MICROWAVE OBSERVATIONS OF W40

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By

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Abstract

Observations of the W40 cloud complex with the Cosmic Background Imager (CBI) at 31 GHz are presented. The data were taken over a span of two years (2006 - 2008). Data reduction and calibration was performed with the aid of the "inhouse" software CBICAL. Three different types of calibration are applied to the data: quadrature, noise and antenna calibration. Difference mapping is used to extract flux from the visibilities. The final CLEANed combined map of W40 at 31 GHz after spillover rejection is presented. The synthesized beam is equal to 4.4×3.8 arcmin and is limited by dynamic range of 321 to 1. It is found that the W40 region consists of a compact source of size 5.8 ± 0.2 arcmin. A low level extended ionised gas that connects W40 to the Galactic plane is also detected from inspection of low frequency (~ 1 GHz) radio surveys. The temperature of the source is 57 ± 4 K. The integrated flux density of W40 is 27.7 ± 1.4 Jy at 31 GHz, as determined by CBI at 5 arcmin resolution. No evidence of extended emission is found. The spectrum of W40 is constrained by the values of its flux density as they were measured by the Effelsberg, WMAP, IRAS, and GB6 surveys at a frequency range between \sim 1 to 5000 GHz. The flux densities are extracted by means of Gaussian fitting and aperture photometry. Gaussian fitting exhibits less sensitivity in its measurements under a variant background and therefore is considered a more robust method of measuring the flux density of W40. We find that a power law with an optically thin free-free spectrum ($\alpha = -0.12$) in combination with a modified blackbody curve with a thermal dust emissivity index ($\beta = 0.1 \pm 0.01$) can sufficiently describe the spectrum of W40. The value of the emissivity is flatter than expected implying that a multiple dust component is more suitable at higher frequencies. We also find that optically thick emission from ultra-compact HII regions is not expected to contribute to the spectrum of W40 using high resolution NVSS data. The upper limit of the flux density at 31 GHz was calculated to be 5 Jy with a 2σ confidence level at 5 arcmin resolution. A tighter limit was derived at 9.4 arcmin, a value of 4.4 Jy. This value cannot account for any spinning dust emission. We interpret this absence as possible destruction of the population of the small grain sizes due to excess heat resulting from active star formation occurring inside the cloud complex. Conversely, for the WMAP data a small bump is observed and the upper limit at 33 GHz is 5.4 Jy with a 2σ confidence level. This only qualifies as a noise fluctuation.

> University of Manchester, **Dimitrios Stamadianos** Master of Science Microwave Observations of W40 06 September 2010

Declaration

No portion of the work referred to in this thesis has been submitted in support of an application for another degree or qualification of this or any other university or other institution of learning.

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Chapter 1

Introduction

In Section 1.2 of this chapter we discuss the general principles behind the radiation emitted by astronomical objects. In Section 1.3 we describe each foreground starting from free-free emission and moving to synchrotron, thermal (vibrational) dust and the anomalous foreground as well as emission from spinning and magnetized dust grains. Section 1.4 and Section 1.5 look into the main assumptions/relations in the field of synthesis imaging and the Cosmic Background Imager interferometer array respectively. Section 1.6 presents the most significant properties of the W40 cloud complex and Section 1.7 concludes the chapter.

1.1 CMB and Foreground contamination

Penzias & Wilson (1965) describe the CMB as "Black-body radiation filling the universe" and actually this definition is perfectly fitting as the CMB radiation (CMBR) dominates the radiation content detected in the universe. Most importantly, the CMB is the most solid evidence that the Big Bang actually occurred as it provides the early picture of Universe. This is easy to understand if one thinks that the further away we look from the Earth, the more we look back in time (due to the finite speed of light). Since we know that the universe expands, in the past (*i.e.* shortly after the Big Bang) it must have been contained in a much smaller volume and as such its temperature would have been larger. That is exactly how the CMB



Figure 1.1: CMB and foreground anisotropies as functions of frequency plotted on the same axis. Synchrotron dominates at low frequencies (< 40 GHz), free-free at 53 GHz while dust emission at high (>90) frequencies. GHz (Bennett et al., 2003)

is observed from the Earth: a hot, dense "fireball" that expanded to the cooler and more stable state now surrounding us.

The CMB anisotropies can be best observed by means of radio interferometry, where an array of telescopes connected to each other ("interferometer") is used to make observations of astronomical objects. These observations will most certainly be hindered by "foreground contamination". This means that the observed signal will not be entirely due to the CMB anisotropies/fluctuations but it will be affected by other types of emission. It has been established that this type of "contamination" originates from diffuse Galactic emission, unresolved point sources and the Sunyaev - Zeldovich (SZ) effect (Tegmark et al., 2000).

Therefore, correctly identifying the characteristics of the foreground contaminants and their respective importance is crucial, since it enables their accurate removal. At the moment, the scientific community accepts the existence of three types of diffuse Galactic emission: free-free, synchrotron and thermal (vibrational) dust. These are presented in **Figure 1.1**. Additionally, there is a debate over the existence of a fourth "anomalous" component related to, most possibly, spinning dust grains. The first two components are important for frequencies below 60 GHz; for frequencies above this value, thermal dust becomes the dominant contributor (de Oliveira-Costa et al., 1997). The above statement should, theoretically, be sufficient to understand which foreground is responsible for the respective signal we detect. In practice though each foreground's frequency limits are not invariant under the region of observation and, as such, "frequency coherence" models have been developed to account for this inconsistency (Tegmark et al., 2000). In principle, the spectral index of each foreground (commonly defined by Equation 1.4) should also be determined as accurately as possible. Further complications may arise when the foreground components mix with each other, or "correlate", thus making distinction almost impossible unless a template for each component is used.

There are two types of the SZ effect; thermal and kinetic. According to Tegmark et al. (2000) and Holder & Carlstrom (1999) only the thermal component should be treated as a foreground. The thermal SZ effect, observed on the CMB, arises from the energy transfer between highly energetic electrons (originating from hot, ionized gas) and CMB photons via collisions *i.e.* via inverse Compton scattering. The CMB photons, having lower energy, will become excited and the overall effect will be distortions in the CMB spectrum. There is not a particular frequency range that this effect dominates; rather it depends on the region of observation and whether there are any galaxy clusters or other large scale structures nearby (Tegmark et al., 2000).

As far as point sources are concerned, they become more significant as one moves to smaller angular sizes ($\leq 1^{\circ}$). At these sizes, their removal presents many complications and becomes a problem more important than any foreground. Tegmark & de Oliveira-Costa (1998) state that this is mainly because there are too many unknowns related to them at the moment such as a varying spectral index and that their population in the Universe has not been estimated to a satisfactory degree yet.

An important aspect of foregrounds is the amount of polarization they carry, since it collectively hinders the determination of the magnitude/amplitude of the total CMB polarization. We briefly discuss this aspect for each foreground individually in the ensuing sections.

In general, it is agreed that foreground removal is a very important process as it allows for more accurate studies which will lead to a higher degree of fidelity in cosmological conclusions.

1.2 Source Spectra - Radiation from Astronomical Objects

We begin by saying that, in general, a source can emit either thermally or nonthermally. In the first case, the radiation emitted is due to any temperature change of the body while the emission in the latter case is due to effects independent of changes in temperature . The free-free emission, mentioned in the introduction, is the most characteristic thermal-type emission while synchrotron is the most usual non-thermal process (Kraus, 1986; Wilson et al., 2009).

The spectrum of a radio source may be determined by its flux density *i.e.* the energy flow over the solid angle of the source (Kraus, 1986; Wilson et al., 2009). A relationship between the flux density, S and wavelength, λ , has been established for that purpose:

$$S = \frac{2k}{\lambda^2} \iint T d\Omega \tag{1.1}$$

using the Rayleigh-Jeans (R–J) approximation which applies in the low frequency limit, $h\nu \ll kT$. In Equation 1.1, S is the flux density, T is the temperature of the blackbody and $d\Omega$ is the solid angle element. Note that S is equal to the measured flux density only when the solid angle of the source, Ω_s , is small relative to the beam.

Equivalently, we can say (Goldsmith, 2002):

$$S(\theta,\phi) = \int I_{\nu}(\theta,\phi) d\Omega \qquad (1.2)$$

where 1.2, I_{ν} is the brightness of the source in the direction (θ, ϕ) . The rest of the symbols have their usual meaning.

Assuming a uniform temperature distribution over Ω_s , Equation 1.1 reduces to:

$$S = \frac{2k}{\lambda^2} T\Omega_s \tag{1.3}$$

Conway et al. (1963) have reached a more general power law between the flux density and wavelength:

$$S \propto \lambda^{-\alpha} \propto \nu^{\alpha}$$
 (1.4)

where α is the spectral index of the source.

It is obvious that we can re-arrange Equation 1.3 to the form $T \propto \nu^{\beta}$ where the relation between β and α is $\alpha = \beta + 2$ in the R-J limit. The actual definition of the spectral index is given in Kraus (1986). For two measurements at frequencies ν_1, ν_2 we have :

$$\alpha = \frac{\log \left(S_{\nu_1} / S_{\nu_2} \right)}{\log \left(\nu_2 / \nu_1 \right)} \tag{1.5}$$

The spectral index is a dimensionless quantity that can be used to distinguish between the different types of emission (as mentioned in Section 1.1) and hence provide information on the properties of the emitting source.

Lastly, for a thermal source (*i.e.* a source in thermodynamic equilibrium) where the electron temperature, T_e , is a constant and not a function of the optical depth, τ_{ν} , the radiative transfer equation can be solved to the following form (Kraus, 1986):

$$T_b = T_{b,0} \ e^{-\tau_{\nu}} + \ T_e \ (1 - e^{-\tau_{\nu}}) \tag{1.6}$$

where T_b is the brightness temperature and $T_{b,0}$ is the initial brightness temperature.

For an optically thick component, where $\tau_{\nu} \gg 1$:

$$T_b \approx T_e \tag{1.7}$$

Conversely, for an optically thin component, where $\tau_{\nu} \ll 1$:

$$T_b \approx \tau_{\nu} T_e \tag{1.8}$$

1.3 Diffuse Foregrounds

1.3.1 Free-Free Emission

The most important of the thermal emission mechanisms at radio frequencies is free-free emission or bremsstrahlung. Such emission is a result of the acceleration of electrons, generally, but more specifically it arises due to interactions (collisions) of electrons with ions. When an electron moves past an ion at a sufficiently close distance, the ion's field will deflect the electron thus causing a change in its direction. Consequently, the electron will gain an acceleration which will result in the emission of radiation (a detailed proof is contained in Rybicki & Lightman (1986)). It is worth noting that the acceleration gained can be approximated by a delta function in the spectrum of emission since the whole process occurs in a very short amount of time.

This emission derives its name from the fact that the electron was, and will remain, free (i.e unbound) even after the interaction. This is contrary to other processes where a free electron will result in a bound state (free-bound emission) or where the emission occurs for electron transits between bound states (bound-bound emission) (Smoot, 1998).

In radio astronomy we have to extend this mechanism to account for multielectron and ion distributions like those found in ionized gas clouds and plasmas. Here we will only state the most important results without proofs.

The general expression of the energy radiated in free-free emission is given by



Figure 1.2: Variations of the free-free spectral index as a function of frequency. The graph is based on Equation 1.11. The different curves are for increasing electron temperatures equal to 2, 4, 6, 8 (*heavy*), 10, 12, 14, 16, 18, 20×10^3 K(Dickinson et al., 2003).

