the stars of different frame and filters is required. This is performed after using ‘DAOMATCH’
programme of DAOPHOT II. This routine considers about 30 brightest stars between the two frames and provides the coordinate transformation equations between the various frames with respect to a reference frame. Using the approximate transformation equation provided by \textit{DAOMATCH}, another programme ‘\textit{DAOMASTER}’ is then used to identify all the stars in a field and to cross-match all the stars by spatial proximity. If, after transformation to the coordinate system of the master frame, a star lies within the specified distance of a star in the master list, it is provisionally identified with that star; if it lies near no star in the master list it is added to the list as a possible new detection.

\textbf{Transformation to the standard system}

To transform the instrumental magnitudes to the standard magnitudes a set of transformation equations are used.

\begin{align*}
    u_{CCD} &= U + a_0 + a_1.(U - B) + a_2.X \\
    b_{CCD} &= B + b_0 + b_1.(B - V) + b_2.X \\
    v_{CCD} &= V + c_0 + c_1.(B - V) + c_2.X \\
    r_{CCD} &= R + d_0 + d_1.(R - I) + d_2.X \\
    i_{CCD} &= I + e_0 + e_1.(V - I) + e_2.X
\end{align*}

Here \( a_0, b_0, c_0, d_0, e_0 \) are the zero points; \( a_1, b_1, c_1, d_1, e_1 \) are the colour coefficients; \( a_2, b_2, c_2, d_2, e_2 \) are the earths atmospheric extinction coefficients and \( X \) is the air mass. \( U, B, V, R \) and \( I \) are the standard magnitudes and \( u_{CCD} \).
$b_{CCD}$, $v_{CCD}$, $r_{CCD}$ and $i_{CCD}$ are the corresponding aperture instrumental magnitudes. The second order colour correction terms are ignored as they are generally small in comparison to other errors present in the photometric data reduction.

The light of celestial objects get reduced during its propagating through the terrestrial atmosphere due to absorption and scattering. The reduction of incident flux is called *atmospheric extinction*. It depends on the atmospheric path length through which the light is travelling and hence on the zenith angle ($Z$) of the object, wavelength of radiation, atmospheric conditions during the observation and altitude of the observatory. Atmospheric extinction correction is an integral part of photometric data reduction. For this, the air mass ($X$), i.e., the thickness of the atmosphere crossed by the light rays along the line of sight, is calculated using the equation [Hardie and Ballard, 1962]

$$X = \sec Z - 0.0018167 (\sec Z - 1) - 0.002875 (\sec Z - 1)^2 - 0.0008083 (\sec Z - 1)^3$$

and

$$\sec Z = \left[ \sin(\phi) \sin(\delta) + \cos(\phi) \cos(\delta) \cos(h) \right]^{-1}$$

Where, $\phi$, $\delta$ and $h$ are observers latitude, the declination and the instantaneous hour angle of the star respectively. By definition, air mass value is 1 at the zenith and increases with the zenith angle. The atmospheric extinction coefficients for each filter are determined from the plot of aperture magnitudes observed versus air mass for the comparison stars on each night using the method of linear least square fit.

The transformation equations and DAOPHOT II routines ‘COLLECT’, ‘CCDSTD’ are used to determine the values of transformation coefficients. Using the obtained
Figure 2.4 Standardization of SN 2010jl field using the standard stars of the field SA98. The detail properties of this object has been discussed in chapter 6.

Transformation coefficients and the standard magnitude values from Landolt [1992], secondary standards in the object frame can be generated. ‘CCDAVE’ program in DAOPHOT II is used to generate the secondary standards. The final transformation of all the stars in the object frame is done by the ‘FINAL’ routine. This program determines a single consistent photometric zero point using all the frames from the local secondary standard stars. Thus, a list of stars and the corresponding positions, standard magnitudes and colours for a given CCD frame of interest is obtained.