(Rybicki & Lightman, 1986):

$$\varepsilon_{\nu} = \frac{dW}{dtdV} = 1.4 \times 10^{-27} \ T^{1/2} \ n_e n_i Z^2 \ \overline{g_B}$$
(1.9)

where $\overline{g_B}$ is a function of temperature T and is the frequency-averaged Gaunt factor for free-free emission, n_i is the ion density, n_e is the electron density and the rest of the symbols have their usual meaning. The Gaunt factor works as a correction to the classical results due to quantum mechanical effects. The above equation is stated in units of $[erg \ cm^{-3} \ s^{-1}]$.

A more refined version of this equation can be found in several papers and books (Dickinson et al., 2003; Oster, 1961; Rybicki & Lightman, 1986) and is quoted below:

$$\varepsilon_{\nu}^{ff} = 6.82 \times 10^{-38} \ T_e^{-1/2} \ Z^2 n_e n_i \ e^{-\frac{h\nu}{KT}} \ \langle \overline{g_{ff}} \rangle \tag{1.10}$$

in units of $[erg \ cm^{-3} \ s^{-1}Hz^{-1}]$, for T_e below 550,000 K where $\langle \overline{g_{ff}} \rangle$ is the velocityaveraged Gaunt factor.

The optical depth of free-free radiation is:

$$\tau_{\nu} \approx 8.235 \times 10^{-2} T_e^{-1.35} \nu^{-2.1} \int n_e^2 dl \qquad (1.11)$$

where T_e is the electron temperature and the integral denotes the Emission Measure (EM), a parameter associated with plasmas, with units $[cm^{-6} pc]$.

Free-free emission is the major contributor in foreground contamination at the frequency range of 10-100 GHz (Dickinson et al., 2003) near the Galactic plane. At the same time, its impact at higher Galactic latitudes is weaker when compared to other foregrounds, such as synchrotron emission (Davies et al., 2006).

The spectral index of this emission is fairly constant under variations of the electron temperature and frequency (Bennett et al., 1992; Dickinson et al., 2003) and is well known. Its value is $\beta \approx -2.1$ with the associated error ranging from 0.03 to 0.1. We can conclude then that the spectrum of free-free emission is expected to be relatively flat as shown in **Figure 1.2**.

A general expression of the free-free emission spectral index is given in Bennett et al. (2003):

$$\beta_{ff} = 2 + \left(\frac{1}{10.48 + 1.5 \ln \frac{T_e}{800 \ K} - \ln \nu}\right) \tag{1.12}$$

where the frequency, ν , is in GHz units (Burke & Graham-Smith, 2002).

$\mathbf{H}\alpha$ and Free-Free emission

The H α line is the spectral line created due to transition of the electron from the n=3 to the n=2 state in atomic hydrogen. The H α emission occurs mostly from the Interstellar Medium (ISM). The surface brightness of the H α emission is given by (Leitch et al., 1997):

$$I_{H_{\alpha}} = 0.36(EM) \ T_4^{-0.9} \tag{1.13}$$

where T_4 is the electron temperature in units of 10^4 K. Here $I_{H_{\alpha}}$ is given in units of Rayleigh, R, a unit of light intensity $(1 R = \pi/4 \times 10^{10} photons \times s^{-1} m^{-2} sr^{-1})$ (Baker, 1974). Dickinson et al. (2003) use a different, more useful expression:

$$I_{H_{\alpha}} = 9.41 \times 10^{-8} T_4^{-(1.017)} \times 10^{-0.029/T_4} (EM)_{cm^{-6}pc}$$
(1.14)

in units of $[erg \ cm^{-2} \ s^{-1} \ sr^{-1}]$.

Now, we see that both Equation 1.13 and Equation 1.14 are proportional to the EM. For this reason, along with the fact that free-free emission also originates from the ISM, it is generally accepted that a relation between free-free and H α emission can be established. The two aforementioned equations can be used to derive that relation; (Smoot, 1998):

$$T_b^{ff} = 1.68\mu K \langle g_{ff} \rangle \left(\frac{T}{10^4 K}\right)^{0.4 \ to \ 0.7} \left(\frac{\lambda}{1 \ cm}\right)^2 \left(\frac{I_{H_{\alpha}}}{R}\right)$$
(1.15)

or,

$$\approx 7 \ \mu K \left(\frac{T}{10^4 \ K}\right)^{0.55 \ to \ 0.85} \left(\frac{\lambda}{1 \ cm}\right)^{2.1} \left(\frac{I_{H_{\alpha}}}{R}\right). \tag{1.16}$$

Still, the connection between H α and free-free emission goes beyond Equation 1.15 or 1.16. The H α line is considered a good "tracer" of free-free emission i.e it can provide us with information concerning the regions where the latter is most likely to be found (Dickinson et al., 2003; Smoot, 1998). This information can then be used to create a free-free template (**Figure 1.3**). This is important since free-free is hard to detect or "separate" from synchrotron radiation, especially at high latitudes, as previously mentioned.

As for all the other foregrounds, observations over a wide frequency range are considered one of the best approaches to separate them from free-free emission. A second approach involves the use of radio recombination lines from which one can derive an appropriate value for the electron temperature of the region under observation, T_e .

Bennett et al. (2003) includes the lack of certainty in the behavior of the electron temperature as a severe limitation in mapping H α and hence in creating a reliable template for free-free. Other limitations include the geocoronal H α emission of the



Figure 1.3: Free-free brightness temperature template at 30 GHz at 1 °resolution as derived by Dickinson et al. (2003). Regions where the template cannot account for with certainty are white. The dot-dashed lines are contours (Dickinson et al., 2003).

Earth that may be confused for Galactic emission of the same kind and inaccuracies associated with the Balmer atomic rates.

Lastly, the free-free polarization contribution is generally considered to be about 10% at maximum, arising due to effects of Thompson scattering (Davies & Wilkinson, 1999).

1.3.2 Synchrotron Emission

Another very important foreground, synchrotron emission, is quite similar to freefree (it was even called magneto-bremsstrahlung initially) in that ultimately the origin of the emission is the acceleration of the electron. Unlike to free-free emission though, the electron will accelerate inside a magnetic field, or more correctly, it will accelerate due to the change in its direction as it moves in circular loops (about the magnetic field lines) induced by the magnetic field force. An important point that needs to be clarified is that the electron is traveling inside the magnetic field with relativistic velocities. For a non-relativistic electron, the same process will lead to another type of radiation called cyclotron emission. A mathematical derivation of the synchrotron emission formula is given in many textbooks; a rather simple and straightforward one is contained in Wilson et al. (2009). Here we will just present the most important results. The electrons will move inside the magnetic field with a frequency,

$$\omega_B = \frac{eB}{\gamma m} \tag{1.17}$$

where we can also define the gyrofrequency (or cyclotron frequency)

$$\omega_G = \frac{eB}{m} \tag{1.18}$$

such that

$$\omega_B = \omega_G / \gamma \tag{1.19}$$

In the above equations, B denotes the magnetic field strength, γ is the Lorentz factor and the rest of the symbols have their usual meaning.

It will radiate with an intensity:

$$I(\nu) = \frac{\sqrt{3}e^3}{8\pi mc^2} \left(\frac{3e}{4\pi m^3 c^5}\right)^{(p-1)/2} LN_0 B_{eff}^{(p+1)/2} \nu^{-(p-1)/2} a(p)$$
(1.20)

where N_0 is the relativistic electron density, L is the emission depth, B is the magnetic field strength, ν is the frequency and a(p) is just a function of the electron energy spectrum, p. The rest of the symbols have their usual meaning. The above equation is given in SI units.

The dependence of the intensity from the magnetic field strength is to be expected as for stronger fields, the magnetic field lines will be narrower and the rate which the electrons spiral around them, *i.e.* their angular acceleration, will be larger. Note that we assume that the electrons are moving randomly in space and also that their energy spectrum can be related to the electron density by :

$$\frac{dN}{dE} = N_0 \ e^{-p} \tag{1.21}$$

The total power radiated will be (Rybicki & Lightman, 1986):

$$P = \frac{4}{3} \sigma_T c \beta^2 \gamma^2 U_B \tag{1.22}$$

where $\sigma_T = \frac{8\pi r_0^2}{3}$ is Thompson's cross section, β is the electron's velocity at a specific point in time and space relative to the speed of light (*i.e.* v/c) and $U_B = \frac{B^2}{8\pi}$ is the magnetic energy density.

Finally, the synchrotron spectrum will die out after a critical or "cut-off" frequency ω_c has been reached, equal to (Rybicki & Lightman, 1986):

$$\omega_c = \frac{3}{2}\gamma^2 \omega_B \sin \alpha \tag{1.23}$$

where α is the electron direction pitch angle to the magnetic field and the rest of the symbols have their usual meaning.

The intensity of the spectrum (e.g. Miville-Deschênes et al. (2008)) is given by a power law where, based on the dependence of the synchrotron emission on the magnetic field and the electron density, we have:

$$S(\nu) = S(\nu_0) \left(\frac{\nu}{\nu_0}\right)^{\beta} \tag{1.24}$$

The origins of synchrotron radiation are well understood. Bennett et al. (2003) states that an overwhelming 90% is emitted as a mechanism of energy loss by cosmicray particles, as a result of their encounters with B-fields possibly during their escape. The remaining percentage is found mainly near supernova remnants (SNRs) of Type Ib or Type II, or other high energy environments. This is logical since all energy loss of the electron while spiraling the magnetic field needs to be replenished in order for the electron to continue to emit.

The spectral index of synchrotron can be most easily measured at low Galactic latitudes where the synchrotron spectrum is easily distinguishable from other emission spectra like the free-free one. Davies & Wilkinson (1998) state that it has a mean value of -2.7 however it presents fluctuations of typically 0.3 units in the frequency range of 38-1420 MHz. According to Lawson et al. (1987) typical values of the spectral index in a similar frequency range (38-800 MHz) are $\beta = -2.55$ and $\beta = -2.8$. For the range 408-1420 MHz, Reich & Reich (1988) report $\beta = -2.3$ to -3.0 with an associated error of about ± 0.15 . These values agree with the conclusion expressed by Bennett et al. (2003) that 90% of the observed synchrotron radiation has a spectral index that lies between -2.5 and -3.1. Note that β here denotes the temperature spectral index (just like in Section 1.2) for which Davies & Wilkinson (1998) provide us with an expression: $\beta = 2 + \frac{p-1}{2}$.

In general, the value of the spectral index provides a wealth of information for the nature of the source of synchrotron emission. For example a steep spectral index can mean that the CR electron loses large amounts of energy very fast and thus will be unable to travel very far. Moreover the age of the source can inferred from a "break" in the emission spectrum (Bennett et al., 2003). This synchrotron break has been studied and observed by a number of authors such as Voelk (1989) and Green & Scheuer (1992). It is widely accepted that the synchrotron spectral index will become steeper over time, meaning that the amount of synchrotron emission will decrease, as the low energy CR electrons do not interact as strongly as the high-energy ones and hence survive longer.

It is worth noting that the most widely used tool in creating templates for the synchrotron emission is the 408 MHz Haslam all-sky survey (Haslam et al., 1982). This is presented in **Figure 1.4**.

Lastly, polarization from synchrotron emission arises by definition; the emission coming from a charge that moves inside a magnetic field has been found to be elliptically polarized (Rybicki & Lightman, 1986). The contribution from multiple electrons, however, comes out as partial linear polarization (rising up to 60-75 % depending on the spectral index). The synchrotron foreground is the most polarised for low and intermediate radio frequencies *i.e.* \leq 50 GHz and therefore it is most significant when determining the overall CMB polarisation signal.



Figure 1.4: The Haslam 408 MHz map used as a template for synchrotron emission (Haslam et al., 1982).

1.3.3 Thermal Dust

In principle, there are several emission mechanisms that originate from a dust grain depending on its nature; for instance the radiation from a rotational grain is referred to as "rotational" emission. If an, initially, stationary grain starts vibrating (usually caused by heating from starlight) then the radiation emitted is called "vibrational emission" (most commonly revered to as "thermal dust emission"). In the following sections we will also look in more detail other dust emission mechanisms namely due to spinning and magnetized grains (Section 1.3.5 and 1.3.6 respectively).

It is highly suspected that the material of the dust grains that accounts for this type of emission involves silicate trailite and foyalite (Finkbeiner & Schlegel (1999)(hereafter FS99). The general practice as far as the size distribution of the grains is concerned (for all types of dust emission) is that it will follow the relation (FS99; Draine & Lazarian 1998; Draine & Lee 1984):

$$\frac{dn}{da} \propto a^{-3.5} \tag{1.25}$$

In general, thermal dust will radiate mostly in the FIR spectrum and as mentioned previously it will dominate at higher frequencies close to and above 100 GHz. The intensity of emission will be given by:

$$I_{\nu} \propto B_{\nu}(T) \cdot \nu^2 \tag{1.26}$$

where a ν^2 emissivity was used.

FS99 and Finkbeiner et al. (1999)(hereafter FDS99) comment that even though the above law can be theoretically supported as the low-frequency dependence of a forced oscillator it still will not hold for "bigger" grains whose size will impact heavily on the amplitude of emission. Values for the emissivity index vary from 1 to 3; FDS99 proposes a ν law while FS99 states that a $\nu^{1.5}$ dependence describes more accurately silicate grains. The standard law is ν^2 .

For a multi-component model (where component denotes the material that accounts for the dust grains), derived using the DIRBE-calibrated 100 μm map, FDS99 give:

$$I_{p,\nu} = \frac{\sum_{k} f_k \epsilon_k(\nu) B_\nu(T_{pk})}{\sum_{k} f_k \epsilon_k(\nu_0) B_{\nu_0}(T_{pk}) K_{100}(a_k, T_{pk})} I_{100}$$
(1.27)

where f_k is a normalisation factor for the kth component, T_{pk} is the temperature in pixel p of the map of the component k, K_{100} is the DIRBE color-correction factor, $I_{p,100}$ is the 100 μm flux at pixel p and $\sum_k f_k$ is equal to unity. We should point out that the proposed model by FDS99 involves only 2 components; silicate and carbonite.

The spectral index of the emission has been found to vary with an upper and lower limit of -1.5 and -2 respectively. It can be used to identify the material the dust grains are made of, as shown by Dupac et al. (2001). In their research they agreed and extended the FS99 results (Bennett et al., 2003). The close relation between thermal dust and synchrotron emission has been established somewhat indirectly, though it is considered a robust conclusion, nonetheless (Bennett et al.,



Figure 1.5: FDS thermal dust map at 94 GHz based on data from IRAS and COBE-DIRBE (Finkbeiner et al., 1999).

2003). More specifically, since thermal dust emits strongly in the radio/ FIR spectrum and synchrotron and FIR emission correlate, it follows that both of these foreground components should also show a degree of correlation between them.

The template used most frequently to map the thermal dust emission was created by Schlegel et al. (1998). The derived map can be seen in **Figure 1.5**.

Last, but not least, the polarized thermal dust component only becomes significant at higher frequencies than the ones dominated by synchrotron radiation. These are, typically, equal or greater than 100 GHz. The levels of polarization vary from 5 - 10 %. It is worth noting that, although the general mechanism behind polarised dust emission has been accepted as grain alligment by magnetic forces (Tegmark et al., 2000; Hiltner, 1949; Hall & Mikesell, 1949) the specifics have not yet been determined.

1.3.4 Anomalous Dust Emission

The existence of an anomalous component of dust emission has been well established by a large number of authors and experiments such as Kogut et al. (1996), Leitch et al. (1997), de Oliveira-Costa et al. (1998) and Finkbeiner et al. (2004). This emission is highly correlated with the 100 μ m (*i.e.* far-IR) thermal emission from interstellar dust, as reported by Leitch et al. (1997), and emits in the range ~ 10-60 GHz.

Over the course of years there have been multiple detections (in HII regions) mostly although some of them are characterized too "tentative" while others are just "statistical". Thus this emission has been termed Foreground X (de Oliveira-Costa et al., 2004) due to its mysterious nature.

First detection dates in 1996 when Kogut et al. noticed a rise in the amplitude of the signal, contrary to their theoretical predictions, at 31.5 and 53 GHz in the Differential Microwave Radiometer (DMR) maps. Based on the spectral index derived, they explained this anomalous rise as a possible mixture of free-free and dust emission but certainly not a new component of dust emission only. At the same time, they overruled synchrotron emission as a possibility based on comparisons with the 408 MHz survey. A similar detection was later recorded by Leitch et al. (1997) using the Owens Valley Radio Observatory (OVRO). This team saw rises at 14.5 and 32 GHz which they explained via a very hot free-free emission model (mainly due to the value of the spectral index they derived) as well as strong correlation with dust emission.

Other projects include the Tenerife survey where the images produced were of low resolution (Davies et al., 1996) and the 19 GHz survey (Boughn et al., 1990).

In 1998, Draine & Lazarian published two investigations (Draine & Lazarian, 1998, 1999) of emission from spinning dust grains and magnetic dipole emission from magnetic grains due to thermal fluctuations (both of these are discussed in Sections 1.3.5 and 1.3.6 respectively) and suggest that these mechanisms can explain the Foreground X emission. In the meantime, they disagree with the model proposed by Leitch et al. (1997) based on arguments concerning the energy required for such a hot free-free emission to be present. Dickinson (2008) comments that the reason anomalous emission is commonly found in HII regions may be because of the high amounts of spinning dust emission by the very dense clouds found there. It is plausible, therefore, that a significant fraction of Foreground X could be due to spinning dust grains..

Since then, there have been several detections of spinning dust coupled with an anomalous component such as the dark cloud LDN1622 (Finkbeiner et al., 2002) the region inside the Perseus Molecular Cloud G.159.6-18.5 (Watson et al., 2005) and the supernova remnant 3C396 with the Very Small Array (VSA) at 33 GHz (Scaife et al., 2007) which is the only detection in this type of region as of August 2010. Most recently, the VSA was again able to detect evidence of anomalous dust emission in nine HII regions (Todorović et al., 2010) while another detection was made at 33 GHz in the RCW175 HII region but by the CBI, this time (Dickinson et al., 2009). Other notable observations include the LDN1111 cloud with the Arc Minute Imager (AMI) at 15 GHz (Scaife et al., 2009) and four Lynds clouds (L675, L944, L1103, L1246; Ami Consortium et al. (2009)).

The detection of anomalous emission coming from the LDN1622 cloud was part of a larger project in which Finkbeiner et al. (2002) used the Green Bank 140ft telescope to observe ten sources of Galactic emission. In their conclusions the research team names another Galactic object, the LPH[96]201.663+1.643 cloud, as a possible anomalous dust agent however subsequent observations with the CBI at 31 GHz by Dickinson (2006) proved that its spectrum was more compatible with free-free emission. On the other hand, the first extragalactic detections of spinning dust by Murphy et al. (2010) from the NGC6946 star forming region were confirmed by Scaife et al. (2010).

The main reason that the existence of Foreground X is still debated is due to the conclusions of the WMAP team in 2003 (Bennett et al., 2003) where they suggest a model that can account for the Galactic foreground contamination without the addition of the anomalous component while instead making use of a flat-spectrum synchrotron (often called "hard") component. Astonishingly, the effectiveness of
such a model (or of the template derived by it) reaches $\sim 98\%$ indicating that indeed the concept of the anomalous foreground is not needed. The frequency range of the observations was 23-94 GHz.

Despite this, further controversy arose later in the same year and the next when a series of experiments by different groups (Lagache, 2003; Banday et al., 2003; Casassus et al., 2004; de Oliveira-Costa et al., 2004) recorded some excess emission in HII regions and mostly in the frequency range of ~ 30 GHz. Perhaps the most important of these research results came from Finkbeiner et al. (2004) when using Green Bank Galactic Plane Survey (GBGPS) data managed to produce images of the anomalous emission from ~ 8 to ~ 14 GHz which were, in fact, in accordance with the spinning dust models described in two papers by Draine & Lazarian (1998a and 1998b) while announcing the spinning dust emission as the dominant foreground below 23 GHz. These spinning dust models, each one based on a different environment, are shown in **Figure 1.6** where we see that the emissivity of spinning dust does not vary significantly between environments.

At the same time, though, none of these were able to overrule the WMAP model. This is because, as noted by de Oliveira-Costa et al. (2004), the spectral signature of some spinning dust models is similar to the hard synchrotron for frequencies larger than 23 GHz; essentially the WMAP observational frequency range. Obviously the images produced by the GBGPS survey cannot account for frequencies above 23 GHz. It should be noted that the WMAP team did not completely rule out the existence of Foreground X; rather they state that the model, and the subsequent map they have produced, characterize in a better way regions of high frequency and high galactic latitudes (Bennett et al., 2003).

In general, it is safe to say that the detection of the anomalous dust component can be a very challenging task. It is suggested that the reason behind this lies in the emission properties of Foreground X (de Oliveira-Costa et al., 2004). Particularly, if the intensity peaks at 10-15 GHz (as was found by Finkbeiner et al. (2004)) then any observations outside this regime will either be overshadowed by the synchrotron emission (on the lower end) or confused with free-free emission (on the higher end)



Figure 1.6: Model dust emissivities of the DL98b in Finkbeiner et al. (2004) where the results from other experiments have been included. Solid lines are total emissivity while dashed lines are rotational emission. The curves peak between frequencies of 20-40 GHz.

due to similarities in the spectrum pattern. Evidently, the best approach in studying the anomalous foreground even further will be more observations in that range.

The spectral index of anomalous dust has been estimated by Leitch et al. (1997) to be $\beta \sim -2$ at 15-30 GHz, with an upper limit of -1.9 and a lower limit of -2.2 with a confidence level of 68%. The first year WMAP results (Bennett et al., 2003) take the spectral index of spinning dust to be around the same value for a frequency domain of 20-40 GHz. The similarity between the spectral indices of synchrotron and anomalous emission explains the high degree of correlation between them that we mentioned above. On the other hand, Davies et al. (2006) in their analysis of WMAP data have found a slightly higher average spectral index of ~ -2.85 between 20 and 60 GHz with a maximum value of -3.8. We conclude that the average spectral index is a strong function of frequency, a conclusion also supported by theoretical models of spinning dust; see **Figure 1.6** (Finkbeiner et al., 2004).

The actual polarization mechanisms for this emission are still debated. A proposed process, under the assumption that the anomalous dust emission is due to spinning dust grains, is similar to the production of polarized emission by thermal dust *i.e.* grain alignment under a magnetic field. Furthermore the contribution of this component to the total CMB polarization is assumed to be 2-3 % at the frequency range where anomalous emission is observed more strongly (Draine & Lazarian, 1998).