After performing the standardization of the observed field, it is necessary to check the quality of this performance. The best way is to check how much accurately we are able to reproduce the magnitudes of the photometric standard found by Landolt [1992]. This can be concluded by plotting the difference between standard
magnitudes and the reproduced magnitudes of these photometric standards against their standard mags and colours. Normally these are plotted against $V$-band mag and $(B-V)$ colours. If the standardization process goes well all the differences should be around zero and within a scatter (standard deviation) $\sigma \lesssim 0.05$ mag. For all the objects, discussed in this work this process have been performed. Figure 2.4 represents the typical nature of this plot. Field of CSS 100217 (see chapter 6) has been calibrated using the stars from the Landolt standard field SA98. The plot shows that the calibration process went very perfectly.

### 2.4 Optical Spectroscopy

During acquisition, optical spectra are also obtained in the form of two-dimensional CCD images. Like photometric data, the long slit spectroscopic data also consists of bias frames, flat frames and object frames along with other calibration frames. These frames are acquired for each observing night. The object frame contains the image of the particular object under interest, focussed at a slit. Unlike photometry, this image is binned over the entire wavelength regime of the optical band (of course, dependent on instrument setup). This image is made by dispersing the light rays of the target on the CCD frame along a particular direction known as ‘dispersion axis’. This normally lies either x-axis or y-axis of the CCD when it is properly aligned with the slit axis. In order to avoid any degradation of resolution as well as the sky background contribution, the slit widths are usually kept equal to the FWHM of the stars seeing limited image. Thus, the spatial profile of the star is governed by the seeing. So, after reduction, the pixels along the dispersion axis give a measurement of wavelength, while the pixels along the spatial axis give a
measurement of counts/flux. To convert these pixels and counts in wavelength and flux respectively, calibration frames are required. ‘Arc’ and ‘standard star’ frames are required for wavelength and flux calibration respectively.

Spectroscopic data reduction has been done using the routines of ‘SPECRED’ package provide in ‘IRAF’. The basic pre-processing is similar to photometric reduction (except detharing and template subtraction). The cosmic ray rejection task can be improved using the Laplacian kernel detection method [van Dokkum, 2001]. After getting a cleaned 2D frame, the 1D spectrum is extracted and then calibrated. The three main steps are ‘Aperture Extraction’, ‘Wavelength Calibration’ and ‘Flux Calibration’. These are briefly discussed here.

2.4.1 Aperture Extraction

The one dimensional spectra are extracted using ‘APALL’ task. This is based on the extraction algorithm proposed by Horne [1986]. APALL is a collection of many tasks. After selecting the aperture around the object part and background, which comes from sky regions, it traces the spectrum along the dispersion axis and sums it along a one-dimensional line. This routine eliminates the sky noise, delivers maximum possible signal-to-noise ratio and takes care of the effects of moderate geometric distortion and cosmic ray hits. The output comes as an one dimensional spectrum in the form of pixel vs intensity values.

In order to calibrate the pixel number in terms of wavelength, it is necessary to take spectra of a laboratory standard source (e.g., Fe-Ar and Fe-Ne, He-Ne Arc-lamp). Similarly to convert the counts to flux spectra of spectroscopic-standard stars [Hamuy et al., 1994; Massey et al., 1988] are taken during the same observing
night. Extraction of 1D-spectra of spectroscopic-standard star and that of Arc-lamp are also done through \textit{APALL}.

### 2.4.2 Wavelength Calibration

The 1D spectrum of the Arc-lamp is now used to convert the pixel to wavelength for both 1D spectra of object and standard star. Since Arc-lamp is a laboratory standard source, wavelength of different spectral features such as emission in Arc-spectrum is known and provided by the observatory as an ‘\textit{Arc Atlas}’. The emission lines of 1D extracted spectrum of the Arc is then identified using that atlas by running the ‘\textit{IDENTIFY}’ task which calculates the pixel-to-wavelength solution for Arc spectrum. ‘\textit{HEDIT}’ task is used to incorporate this information in 1D extracted spectra of object as well as of standard. Task ‘\textit{DISPCOR}’ is then executed to apply this solution on the spectra of standard and object simultaneously. It gives dispersion-corrected spectra for object and standard. Usually a high order polynomial is employed to fit the identified pixels against wavelength. It is always better to check the wavelength calibration before going for flux calibration. For this the standard atmospheric emission lines like 5577Å, 6300Å and 6364Å are used. If any correction is required then the necessary offset can be applied to dispersion-corrected standard and object spectra using the task ‘\textit{SPECSHIFT}’ or by applying the offset to the \textit{header} – \textit{adverb} ‘\textit{CRVAL1}’ mentioned in the header of dispersion-corrected spectra. ‘\textit{HEDIT}’ task is used for this purpose.
2.4.3 Flux Calibration