1.3.5 Electric Dipole Radiation from Spinning Dust Grains

Damping Mechanisms

In explaining the properties and physics of spinning dust grains we will follow closely the model presented by Draine & Lazarian (1998) (hereafter DL98b) while adopting most of their assumptions. Rotational excitation of the grain depends on the grain size and the environment in which the grains are located. More specifically, spinning grains will be affected by ions in several ways but mostly by transfer of angular momentum and collisional excitation while plasma oscillations can also cause excitation (or damping). DL98b consider the dynamics of the grains as if they were spherical; their actual geometry is, as of now, uncertain. In this section we will consider the properties of the emission only from "very small" grains (atom population of $10^2 - 10^3$), since the frequency range in which the anomalous emission is observed (10-100 GHz roughly) is dominated by such grains.

The power radiated by a spinning dust grain is given by:

$$P = \frac{2}{3} \, \frac{\omega^4 \mu^2 \sin \theta}{c^3} \tag{1.28}$$

where μ is the electric dipole moment and θ is the angle between the angular velocity ω and μ . The above can be simplified if we assume that μ is directionally independent of ω ; the average value of $\sin \theta$ will be equal to $\frac{2}{3}$ and thus:

$$P = \frac{2}{3} \times \frac{2}{3} \frac{\omega^4 \mu^2}{c^3} = \frac{4}{3} \frac{\omega^4 \mu^2}{c^3}$$
(1.29)

Consider the different types of damping that a rotational grain might undergo

depending on the environment it is located in. These could be: gas-grain interactions, plasma-grain interactions, IR and radio emission. If we define a dimensionless quantity F that describes the total drag contribution to the total damping then we can write :

$$F = F_n + F_i + F_p + F_I R \tag{1.30}$$

where $F_n, F_i, F_p, F_I R$ are the contributions related to each damping process stated above.

Generally the contribution to a drag process, F_j , to the total drag torque is defined as:

$$I\left(\frac{d\omega}{dt}\right)_{j} \equiv -\left[n_{H}\left(\frac{8kT}{\pi m_{H}}\right)^{1/2} \frac{2\pi \alpha_{x}^{4} m_{H}}{3}\right] \omega F_{j}$$
(1.31)

where n_H is the nucleon density, m_H is the mass of the hydrogen atom, α_x is the rms distance from the grain's surface to the center of mass and the rest of the symbols have their usual meaning.

Collisional Drag due to Gas-Grain Interactions In collisional damping, a particle (*e.g.* a gas molecule) will collide with a grain plastically thus decreasing the grain's rotational period by providing additional angular momentum. Eventually, the particle will be expelled and upon expulsion it will take an amount of the surface's angular momentum thus decreasing its overall rotational velocity. In fact, the angular momentum of the particle will be proportional to the grain's rotational velocity.

Plasma Drag Alignment of dust grains depends primarily upon two factors: the magnetic field force and the amount of collisions with moving particles comprising the environment where the grain resides. In general, the former is stronger than the latter and hence the rotational axis of a spinning grain will align with the magnetic field lines. Moreover, if the grain is inside the plasma, such as an HII region, then it will be affected by ions and electrons that go past it thus increasing the chance of possible de-alignment. This effect however is not very significant for dipole moments (which we are interested in) while even for higher order moments (e.g. quadrupole) it is negligible.

IR, Radio Emission & Electric Dipole Damping This process involves the absorption of a UV photon by the grain and subsequently the emission of a number of IR photons such that the number of emitted photons is significantly larger than the number of absorbed photons. Of course, any absorption/emission means that the photons will give/take away angular momentum from the grain. Since the number emitted is larger, the total angular momentum will decrease after the IR photon emission thus causing rotational damping.

Similar procedure holds for radio emission with the only difference being in the emitted particle as, in this case, it is a radio photon while in electric dipole damping the grain emits electric dipole radiation which damps the rotation in a similar manner as the one described above.

Summing up, all mechanisms damp the rotational motion of the spinning dust grains but Davies & Wilkinson (1999) note that, for grains of our assumed size, plasma drag is the major contributor.

Rotational excitation

The basic principle here is that the grain's angular momentum changes are due to impacting particles and due to the recoil from subsequent evaporation (emission) of those particles or others.

In general under this category we have four different ways of excitation based on the particles responsible. As such we have: excitation from collisions with neutrons and ions, excitation from the plasma (where the ions and electrons that pass by set up an electric field thus disturbing the grain's rotation), excitation from IR emission (where the grain is excited by the recoil from the emission of IR photons - note that this emission also causes damping as we explained previously) and excitation from the recoil due to emission of photoelectrons. The rate of increase of the rotational kinetic energy of a stationary grain is given by:

$$\frac{d}{dt}\left(\frac{1}{2}I\omega^2\right) = n_H\left(\frac{8kT}{\pi m_H}\right)^{\frac{1}{2}} \left(\frac{2\pi\alpha_x^4m_HkT}{I}\right) \left(G_n + G_I + G_p + G_IR\right)$$
(1.32)

where G is the normalized excitation rate of each mechanism and the rest of the symbols have their usual meaning.

 H_2 formation Incoming H atoms may recombine on the surface of the grain to form H_2 atoms and affect its rotation due to their additional translational kinetic energy. This effect is negligible and can thus be ignored.

To sum up, the dominant excitation mechanism is through collisions with ions.

Emissivity

The emissivity per H atom due to dust grains is given by:

$$\frac{j_{\nu}}{n_H} = \left(\frac{8}{3\pi}\right)^{\frac{1}{2}} \times \frac{1}{n_H c^3} \int da \frac{dn}{da} \frac{\mu^2 \omega^6}{\langle \omega^2 \rangle^{\frac{3}{2}}} \exp\left(\frac{-3\omega^2}{2\langle \omega^2 \rangle}\right)$$
(1.33)

where $\langle \omega^2 \rangle$ is the mean square angular velocity and ω follows a Boltzmann distribution, a is the grain size, n_H is the H molecule nucleon density, μ is the grain dipole moment which is proportional to the electric dipole moment, $\frac{du}{da}$ is the size distribution of grain species.

The expected value of emissivity is

$$\frac{j_H}{n_H} = 1.1 \times 10^{-15} \times \left(\frac{\nu}{100 \ GHz}\right)^{2.8} \ Jy \ sr^{-1} \tag{1.34}$$

This value is found by both DL98b and Ferrara & Detmar(1994) however their method and assumptions are very different and often contradicting. DL98b comment that this value is indeed large and therefore it should make quite an impact in observations in the frequency range of 10-100 GHz. It is worth noting that for frequencies above 70 GHz, thermal emission from dust will overpower the rotational emission while this picture is reversed for lower frequencies. Especially in the frequency range of 30-50 GHz, as pointed out by DL98, the results reached by a number of authors (Leitch et al., 1997; Kogut et al., 1996; de Oliveira-Costa et al., 1998) can be sufficiently explained by considering rotational emission as the majority contributor of the anomalous component.

1.3.6 Magnetic Dipole Emission from Dust Grains

This is the second possible model that could account for the anomalous emission as proposed by DL98a and described in Draine & Lazarian (1999)(hereafter DL99). We will follow closely their model.

The key idea here is that magnetic dipole radiation may arise due to "thermal fluctuations of the magnetization of the dust grain". In particular, when a change in the grain's temperature will occur (due to the heat produced by the magnetization) it will cause a change in the internal energy of the system (for a constant heat capacity). A variation of this kind is termed a thermal fluctuation. It is suspected that this kind of emission accounts for part or all of the anomalous emission observed at microwave/radio frequencies. Of course the magnitude of the fluctuations (and hence of the radiation emitted) depends on the grain material.

As noted by DL99, a 30% of the grain mass could be Fe and Ni which are highly magnetic materials thus making this model quite plausible, even more so when considering the abundance of grains made from Fe.

The general formula of the fraction of the total emissivity due to magnetic dipole emission is:

$$\frac{j_{\nu}}{n_H} = \frac{n_{gr}}{n_H} \langle c_{abs}{}^{md} \rangle B_{\nu}(T)$$
(1.35)

where $\langle c_{abs}{}^{md} \rangle$ is the angle-averaged magnetic dipole absorption cross section, n_{gr} is the grain number density, n_H is the grain number density and $B_{\nu}(T)$ is the Planck function. All types of interstellar dust grains that have been considered as possible candidates for this emission involve Fe in their structure. DL99 conclude that the only way for emission due to thermal fluctuations in the magnetization of grains to account for the anomalous emission observed would be if the emitter were in the form of a hypothetical material with certain properties that they call "X4". This material, which is assumed to consist of Fe in a high degree, has not been discovered yet.

The polarization of this component has been found to be quite large in the 14-90 GHz range and also highly frequency dependent. Like the rest of the dust components it arises due to grain alignment by the magnetic field (DL99).

1.4 Synthesis Imaging

In this section we will review the basic principles of synthesis imaging. The main aim of this field of interferometry is to produce the image of a stellar object by combining the different signals detected by an interferometer. Some basic concepts of interferometry will also be reviewed.

The correlation between two signals often means the "product" of these or their combination and it often hints information on their relationship. We can find the correlation properties of a field by simply measuring the spatial coherence function (or, visibility function) which is defined as (Clark, 1999):

$$V_{\nu}(\boldsymbol{r_1}, \boldsymbol{r_2}) \approx \int I_{\nu}(\boldsymbol{s}) e^{-2\pi i \nu \boldsymbol{s} \cdot (\boldsymbol{r_1} - \boldsymbol{r_2}/c)} d\Omega \qquad (1.36)$$

where r_1, r_2 are the position vectors, $I_{\nu}(s)$ is the observed intensity, s is the unit vector and the rest of the symbols have their usual meaning.

By defining the spatial coherence function (SCF) we have succeeded in defining the interferometer as its primary purpose is to measure this function. Measurement of the SCF, and consequently synthesis of the image, usually occurs at the "u-v" plane where v points along the N-S direction as shown in **Figure 1.7**. In practice, the u,v plane is defined as $[r_1 - r_2]/\lambda(u, v, w)$ in 3-D synthesis imaging. Here λ is



Figure 1.7: Tracking of the source in the u-v plane as the Earth rotates. Taken from: http://gmrt.ncra.tifr.res.in

the wavelength. It is also convenient to define an "l-m" plane, which can give the coordinates on the sky, such that: $\mathbf{s} = (l, m, \sqrt{1 - l^2 - m^2})$ again in 3-D synthesis. We can rewrite the SCF in this coordinate system as:

$$V_{\nu}(u,v) = \iint A_{\nu}(l,m) I_{\nu}(l,m) e^{-2\pi i (ul+vm)} dl \ dm$$
(1.37)

where $A_{\nu}(l,m)$ is the "primary beam or normalised reception pattern of the interferometer" and the rest of the symbols are as defined previously. The quantity $V_{\nu}(u,v)$ is often called the "complex visibility" (Taylor et al., 1999). Its units are $[Wm^{-2}Hz^{-1}]$. The above relation clearly shows that the Fourier Transform (FT) of the SCF is equivalent to the intensity distribution. This is essentially the basis of all the methods used in synthesis imaging.

Another useful function is the "sampling function" defined as (Clark, 1999):

$$I_{\nu}^{D}(l,m) = \iint V_{\nu}(u,v)S(u,v)e^{2\pi i(ul+vm)}du \ dv$$
(1.38)

where $I_{\nu}^{D}(l,m)$ is referred to as the "dirty image" and $\iint S(u,v)e^{2\pi i(ul+vm)}du dv$ is the point-spread function (PSF). In words, the sampling function is the "ensemble of elliptical loci" which are the projection of the baseline vector while tracing out the source on the *u*-*v* plane. It is obvious that the FT of the sampling function is equal to the PSF. Another fundamental statement is that the convolution of the intensity distribution, I_{ν} , with the PSF is equivalent to the dirty image (Clark, 1999), *i.e.*:

$$I^D_{\nu} = V_{\nu} * PSF \tag{1.39}$$

where the (*) means convolution.

We now describe the way an interferometer functions. For this purpose we will consider its most basic form, a 2-element interferometer. The signal from a radio source is picked by the antennae, then forwarded to the amplifiers (where the signals are amplified separately) and the correlator (where their combination occurs). The final signal is then visible in the form of a fringe pattern while the correlator output is given by:

$$r(\tau_g) = v_1 v_2 \cos(2\pi\nu\tau_g) \tag{1.40}$$

where τ_g is the geometrical delay, v_1, v_2 are the amplitudes of the incoming signals and the cosine term accounts for the Earth's rotation that will affect the visibility.

The sensitivity of an interferometer i.e "the time τ it requires to achieve a particular noise level, σ_{ν} , with a single baseline" (O'Neil, 2002) is given by:

$$\sigma_{\nu}(Jy) = \frac{\sqrt{2}k_B T_{sys}}{A_{eff}\eta\sqrt{\Delta\nu \times \tau}}$$
(1.41)

where T_{sys} is the system temperature which is defined as the temperature of all the factors that affect our observations, $\Delta \nu$ is the noise bandwidth, A_{eff} is the effective area of the antenna which determines the percentage of the collected radiation compared to the incident, η is the correlator efficiency, τ is the integration time and the rest of the symbols have their usual meaning.

Now, higher sensitivity will mean less time to reach a particular noise level so if one aims in designing a sensitive interferometer, τ should be kept as low as possible. However, from Equation 1.41 this would mean an increase in the noise levels $\propto \frac{1}{\sqrt{\tau}}$. It follows then, that T_{sys} should also be kept as low as possible to prevent this from happening. Another factor that will affect the sensitivity of the instrument, by reducing the observed flux density (Condon, 2002) and, in particular, the phase of the visibilities (Cartwright, J.K., 2003), is the pointing errors. This reduction is described by the mean observed intensity, $\langle G \rangle$, and it is given by (Condon, 2002):

$$\langle G \rangle = [1 + (8ln2 \times \epsilon^2)]^{-1}$$
 (1.42)

where ϵ is the pointing error in one coordinate. Typically, the rms pointing error $\sigma_{\alpha} \approx \sigma_{\delta}$ in each coordinate is given by $\epsilon\theta$ where θ is the beamwidth. The reduction caused will be of a factor of $\langle G \rangle / G$ where G is the directivity of the antenna. Typical values for ϵ are 0.05 and 0.1.

The length of the baseline affects the sensitivity of measurements. In general, longer baselines provide very high resolution (*e.g.* MERLIN) but this is a "doubleedged sword" since it decreases the surface brightness sensitivity (Padin et al., 2002). Obviously for shorter baselines the opposite will hold. Note, also that the interferometer can be far more effective than a single dish in observing the CMB anisotropies as it is more capable of shielding noise. At the same time it does not require apertures of large dimensions to achieve greater resolutions (Padin et al., 2002).

In imaging a source, the small-field approximation is often used :

$$(\sqrt{1-l^2-m^2}-1)\ \omega \approx -\frac{1}{2}(l^2+m^2) \approx 0$$
 (1.43)

where l, m are the sky coordinates. Equation 1.43 stems from the assumption that |l|, |m| are small and in particular that $|\pi(l^2 + m^2)| \ll 1$. As a matter of fact, applying this condition ensures minimum distortion in the produced image (Taylor et al., 1999).

A very important part of the data analysis (Chapter 2) that needs to reviewed early on is the concept of closure errors. Closure errors reflect how accurate is the description of the complex gain as the product of an antenna-based amplitude and phase, based on the following equation (Fomalont & Perley, 1999):

$$G_i(t) = a_i(t)a_j(t)e^{i(\phi_i(t) - \phi_j(t))}$$
(1.44)

where $a_{i,j}(t)$ is an antenna-based amplitude correction and $\phi_{i,j}(t)$ is the antennabased phase correction. Since the larger the magnitude of the errors, the less valid the above equation is, it is best to reduce them as much as possible. Fomalont & Perley (1999) outline a few problems that may be the reason behind closure errors; for instance, amplitude and phase variations and non-application of the quadrature calibration on the data.

To conclude, we state that the convolution of the true intensity distribution with the PSF is equivalent to the dirty image. This can be seen by considering that the PSF will limit the resolution of the telescope and will cause "deviation" from the true value of the intensity thus producing the image we observe from the Earth (as described by Equation 1.37)

A solution to the above problem is derived using the CLEAN algorithm; essentially a method developed by J.Hogböm (1974) to deconvolve the dirty image where it is assumed that the sky is made of a number of delta functions. Cornwell, Braun & Briggs (1999) in Taylor et al. (1999) summarize how the algorithm works in five steps. We follow closely their account and describe it below. After recording the properties of one bright point on the dirty image, one subtracts from that point a factor equal to: the dirty beam times the intensity of that particular point times γ , where γ is a damping factor called "loop gain". The properties (position and intensity) of the new point are recorded and the procedure is repeated for the rest of the bright points (peaks). By the end, the image will consist only of the "residuals". At the same time, the new peaks are recorded in a model which needs to be convolved with an "idealized CLEAN beam"; usually an elliptical Gaussian. This will produce the final "CLEAN" image. The final step involves the addition of the residuals to the CLEANed image to complete the process. We discuss CLEAN in a more practical manner in Sections 3.1.1 and 3.1.3.

For completeness, we just mention here that the other method which is widely



Figure 1.8: Image of the original Cosmic Background Imager. Taken from: http://www.astro.caltech.edu/~tjp/CBI/project/Programs/cbical/

used for deconvolution is the Maximum Entropy Method (MEM) algorithm (Taylor et al., 1999).

1.5 The Cosmic Background Imager

We have already given a basic overview of how a 2-element interferometer works. We will now review the properties of a more complicated instrument: the Cosmic Background Imager (CBI; **Figure 1.8**) which was used to measure the CMB radiation fluctuations. Our summary is based on the technical paper by Padin et al. (2002).

The CBI is an "interferometer array" *i.e.* an array of 13 antennae each with a diameter of 0.9 m. These are mounted on a rotating platform and they have been set to operate at a frequency range of 26-36 GHz. As explained by Padin et al. (2002), this was not a random selection as outside these limits, foreground contamination and noise contribution to the data becomes very hard to remove. This 10 GHz range is divided into 10 channels of 1 GHz width and that permits measurements of the spectral index. As we have already said, the spectral index is crucial in determining

the nature of foreground contamination and hence the CBI measurements can be used to remove this source of confusion. We ought to point out, though, that point sources and their removal pose a significant problem to the CBI. On the other hand, for the removal of gain and phase variations, a noise calibration source has been implemented to the system; its flux density is calibrated against several celestial and planetary objects. This internal source is discussed in more detail in Section 2.3.2.

The CBI signal processing is quite complicated as it involves multiple components, however, it can be divided in 3 major steps: the "receivers", the "downconverter" and the "correlator". Whilst in the receivers stage, the signal will be detected and amplified while in the correlator stage, the signal will be split between two branches and multiplied with a phase shifter in only one branch. This will create real and imaginary visibilities that will be digitized and saved in the CBI archive. In the intermediate stage, the downconverter will manipulate the signal accordingly so that at the output, it will be at an intensity of 16 dB of the input value; this will ensure that the correlator will work reasonably. A much more detailed depiction of the signal path described above is shown in **Figure 1.9**.

The distance between the antennae varies from a maximum of 5.5 m to a minimum of 1 m. This allows a resolution between 6 and 30 arcmin. Based on observations of Tau A, the primary beam width (FWHM) of the telescope is 28.2 arcmin (at a reference frequency of 31 GHz/ ν for the CBI2; for CBI, the FWHM was equal to 45.2 arcmin) (Dickinson, Clive; private communication).

Due to the rotating platform, the CBI has a very good u, v plane coverage. An additional privilege that the rotation offers is the easy distinction between true and false signal since the latter will rotate with the detector while the former will remain stable.

The location of the CBI is at 5000 metres altitude in the Chilean Andes. The high altitude results in less atmospheric noise in the observations. Significant drawbacks include the harsh weather conditions that have repeatedly hindered the smooth operation of the telescope.



CBI Signal Chain Overview

Figure 1.9: The CBI signal path (Cartwright J.K.,2003). The colourbar at the left indicates the changes in frequency that the signal undergoes as it moves through the components.

Source	T(K)
CMBR	3
atmosphere	1
ground spillover	2
optics	2
HEMT	18
cold downconverter	1

Table 1.1: The temperature of all the CBI noise contributors (Cartwright, J.K., 2003)

We consider the sensitivity of CBI, as was defined in Equation 1.41 in Section 1.4. For the CBI, $A_{eff} = 4860 \ cm^2$ and $\eta = 0.86$ while the bandwidth will remain fixed for the rest of the observations as it depends on the low-noise high electron mobility transistor (or, HEMT) amplifiers and hence it is a property inherent to the system. Thus it becomes evident that the noise level, σ_{ν} , depends only on the time τ and the system temperature, T_{sys} .

In CBI's case, T_{sys} depends on the temperature of the CMB, the atmosphere, the amplifiers and the downconverter while the ground spillover and the antenna optics also contribute to increasing it. In other words:

$$T_{sys} = T_{CMB} + T_{LNA/HEMT} + T_{atmosp}(1 - e^{-\tau}) + T_{spil} + T_{opt} + T_{DC}$$
(1.45)

where T_{CMB} denotes the CMB temperature, $T_{LNA/HEMT}$ is the temperature of the amplifiers, T_{atmosp} is the temperature of the atmosphere, T_{spil} and T_{opt} are the contributions from the ground spillover and the optics respectively, and T_{DC} is the temperature of the downconverter. **Table 1.1** summarizes the CBI's temperature budget, with the amplifiers accounting for more than 65% of the whole system temperature.

We now draw the attention to the term "ground spillover" that was mentioned previously. This is an important consideration, both when making observations as well as during the data reduction and calibration (Chapter 2). In essence, ground spillover describes the power that leaks in to the system from the ground (commonly assumed to be at a temperature of ≈ 300 K) via the sidelobes. Of course not all of the 300 K power gets through but the amount of radiation that does depends highly on the elevation of the CBI. At maximum, when the antennae are facing at the plane parallel to the horizon one expects the spillover to contaminate the signal with an "extra" flux density of a few Jy (Cartwright, J.K, 2003).

Padin et al. (2002) notes that this problem stems from the lack of any shielding from ground emission; the construction of such a shield was not possible due to the limited budget and the heavy weather conditions. Cartwright (2003) reports that even when the telescope is pointing at the zenith, spillover is not completely eradicated indicating that the radiation reflects off the antenna surfaces to the feed.

As we will see in the figures of Chapter 2 ground emission becomes more prominent for shorter baselines where the size of the sidelobes is larger (Cartwright, J.K.,2003).