After getting the wavelength calibrated spectra of standard and object, they are calibrated in flux after knowing the standard-flux of spectroscopic standards. These standard fluxes are provided in IRAF itself and can also be found in the web interface of several observatories like ESO. Flux calibration is performed using the tasks ‘STANDARD’, ‘SENSFUNC’ (used to fit a polynomial to the observed magnitude as a function of wavelength after applying extinction correction in the standard and programme star spectra) and ‘CALIBRATE’. The final flux calibrated one-dimensional spectra give distribution of flux over wavelength.

2.5 Acquisition and reduction of Swift/UVOT data

The Ultraviolet Optical Telescope (UVOT), on-board the Swift satellite, is a 30 cm modified Ritchey-Chretien Ultraviolet (UV)/Optical telescope Co-aligned with the X-ray Telescope and mounted on the telescope platform common to all instruments of Swift. The photometric data from UVOT, was used to access the UV and Optical data. This is the only space-based facility that has been used for this work. UVOT provides simultaneous ultraviolet and optical coverage (170–650 nm) in a $17' \times 17'$ field. Despite its limited aperture, UVOT is a powerful complement to other instruments because of its UV capabilities and the absence of atmospheric extinction, diffraction, and background.

UVOT has six photometric filters, three near-UV (NUV) bands $-uvw2, uvm2, uvw1$ and three Optical bands $-u, b, v$. It is also equipped with a White filter, having a response curve extended between NUV and NIR. The detail information of UVOT filter system can be found at Poole et al. [2008]. The UVOT data has be obtained
from the *Swift* Data Archive through NASA’s *High Energy Astrophysics Science Archive Research Center* (HEASARC) on-line services. The photometry was done using standard processes of HEASOFT routines.

The *Swift* data have had bad pixels identified, mod-8 noise corrected, and have been transformed into FK5 coordinates. We used the standard UVOT data analysis software distributed with HEASOFT 6.10 along with the standard calibration data. As long as the source had a count rate greater than 0.5cts/s, photometry was done using the command ‘*uvotsource*’ with a standard circular aperture of radius 5″ and a circular background region with a radius of 15″. Below this threshold, a 3.5″ radius was used and an aperture correction was applied. The background region was selected to have similar background properties to those at the location of the transient, and to be free of contaminating sources. The photometry was calibrated to the UVOT photometric system following the procedure described in Poole et al. [2008]. To remove the contribution from the underlying host galaxy we measured the host galaxy flux at the position of the SN from late UVOT observations, where there was no longer a contribution from the SN. This additional flux was then subtracted from SN photometric measurements. For all observations the source was close to the centre of the field-of-view, and differences in the PSF between observations were, therefore, negligible. A detail description of UVOT photometry can also be found elsewhere [Brown, 2009].

### 2.6 Data acquisition at Radio wavebands

The archival radio data have been used for analysis of Type Ib SN 2007uy. Mainly archival data from radio telescope VLA has been used in this regard. VLA [Thomp-
son et al., 1980] is a radio astronomy observatory located on the Plains of San Agustin, between the towns – Magdalena and Datil, about fifty miles (80 km) west of Socorro, New Mexico, USA (Latitude = 34°04′43.497″N, longitude = 107°37′03.819″W). It is consisted with 27 independent antennas, arrayed along the three arms of a Y-shape, each of which measures ~21 km/13 miles long. Each antenna is 25 m in diameter and the weight of each solid dish is about 230 tons (see figure 2.5). There are four array configurations. The maximum antenna separations are respectively 36 km, 10 km, 3.6 km and 1 km for A array, B array, C array and D array. These configurations change in every four months or so.

The earth atmosphere is transparent with respect to radio waves having wavelength range between ~1 cm to few tens of metre. So, by principle big reflecting telescopes can be used to get an image of distant celestial objects emitting in radio waves. The procedure for image production at different frequencies will be of
course different from optical telescopes, though they are mathematically identical. In fact, many such huge reflectors had been used (and still being utilized) to image the radio sky. Some of them are – 300-m Arecibo radio telescope, USA; 600-m RATAN radio telescope, Russia; 530 m long and 30 m wide cylindrical-paraboloid radio telescope at Ooty (ORT), India; 10.4-m Caltech Submillimeter Observatory (CSO), USA etc. On the other hand according to Rayleigh’s criterion the angular resolution of a telescope is ultimately diffraction limited by the expression $\theta \sim \frac{\lambda}{D}$.

Where $\lambda$ is the observed wavelength and ‘D’ is the measure of aperture size (size of reflector/refracter) (The actual resolution of optical telescopes is however limited by atmospheric fluctuation, known as ‘seeing’). Hence, to get a better angular resolution at higher wavelength, larger aperture size is required. For e.g., while the human eye has a diffraction limit of $\sim 20''$ and the modest optical telescopes have this value as low as $0.1''$, the large ($\sim 300$-m) single dish radio telescope has angular resolution of only $\sim 10'$ at 1 metre wavelengths. To increase the angular resolution in radio wave, interferometry technique is adopted.

The telescopes like VLA, Giant Metrewave Radio Telescope (GMRT), Westerbork Synthesis Radio Telescope (WSRT) and all other upcoming radio telescopes work / will work on the basis of interferometry. In this technique, each single dish antenna behaves like a slit and each combination of two antennas behaves like a setup producing double-slit diffraction pattern, where the projected distance between two antennas toward the object is equivalent to slit separation. A very large area is covered by several small antennas and after interferometry, they act coherently like a big telescope. The observed quantities in a radio-interferometric observations are the complex visibilities which is the fourier transformation of the sky brightness distribution. A radio-interferometer consists of several antennas (say $N$ antennas)
and each pair records one Fourier component. Hence at any given instant of time only $^\text{N}\text{C}_2$ Fourier components are recorded. With the motion of the object on the sky from east to west, projected separations between pair of antennas changes and within a moderate time span, a considerable number of discrete samples of visibilities are observed and hence the sky brightness distribution can be reconstructed. This procedure is known as ‘earth rotation aperture synthesis’ and is used widely to synthesize large aperture telescope even with a finite number of antennas. The basic parameters for VLA are mentioned in table 2.3. For the analysis of SN 2007uy radio data mainly ‘C’ and ‘X’ band VLA archival data, along with other literature data have been used.
Table 2.3 Basic system parameters of Very Large Array (VLA)

<table>
<thead>
<tr>
<th>Characteristic</th>
<th>4 Band</th>
<th>P Band</th>
<th>L Band</th>
<th>C Band</th>
<th>X Band</th>
<th>U Band</th>
<th>K Band</th>
<th>Q Band</th>
</tr>
</thead>
<tbody>
<tr>
<td>Freq (GHz)</td>
<td>0.07–0.074</td>
<td>0.30–0.34</td>
<td>1.34–1.73</td>
<td>4.5–5.0</td>
<td>8.0–8.8</td>
<td>14.4–15.4</td>
<td>22–24</td>
<td>40–50</td>
</tr>
<tr>
<td>Wavelength (cm)</td>
<td>400</td>
<td>90</td>
<td>20</td>
<td>6</td>
<td>3.6</td>
<td>2</td>
<td>1.3</td>
<td>0.7</td>
</tr>
<tr>
<td>Primary beam (')</td>
<td>600</td>
<td>150</td>
<td>30</td>
<td>9</td>
<td>5.4</td>
<td>3</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>Highest resolution (&quot;)</td>
<td>24.0</td>
<td>6.0</td>
<td>1.4</td>
<td>0.4</td>
<td>0.24</td>
<td>0.14</td>
<td>0.08</td>
<td>0.05</td>
</tr>
<tr>
<td>System Temperature ($^\circ$K)</td>
<td>1000–10,000</td>
<td>150–180</td>
<td>37–75</td>
<td>44</td>
<td>34</td>
<td>110</td>
<td>50–190</td>
<td>90–140</td>
</tr>
</tbody>
</table>
2.7 Data reduction strategy at Radio wavebands