The CBI was upgraded to CBI-2 in 2006. The main upgrade centred in increasing the antennae diameter from 0.9 meters to 1.4 meters (A.C. Taylor et al., 2009, in prep). It is shown in **Figure 1.10**. Observations ceased in June 2008 in preparation for a new CMB polarization equipment called QUIET (Newburgh & QUIET Collaboration, 2010).

The contribution of the CBI was indeed invaluable in the field of radio interferometry. It had observed and imaged a wide range of stellar objects and areas such as molecular clouds, SNRs and HII regions. Amongst its most important discoveries is that it detected a damping tail and a subsequent excess in the high-latitude CMB anisotropies (Mason et al., 2003; Pearson et al., 2003). Finally, CBI was amongst the first to detect the E-mode CMB polarization (Readhead et al., 2004). These and other results have found many applications in further studies.

1.6 The W40 Cloud Complex

The CBI-2 observational data, we will be focusing on, are related to the emission from the cloud complex called W40. This is a region that has not been extensively



Figure 1.10: The improved version of the Cosmic Background Imager (CBI2). The size of the antennae has increased in an attempt to make the instrument capable of more sensitive observations.

Taken from: http://www.astro.caltech.edu/~tjp/CBI/project/Programs/cbical/

studied at radio frequencies in the past but that has caught the attention of radio astronomers recently as it showed indications of anomalous emission; Finkbeiner et al. (2004) reported an emission excess at 33 GHz as shown in the spectrum of W40 in **Figure 1.12** where it is evident that the spinning dust model fits the data points almost perfectly. An image of W40 in the optical and IR regime is shown in **Figure 1.11** (Rodney & Reipurth, 2008).

In equatorial J2000 coordinates, W40 lies at $18^h \ 31^m \ 29^s$, $-2^{\circ}05'4''$. In Galactic coordinates, this corresponds to 28.77 + 3.70 (Rodney & Reipurth, 2008; Rodríguez et al., 2010). Its size has been determined to be ~ 6 arcmin. The morphology is quite typical of an HII region with W40 lying next to a cold molecular cloud of angular size 1°. The space between them is described as "neutral" by Vallee (1987).

There are several energy sources inside W40, with the most prominent being three young and large IR stars giving rise to ionizing UV radiation (Zeilik & Lada, 1978). A more detailed account was given by Rodney & Reipurth (2008). They found a total of 20 radio sources half of which presented fluctuations in their flux activity over time and their results were confirmed recently by VLA observations at FIR frequencies (Rodríguez et al., 2010). The same authors speculate on the nature



Figure 1.11: The W40 cloud complex as observed by two surveys; on the left, the Digitized Sky Survey (DSS) imaged W40 at optical frequencies while on the right the Two Micron All Sky Survey (2MASS) produced an image of the cloud complex in IR frequencies (Rodney & Reipurth, 2008)



Figure 1.12: The spectrum of W40 (Finkbeiner et al., 2004). The dotted line represents the free-free model, the dashed line represents the thermal dust contribution and the thin solid line is the DL98 spinning dust model. The latter seems to be a very good fit to the data indicating a possible connection with anomalous emission.

of these sources and suggest that they may be a combination of ultracompact (UC) HII regions, young stellar objects (YSOs) and shocked interstellar gas.

Finally, a thick dust cloud heavily reduces our visibility of W40. In fact, the observations are obscured to such an extent that we cannot view a significant portion of W40 in the optical & IR bands (Reylé & Robin, 2002; Rodríguez et al., 2010).

The distance to W40 is not well-determined yet. Estimates, given by a number of authors present huge deviations from each other with the highest value given to be 900 pc and the lowest, 300 pc (Radhakrishnan et al., 1972). Rodney & Reipurth (2008) assume a value of 600 pc, which they characterize as "a conservative mean". The mass of the cloud complex has also been estimated up to an order of magnitude; approximately $10^4 M_{\odot}$ (Zhu et al., 2006). A quite accurate measurement of the magnetic field has been recorded by Crutcher et al. (1987) and it was $-14.0\pm 2.6 \ \mu G$.

To conclude, we present two important points. Zeilik & Lada (1978) observed closely the IR sources and deduced that the displacement of those, from the position where the CO emission peak appears to be, is a result of the gradual displacement of the HII membrane away from its associated core. Secondly, observations by Zhu et al. (2006) that found unusual molecular activity in a part of the W40 that no source had been previously recorded, led them to believe that stars are still formed there.

For the above reason, W40 is a very interesting and potentially useful area for studies centered around young stars and their behavior (Smith et al., 1985).

1.7 Conclusion and Thesis Outline

Deriving the spectrum of W40 at radio and IR frequencies is the primary objective this thesis. It must be completed in order to quantify whether there is a potential connection of W40 with anomalous emission.

In Chapter 2 we discuss the form and shape of the CBI W40 data, while we explain the methods we used to edit and calibrate them. Chapter 3 explains how the calibrated data were used to produce images of W40 while Chapter 4 looks into

W40 observations from other telescopes and compares them with the CBI data. Chapter 5 describes the methods used to plot the spectrum of W40. The presentation and interpretation of our results is also included. In Chapter 6 we summarise and conclude this thesis.

Chapter 2

CBI Data Reduction and Calibration

In this chapter, we will describe the observational data as well as the method and software package used to analyze them.

The CBI has already been used in deep field observations of the CMB and the steps taken to reduce and calibrate the data are thoroughly described in a series of papers (Mason et al., 2003; Pearson et al., 2003; Readhead et al., 2004). The method we have employed here is very similar.

Section 2.1 introduces the reader to the observational data. CBICAL, an "inhouse" software (Dickinson et al., 2007) written by Tim Pearson, was used for data reduction and calibration. Section 2.2 examines the procedures followed to perform the automated and manual editing. Finally, Section 2.3 explains the different steps taken to calibrate the data using CBICAL routines.

2.1 The observational data

To begin with, the observations of W40 at our disposal comprise of a total of seven days and were taken over a period of two years, starting from 2006 June 14 (observations occured during commissioning; W40 was viewed as a useful test target for commissioning of CBI2) to 2008 May 2. The duration of these observational sets ranged from 30 minutes up to 4 hours. However, on 2006 June 14, officially the first day of observations, W40 was observed for only three minutes. This made it impossible for any valuable conclusions to be reached as the data were insufficient in quantity. For this reason, in the rest of our analysis, we have always regarded as "the first day" the next available dataset taken on 2006 June 22. The total integration times, in seconds, as well as the starting time of the W40 observations for each day are shown in **Table 2.1**.

Even though there was no set time that W40 was scheduled to be observed daily, most of the observations were occurring early in the morning (contamination from the Sun is more important for CMB data and only becomes significant at angles \lesssim 40 degrees.

The primary calibrating sources used were: Jupiter, Saturn, TauA[3C144]-the Crab Nebula and Vir[3C274] along with a set of secondary calibrators (3C279, 3C273, J1743-038, B1830-210 and J1924-292) (Mason et al., 2003). Selection of these calibrators was based on the following criteria, (O'Neil, 2002) :

- the size of the source has to be significantly smaller than the beam (*i.e.* a point source),
- its flux density has to be well-defined and not varying with time and
- it must have a high signal-to-noise ratio (*i.e.* it has to be bright).

2.2 Editing with CBICAL

The "primary total intensity calibration package" (Cartwright, J.K., 2003) used was CBICAL. A summary of the typical sequence of operations we have followed to perform the data reduction and calibration using CBICAL is schematically shown in **Figure 2.1**.

Typically, the data are read from the CBI archive and plotted on an amplitude/phase vs time axis (**Figure 2.2**). The visibilities plotted can then be edited by either removing ("flag" command) or accepting data ("tracking lax", "thresh tp"

Day	Primary Calibrator	Duration(s)	Start time(U.T.)	Lead/Trail Obs	Comments
2006-06-22	Jupiter	4660.7	03:15:00	No	strong, clear signal
2007-04-03	TauA	2290.1(L)/ 2256.5	10:46:21	Yes	nothing unusual/excellent weather conditions
2007-05-23	Jupiter	1820.3(L)/1799.4	07:41:29	Yes	strong spurious signal for channels 5 and above
2007-06-07	Jupiter	5851.1(L)/5695.9	04:33:04	Yes	very long observation time
2008-04-26	TauA	2743.1(L)/2692.7	09:26:47	Yes	extended phase/amp visibilities for Jupiter
2008-04-30	Jupiter	4660.7	03:15:00	No	all W40 obs. flagged/ unusable data
2008-05-02	Jupiter	2294.3(L)/2252.3	08:24:36	Yes	spurious signal in higher channels

Table 2.1: Summary of the W40 observations made by the CBI for six different days over the span of approximately two years. The calibrator closest to W40 was Jupiter but in some cases TauA was used instead. The (L) stands for "lead" and indicates the lead observations' integration time. The fourth column presents the time of the day that the first W40 observation took place.



Figure 2.1: Schematic representation of the CBICAL data reduction/calibration process. The CBICAL tasks used are shown outside the boxes. The output was exported in UVF files. Words in quotation marks are not commands.

commands) based on various criteria. For instance, *tracking lax* allows data with larger pointing errors while *thresh tp* sets the total power range that is acceptable. Different types of calibration ensue; we start with quadrature calibration ("quad" command), follow with noise calibration ("ncal") and finally astronomical calibration ("antcal"). These are described in more detail in Section 2.3 It is important to first remove contaminated data and then apply the calibration procedures. This is because the calibration (especially the astronomical calibration) is antenna-based and thus its effectiveness will increase or decrease depending on the quality of the data.

As we will explain in more detail below, to remove ground spillover, observations of "trail" fields were performed and subsequently subtracted (from the main field observations using the command "uvsub").

The output is written in UVF files and will be thereafter loaded in the mapping package, DIFMAP (Chapter 3). Eventually these UVF files will be combined in one final map using the UVCON procedure (Chapter 3).

2.2.1 Spillover Rejection

The origin and effects of ground spillover were discussed in Section 1.5. Here we will look into the way we dealt with this limitation. The ground emission is constant over short timescales (of the order of minutes) and as such, two fields at the same Declination and Elevation will carry the same contribution from the ground pickup in their signal. Assuming that one of those fields is the primary observational field, *i.e.* our source of interest ("lead") then by observing another field at a different Right Ascension (RA) that satisfies the aforementioned conditions ("trail") and subtracting them from each other, we can remove the ground contribution. The difference in RA was eight minutes (corresponding to 2 degrees) in accordance with the method suggested by Padin et al. (2002). Still, this subtraction is not without any side-effects. While it will remove other contaminants in the signal (Mason et al., 2003) it will also reduce the sensitivity of the CBI by a factor of $\sqrt{2}$ as we will be measuring the ground half of the time. (Padin et al., 2002). For calibration of W40, Jupiter was mostly used except in occasions where it was observed as having an extended phase/amplitude, possibly due to a pointing error in the system, *e.g.* on 2008 April 26. In such cases, Jupiter became unsuitable as a calibrator and TauA was used instead. An example of such a baseline is shown in **Figure 2.4**.

It should be noted that the flux densities of the primary calibrators were determined directly from observations with the CBI relative to Jupiter, for which a brightness temperature of 152 ± 5 Kelvin in the Rayleigh - Jeans limit was assumed (Readhead et al., 2004). The total uncertainty in the calibration was assumed to be 5 %. This value was derived by a combination of factors: the error on the absolute calibration is ~ few % but an additional small percentage is added due to other limiting factors such as phase errors. Even so, we state that this is a rather conservative value.