In terms of mathematical-algorithms, the processing and analysis of Radio data is similar to Optical, though due to a completely different ‘receiver/detector – technology’ and presence of different kinds of artifacts than optical, reduction process and used softwares are different from optical. The software that is widely used to reduce and analysis of ‘radio interferometric data’ is ‘Astrophysical Image Processing System’ (AIPS)\(^1\). Unlike IRAF, working with radio interferometric data, requires to bring the data under AIPS-environment, before starting any processing. To bring the processed data, outside AIPS also requires particular task, so that it can be fetched from AIPS-environment.

The radio data may or may not be in the FITS format and provided by the observatories mostly in the form of a ‘data – cube’, containing data and informations of multiple observations. In VLA archive, the data are not provided in completely binary format. This data have been downloaded to do further analysis. In principle the spatial correlation function or the visibility, \(V(u,v)\), measured across several baselines using an interferometer can be inverse Fourier transformed to reconstruct the sky brightness distribution \(I(l,m)\). However, the measured visibilities are affected by man made interferences and is scaled by the antenna gains etc. Because of these a series of analysis techniques has to be applied on them before any scientific analysis.

To calibrate the flux of a radio source and its brightness distribution over the sky plane, a standard calibrator is required. After several measurements, NRAO/VLA had chosen few point like objects (not resolvable in the VLA operating radio frequencies) in the radio-sky as flux calibrator. On the other hand, the radio wavelength

\(^1\)Astrophysical Image Processing System (AIPS) has been developed by the National Radio Astronomical Observatories (NRAO), USA
is affected due to earth’s ionospheric turbulence. This turbulences are distributed in the form of patches and every radio observation is scattered when it passes these turbulent zones. This is similar to atmospheric extinction for optical observations. Angular dimensions of these turbulent zones are within few degrees. Since the number of flux calibrator are limited, a secondary calibrator (also listed by NRAO) is used to calibrate the target object. This secondary calibrators are called phase calibrator. While observation, the phase calibrators are chosen in such a way that their angular separation from object remains within few degree. This confirms that both phase calibrator and target are affected due to same turbulent zone. The phase of this secondary calibrator is well known for a given VLA frequency. During calibration the secondary (phase) calibrator is calibrated in flux by the flux calibrator and then the target source can be calibrated in both flux and phase using flux and phase calibrators. If the flux calibrator is located within few degree away from the target object, then for that object use of phase calibrator is not required; only flux calibrator can be used for this purpose.

VLA provides a preliminary calibrated data. This will be discussed further in subsequent sections. The SN radio data reduction in AIPS can broadly be divided in four categories – ‘Pre-processing’, ‘Data editing’, ‘Calibration’, ‘Imaging’.

2.7.1 Pre-processing

The Radio data are first loaded in AIPS-environment. If the the data is in FITS-format task ‘FITLD’ is used. For the VLA data, which was not in FITS format, task ‘FILLM’ is applied. Thus, the basic purpose of these two tasks are same – to bring the Radio data in AIPS-environment.
For loading the data, indexing and a primary calibration may be required for few cases, like low frequency GMRT data. For this task ‘INDEXR’ is used. For VLA data, primary calibration is provided by the observatory, so this task is not required. The listing of all the individual data present in the loaded data cube is done through the task ‘LISTR’. The list should indicate the object data, flux and phase calibrators (surveyed and listed by NRAO in VLA website), their coordinates, observing sequence over time, the observing frequency along with bandwidth and the polarization. So, the data is now ready for further processing.