2.2.2 Automated Editing

Below we summarize the most important points concerning the automated data editing and error removal as it was performed by CBICAL.

- An internal "noise calibration source" was used for the reduction of gain and phase fluctuations as well as for quadrature and atmospheric calibration.
- The datasets were compiled such that the lead/trail field scans are bracketed by observations of the primary and secondary calibrators, typically Jupiter and J1924-292.
- CBICAL provides automatic flagging of "bad" data *i.e.* data that do not conform to a specified antenna representing nominal operational values. By "flagging" one typically means the removal of bad data so that they are no longer needed; essentially CBICAL assigns a zero weighting on those data so that they are ignored in any subsequent use. Fomalont & Perley (1999) note that the term was originally used in a different context; that of marking data of ambiguous nature. Bad data are usually a product of the inability of the

system's components to function properly. Thus, most commonly, CBICAL would flag data during periods where a receiver's temperature was too high (>300 K; this caused the production of a very noisy signal), if there was an unlocked phase circuitry of the local oscillator/synthesizer, if there were errors in the telescope's tracking and if the power meter was out of range.

2.2.3 Manual Editing

After CBICAL had performed all the automated editing, we had to manually examine the data and flag or unflag portions of them. Even though the latter is not generally recommended, in some cases good data were found "buried" in low quality data while at other times we allowed the admission of moderate quality scans simply because the purely good ones were not sufficient.

As a first indicator in this procedure, we used the observers' log, a set of notes that contain information about any noticeable event that might have occurred before and during the observations, usually involving technical difficulties associated with the receivers. Descriptions of the weather conditions of each day were also included. In other words, the notes were never treated in a conclusive manner but rather they were used as a guide. In general, the log was found to be in accordance with CBICAL's automated data reduction procedures. For this reason, most of our involvement in data editing came from careful examination of the baselines for each of the six days that W40 was observed.

Generally, the decision to flag a specific portion of the data was based on three indicators: the working status of the receiver, the form of the phase/amplitude of the calibrator or whether some sort of non-astrophysical (often, called "spurious") signal existed.

For the first case, one can verify the condition of a receiver by the number of automatically flagged frames. If, for a significant number of data generated, this is indeed the case then it is most likely that the receiver in question is problematic.

Further verification can be obtained by investigating the power plots (**Figure 2.5**). Ideally, the power plot should be flat at around 1. Deviations that were



Figure 2.2: CBICAL screenshot of plotted visibilities of baseline [10-12] for channel 6 for an entire night of observations. In the top two panels, the blue points represent the good data on source, the grey points are the flagged data and the red/green points are the good data on the noise diode. Telescope specifics are located at the bottom of the graph (azimuth, elevation, declination, coordinates, diagnostics, informatics). The phase and amplitude of the calibrator sources presents strong localization, are well-defined and significantly more intense than the signal coming from the background indicating the receiver is working. The strong concentration that the visibility data present in the interval bracketed by the calibration observations are regarded as "spurious" signal.



Figure 2.3: CBICAL screenshot of plotted visibilities of the same baseline and dataset as in **Figure 2.2** concentrating solely on the time period where W40 was observed (3 hour observation). Observations of the calibrator (Jupiter) are also included for comparison. The colour scheme is the same as in the previous figure. The green colour indicate that the calibration was successful. The phase of W40 shows significant dispersion which is in accordance with the length of the baseline; another indication that the receiver system works smoothly.



CBI DATA REDUCTION AND CALIBRATION

CHAPTER 2.

Figure 2.4: Plotted visibilities for channel 1 of baseline [1-12]. The colour scheme is the same as in **Figure 2.2** The amplitude and phase of Jupiter are extended and hence cannot be considered a point-source. In comparison, TauA has better defined properties which allow its use as a calibrator.

accepted were up to 10 % of that value. A faulty receiver would usually produce a very unstable signal with values of the power output ranging from 10^{-2} to 10^{-3} which is much lower than our acceptance limits.

Moreover, the functionality of a receiver can be checked through the number of closure errors, amplitudes and phases it produces. The closure errors are recorded in a .log file, which is produced by CBICAL after all the data are read, along with other information related to the processing of the data by the program, in line with proposals by Fomalont & Perley (1999). We generally found that it was best to flag completely all the data from receivers that were associated with multiple closure errors. We regarded as acceptable closure errors anything less than ten degrees, since above that limit they become a significant source of error on the amplitude. The same limit was assumed for closure errors and amplitudes.

In addition, distinguishing between low and high quality data was based on how well-defined the phase/amplitude of the calibrating sources on the plot were. We expect stronger concentration and larger amplitudes in their visibilities when compared to the signal from the background sources which was bracketed in between the observations of the calibrators (**Figure 2.2**).

Noise levels can also be an indicator of a good baseline; these should be higher than any of the plotted visibilities, as shown in **Figure 2.6**.

The background sources produced a random Gaussian noise pattern in phase (or a Rayleigh distribution in amplitude) except if a specifically bright source was observed (such as W40). In these cases, a wider and more complex phase pattern was noticed for longer baselines (*i.e.* for higher resolutions) as the source's structure becomes more resolved and the point-source approximation does not apply anymore. On the contrary, for shorter baselines, this is not a concern and we get a tighter pattern (**Figure 2.3**).

Finally, we investigated the nature of the signal. In general, unless a strong source is observed we should not expect to see strong localization in the phase plot; the rms spread of the data points/visibilities should be about the same as that due to random noise. If such was not the case, we dubbed the signal as "spurious" (Section 2.2.3).



Figure 2.5: Total power data for channel 2 from receivers 0 to 12 in the time interval where W40 was observed. The red line at the top of the graph represents the ambient temperature. The colour scheme of the data points is the similar to the one of the visibility plots. The power output from receivers 6, 8, 11 and 12 is stable and close to 1 indicating that their status is operational. Receiver 1 produces the lowest output, with a very unstable signal and therefore should be flagged.



Figure 2.6: Magnitude comparison between noise levels and visibilities for channel 4 from baseline [1-5] on 2008 April 26. The colour scheme is as described in the previous figures. The red line at the top half of the graph indicates the noise levels. The blue line near the zero level amplitude are the plotted visibilities.

An example is shown in **Figure 2.2**. In essence, this type of problem is accounted by the observation (and subsequent subtraction using UVCON; Section 3.3) of the lead and trail fields we described in Section 2.2.1. In a few cases, however, we manually had to discard some scans that were very strongly contaminated, and therefore may lead to residual contamination even after the lead/trail field observations have been subtracted.

2.3 Calibration

2.3.1 Quadrature Calibration: QUAD

The quadrature calibration is closely related to the output of the correlator, which lies at the bottom of the CBI signal path (**Figure 1.9**). At the output, the correlator signal (*i.e.* the visibilities) will be a complex quantity; it will have a real and imaginary component, which should be identical except for a 90° phase difference, owing to the manipulation of part of the signal by the phase shifter while at the "correlator" stage, as we described in Section 1.5. In principle, this will only be the case for a point source at the centre of the map. Then, realistically speaking, this rarely the case and as such the purpose of the quadrature calibration is to ensure the orthogonality of these components by measuring the amplitude-phase offset and subsequently estimating and applying the quadrature calibration corrections on the imaginary component of the signal, if need be. Based on what has been said in Section 1.4 about the potential causes of closure errors, we expect their number to decrease after applying QUAD.

Failure of QUAD can be due to three reasons: either the amplitude of the signal we are trying to calibrate is too low, the visibilities we are provided with do not suffice in quantity for accurate estimates of the correction factors that will be applied, or the correction factors themselves are too small/large and hence their application is not possible by CBICAL.

2.3.2 Noise Calibration: NCAL

By noise calibration we mean the comparison of the systematic gain fluctuations and phase errors in the receivers with respect to a reference source. This is made possible by the implementation and use of an internal noise calibration source (Section 1.5). The key element here is the stability of the power output of this noise source. Mason et al. (2003) point out that the source's temperature could not be kept easily constant during the observations, especially for extended periods of time. This made the source's output even more unstable than the gain fluctuations. Consequently, to perform noise calibration we used CBICAL's NCAL command which normalizes and averages each one (the source's and the gain fluctuations) before applying a global correction of, typically, a few % at low elevations (~ 50).

This allows NCAL to compute appropriate calibration factors and correct the visibilities accordingly.

2.3.3 Antenna Calibration: ANTCAL

This procedure finds the best fitting value to correct the amplitude and phase to the reference values given by the calibrator. The utility of the ANTCAL task rests in how accurately the antenna-based complex gain has been determined since CBICAL uses Equation 1.44 to determine the correction factors to be applied. From Section 1.5, we know that that the complex gain is directly related to the number of closure errors, hence for a successful result to be reached their number needs to be reduced as much as possible, usually by means of manual editing.

Likewise, plotting the visibility statistics (Figure 2.7) of the calibrator can reveal plenty of information about how successfully the antenna calibration has been performed. As ANTCAL forces the phase of the calibrator to zero, it should not present any deviations from the zero level. At the same time, the rms noise for each baseline should vary by, at most, a factor of 2. Because fluctuations in the amplitude on an individual baseline cannot be fixed by ANTCAL (because they are antenna-based), for a successful antenna calibration, we expect them to be small.
Note that in **Figure 2.7** the slope in the amplitude observed is because of Jupiter's spectrum. Similar to a typical B-B spectrum, it is roughly flat in temperature T at high frequencies while it gets brighter as we move towards the far-IR regime (Readhead et al., 2004).

Lead - Trail Field Subtraction: UVSUB

The program UVSUB was employed to perform the spillover rejection. This is attained by the matching of the u-v coordinates of the lead and trail fields and the subsequent subtraction of their visibilities.



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Figure 2.7: Plot of the visibility statistics for Jupiter (calibrating source) for the 2008 May 02 observations. The individual points represent the baselines, while the numbers to the right indicate the colour scheme assigned to each channel.

Chapter 3

CBI Imaging of W40

In this chapter, the procedure used to image the reduced and calibrated visibility data of Chapter 2 is described. The imaging software DIFMAP (Shepherd et al., 1995), which is often used to extract flux from visibilities as an "easy" alternative to more established imaging packages such as AIPS, plays a central role in achieving this goal.

Section 3.1 introduces DIFMAP while it also describes a typical mapping sequence. In Section 3.2 we inspect the quality of the images produced, based on three parameters (peak flux, average RMS noise and dynamic range), while we attempt to reach to some early conclusions about W40. In Section 3.3 a brief description of the CBI program UVCON, which was used to combine the several outputs of DIFMAP and produce two finalized images of W40 from CBI, is included. The rest of the Section presents and discusses these final images.

3.1 Imaging in DIFMAP

3.1.1 Difference Mapping

Shepherd et al. (1995) explain that DIFMAP employs the technique called "Difference Mapping" (from which DIFMAP derives its name) to form the final image. In difference mapping, one builds up a "source" model through successive applications of the CLEAN algorithm (see Section 1.4) on the dirty map. The subtraction of the flux, *i.e.* the application of CLEAN, is done in DIFMAP by "drawing" CLEAN windows around the source and the brightest points on the map. Eventually, the model is subtracted from the dirty map and the end result is the residual or the "difference" map. The last step involves inverting the dirty beam (or "residual" map) to produce the final map with all the noise effects removed.