2.7.2 Data editing

Radio astronomical measurements particularly at low frequencies are highly affected by the man made radio signals like satellite and mobile phone communication and use of radio receivers near the observatory. These signals basically create ‘Radio Frequency Interference’ (RFI) with the radio-signal coming from the celestial object and makes the data noisy. A good quality image is possible only when the data is fully free from RFI. The process of editing the radio data to remove the RFIs is commonly known as ‘flagging’. Many tasks like ‘UVPLT’, ‘UVFLG’, ‘TVFLG’, ‘VPLT’, ‘WIPER’ are used to look the UV data and to identify the bad signals and RFIs. This is usually done minutely to preserve the good data after removing the bad portion. The defects of data set comes in different form - antenna or baseline based or time based (and also channel based for GMRT). The data editing is first applied on the point sources – flux and phase calibrators and finally on target if required. The visibility amplitude of an unresolved point source should remain constant as a function of baseline length. By the task UVPLT visibility
amplitude vs UV distance can be plotted to locate the inconsistent data points. Using task `V PLOT` the particular antennas or baselines which are responsible for those discrepant points can be located and the task `UVFLG` can be used to flag the contribution of those antennas or baselines either for whole or particular time range within which they were behaving abruptly. If any antenna or baseline remain dead through out the observation of both flux and phase calibrator then it has to be removed from entire data set by the task `TVFLG` as well. Sometime, first few minute data of a scan get corrupted. Task ‘QUACK’ is used to remove this corrupted part from entire data set.

### 2.7.3 Calibration

Through calibration, observed visibilities are corrected to the true visibilities. At the time of observation we observe flux calibrator (∼ 15 minutes) in the beginning and at the end of the observation and the target source observations are interspersed with observations of the secondary phase calibrator. The desirable flux density calibrators have to be unresolved so that they will appear same in all baselines. Furthermore flux density has to be known and be constant with time which helps to find out any systematic atmospheric or instrumental errors [Fassnacht and Taylor, 2001]. Inhomogeneous and dynamic earth ionosphere is responsible for rapid change of refractive index with time and position for radio waves. The ionosphere adds an differential phase rotation to the complex visibilities. To calibrate this atmospheric phase fluctuations we need a secondary calibrator which is near (∼ 20°) to the target source and whose true visibility is known. In calibration process observed visibility is to be divided by the true visibility to get the gain factor. This is applied as a
correction to the visibilities of the target source. To get the gain values in between calibrator observations interpolation needs to be done. However if a flux calibrator is near by our target source then secondary phase calibrator is not needed.

The calibration process starts with setting the flux density of the flux calibrator using the task ‘SETJY’. This task follows standard Baars et al. [1977] formulae and takes frequency information from the header to determine the flux densities of primary calibrators. The task ‘CALIB’ is used for the selected single channel to determine the antenna based complex gains for the calibrators. This creates a solution (SN) table which contains the complex gain solutions. The SN table can be plotted by the task ‘SNPLT’ to check the gain solutions. If the solutions are not reasonable then one can repeat the same process starting from the flagging. The flux density of the phase calibrator was computed by the task ‘GETJY’. This creates a SU table and corrects the amplitude in the SN table. The flux density of the phase calibrator are cross checked with value given by VLA calibrator list. If it differs too much then more flagging is needed for the phase calibrator. The task ‘CLCAL’ is used to interpolate the gains of the target with time. This creates CL table version two which contains solutions derived from SN table by linear vector interpolation. The amplitude and phase deviations of all the flux and phase calibrator has to be checked applying CL table in ‘UVPLT’. If there is any significant deviation then one has to identify the bad data and flag them. Deleting the SN tables and CL table 2 same calibration process has to be repeated again.

The calibration and flagging are repeatedly done in an iterative way to remove the RFI and calibrate the good data in flux and phase. After a complete calibration the target data is fetched out from the data cube using the task ‘SPLIT’. This is further utilized to make a radio image.
2.7.4 Imaging

The visibilities measured by all antennas are all discrete points in uv-plane. ‘Dirty beam’ is the Fourier transform of uv sampling function or uv tracks during our observation. ‘Dirty map’ is convolution between real brightness distribution of the source and Dirty beam. The Dirty beam and the Dirty map comes from fourier transformation of sampling function and visibilities by ‘Fast Fourier Transform’ (FFT) algorithm. The true image can be achieved by deconvolution of Dirty map. The most popular deconvolution algorithm CLEAN was proposed by Högbom [1974] and widely used for imaging of point sources like supernova. The task ‘IMAGR’ is used to make the Cleaned image from a Dirty image.

The supernova flux can be measured from the leaned image, obtained after running IMAGR.