3.1.2 Preliminary Considerations

Polarisation and Mapsize

In a typical mapping sequence, the calibrated visibility data or UV-FITS (.uvf) files were read from the memory (the output of CBICAL and UVSUB; Section 2.2) and loaded into DIFMAP as the input. Choosing the polarization of the data to work with was also required. There are several options that DIFMAP recognizes (*e.g.* left-circular/LL, right-circular/RR, intensity/I) but the pseudo - intensity (PI) was mostly used. The PI polarization assumes LL visibilities are equal to RR visibilities and uses either one since CBI is configured to measure LL polarization data. When both are available, it combines them to find the weighted mean and sets I equal to that value. For unpolarised sources, measuring LL (or RR) is equivalent to measuring the total intensity, I. Therefore selecting PI as the polarisation of the visibilities is appropriate for these data.

The map size used was 1024×1024 with a pixel size of 0.5 arcmin. This map size covered several times the primary beam, which sets the field-of-view (4.1°, in this case). The calibrated visibilities were then binned into a 2-D array and Fourier inverted (this essentially involves the inversion of Equation 1.37). Thus the production of the distribution of the sky brightness or the "dirty map" became possible. Typically, after formation of the dirty map, one has to move the source to the centre of the map; here the source was centred by default.

Weighting

There are two weighting schemes available in DIFMAP: natural and uniform weighting. The latter is the default option. Natural weighting assigns the same weighting factor to data from all baselines. As shorter baselines are associated with data of lower resolution, and since shorter baselines are often more in number than the longer ones, it follows that treatment of all visibilities in the same way will only result in degrading the resolution to that of the shorter baselines. At the same time, this means that noisy data will average out causing a general reduction in the noise levels of the final image.

On the contrary, when using uniform weighting DIFMAP assigns a larger weighting factor to areas with a smaller number of visibilities (and vice versa). Despite, allowing finer details to be more visible in the final image (*i.e.* increases the resolution) this procedure also introduces the risk of assigning a large weighting factor to data from a noisy baseline. The result is a larger noise level, relative to the one produced if natural weighting is used, at the expense of a smaller beam.

In imaging W40, the author experimented with both weighting methods. In general, it was found that uniform weighting decreased the peak flux density (due to the small beam) compared to the results of natural weighting. Hence, uniform weighting was selected to provide more resolution in the image; this proved to be the ideal choice for the CBI data for which the main limitation is dynamic range not thermal noise (Section 3.2.1).

3.1.3 Practical Application of CLEAN

To CLEAN, we had to choose values for three parameters:

- niter (= 100), which is the maximum number of CLEAN repetitions that will be performed, for each iteration of CLEAN
- loop gain (= 0.03), which is the amount of flux subtracted after each iteration, and

• cut-off (= 0), which determines the flux density at which the CLEANing will stop.

The selected values are shown inside the parentheses and they were kept fixed for the whole of the duration of the CLEANing process.

As mentioned at the start of this chapter, the flux subtraction is performed by drawing windows (boxes) around the brightest points in the map. As the brightest point will always be the source, traditionally one draws a single window around it to CLEAN the maximum amount of flux straight away.

Most importantly, after each CLEAN iteration, the deconvolved map remains in memory so that it becomes possible to switch back and forth between the data and dirty map and observe how changes in the former affect the latter. As a result, re-editing of the data occurred several times in order to improve the quality of the dirty map (*i.e.* to reduce the noise levels).

A compilation of the maps produced for each day of observation of W40, after the subtraction between the lead and trail fields, is shown in **Figure 3.1** and are discussed in Section 3.2.

3.2 Image Quantification

In Figure 3.1, W40 appears to be a compact source, with mostly spherical structure presenting a slight extension towards the South. No significant extended emission is observed. The detection of artifacts on 2007 April 03 and 2008 April 26 could be attributed to particularly sensitive measurements on those particular days of observation. At the same time, the negative values observed could have been caused by many reasons; most commonly they are the result of gaps in the u-v plot due to the removal of low quality/problematic data during the calibration and editing procedures (Sections 2.2 and 2.3 respectively). However, often they arise because of gaps in the baseline coverage. Filtering effects, such as the subtraction of the sky emission (Section 2.2.1) can also cause the appearance of negative brightness. Slight phase errors (typically of 10° to 20°) are only documented on 2006 June 22, while in



Figure 3.1: Maps of W40 produced by DIFMAP for each one of the six days of observation. The trail field observations have been subtracted. The colourbar applies to all of the pictures and the units are Jy/beam. The size of the image is $2.4^{\circ} \times 3.5^{\circ}$

Observation	Peak(Jy/beam)	Av. RMS Noise (Jy/beam)	Dyn. Range
2006-06-22	7.8	0.05	152
2007-04-03	9	0.11	86
	(SR) 9.36	0.09	104
2007-05-23	10.5	0.71	15
	(SR) 10.79	0.03	317
2007-06-07	10.78	1.86	6
	(SR) 11.47	0.03	364
2008-04-26	10.54	0.11	98
	(SR) 11.0	0.09	122
2008-05-02	10.0	0.05	192
	(SR) 10	0.04	260
Unedited data	10.16	0.04	256
Combined data	8.2	0.033	248
	(SR) 8.6	0.027	321

Table 3.1: Comparison of the values of the peak brightness, average RMS noise and dynamic range for each of the CBI CLEANed maps produced with (bottom line) and without (top line) spillover rejection. The sign (SR) denotes that spillover rejection has been applied. No trail field observations occurred for the 2006 June 22 dataset. Values for the completely unedited data (taken on 2008 May 02; see text) are also shown. The last line presents the values of the four parameters for the UVCON combined image before and after spillover rejection (**Figure 3.5**)

all the other cases, some amplitude errors are present at around the same levels as the phase errors. These can be distinguished by the degree of symmetry they hold; amplitude errors are usually asymmetric in the map while the opposite is true for phase errors. The low noise levels observed on 2007 May 23, 2007 June 07 and 2008 May 02 were attributed to the larger, comparatively, number of visibilities available for mapping *i.e.* the data on these particular days were not edited as heavily as in others. Therefore, this resulted in better u-v coverage.

A comparison was also made between the final maps produced in Section 3.1.3 after spillover rejection had been applied (**Table 3.1**). Values of the preliminary maps, before the aforementioned subtraction occurred, are included to quantify the difference between the two methods. The analysis is based on the following parameters: average rms noise, peak flux density and dynamic range. **Table 3.1** presents the values of the four parameters for each day of observations of W40. Certain aspects of the maps with regards to those parameters are also discussed.

The average RMS noise was found by calculating the average value between the



Figure 3.2: A schematic representation of the way the average rms noise was calculated (see text). The size of the boxes was kept small so that no flux is included in the noise estimates. Here an arbitrary date has been chosen; 2007 May 23.

noise levels in each of the four corners of the map as that was returned by DIFMAP's "statistics" command (Figure 3.2). Because these regions lay the furthest from the image, they hold the lowest possibility of getting contaminated by the source's emission. Thus an objective estimation of the noise levels is allowed. The average noise levels derived were of the order of $\sim 1\%$ of the peak flux density. This indicates that at 31 GHz, the background contribution is largely invisible for the W40 cloud complex. The lowest values for the average RMS noise belong to data taken on 2007 May 23, 2007 June 07 and 2008 May 02; a result which is agreement with what is seen in **Figure 3.1**. The shape of the source remained similar over the duration of observations with the exception of the 2008 April 26 map where an elongation was observed. Several hypotheses were examined regarding this observation. First of all, if this was a detection of a true extension in the shape of W40 then one would expect more than one detection to have occurred in the eight samples presented here. Another possibility lies in the shape of the beam; a discrepancy in the shape of the beam could indeed have caused such an elongation. However the beam, as drawn by DIFMAP during the mapping procedure of W40, did not present any significant deviation from its usual shape. It follows then that the most plausible explanation at the moment is that this is the result of an observation cut short for unknown reasons; an abrupt stop while CBI was scanning the sky could have caused this anomaly.

The statistics command also returned values for the peak brightness of W40, essentially the value of the brightest point on the source. This varied from ≈ 9 to ≈ 11.5 Jy, leading to an average of ≈ 10.5 Jy over all days of observations after rejection of ground spillover.

3.2.1 Dynamic Range

This quantity is defined as the ratio between the peak brightness and the average RMS noise (Perley, 1999). In principle, the dynamic range is a tricky concept, because it does not necessarily quantify how accurate the deconvolved image is compared to the true image. Perley (1999) states that, generally, a high dynamic range means lower errors in the image only because its meaning is associated with reduction of potential data contamination factors. As such, dynamic range is assumed to be a "good indicator" of the difference between the deconvolved image and what truly lies out there, although it must be pointed out that sensitivity to faint features has often been recorded.

It is possible to compare the noise with the sensitivity of the interferometer and determine the extent of the dependence of the CBI observations from thermal noise. A variation of the standard formula for the sensitivity (Equation 1.41) can be used for that purpose:

$$\Delta S = \frac{2k_B T_{sys}}{A_{eff} \eta \sqrt{n(n-1)(\Delta \nu \times \tau)}}$$
(3.1)

where all the symbols are as defined in Equation 1.41, ΔS is the minimum flux density in Jy, n is the number of receivers and A_{eff} is the effective area for each disk of the CBI2. Substituting the appropriate values:

- $\Delta \nu = 10 \text{ GHz} \text{ (Section 1.5)}$
- $\eta = 0.86$ (Section 1.5)
- $A_{eff} = 1.54 \ m^2$ (Section 1.5)
- $\tau = 24318$ s, by summing all the integration times of the lead field observations shown in Table 2.1
- n = 13, since there are 13 receivers on CBI (Section 1.5)
- $T_{sys} = 27$ K, by summing all the elements of **Table 1.1**

it is found that the theoretical noise level is ΔS is 8.61×10^{-3} Jy, where $\langle RMS \rangle$ is the value of the average RMS noise of W40 as shown in **Figure 3.5**.

Then by comparing with the effective noise level (**Table 3.1**):

$$\frac{\langle RMS \rangle}{\Delta S} \approx 3.13 \tag{3.2}$$

Hence the effective noise was found to be greater than the theoretical noise which means that the CBI data are not limited by thermal noise but by the alternative: dynamic range.

As it is evident in **Figure 3.1**, the removal of ground spillover increased (significantly in most cases) both the dynamic range and the peak brightness. In parallel, the noise levels were reduced and the resolution was improved.

We were also interested in finding out how good CBICAL was in data editing and calibration without the interference of the human factor. To do this, we produced an image of W40 based on the visibilities of a randomly selected day (2008 May 02) without manually editing out any bad baselines. The results are shown in the last line of **Table 3.1**. Evidently, the improvement in the dynamic range is marginal (less than 2%) indicating that the default automatic editing within CBICAL works well for these data. It is interesting to note that the value of the dynamic range for this particular day is average in comparison to the ones of the other days; the lowest dynamic range was ≈ 103 and the highest was ≈ 364 . This means that there was nothing special about the specific dataset making this conclusion even more robust. As a result, we can confidently argue that this is evidence of the remarkable brightness of W40 as a celestial body.

3.3 Final Map of W40 at 31 GHz

Combining The Images: UVCON

After imaging the visibilities, all the FITS files were combined and exported by DIFMAP into one final map using the CBI program, UVCON. To produce the combined version, UVCON first checks if the files provided at the input have the same properties. For example, whether the observational target is identical and if the number of channels/receivers is the same. Once it asserts that all input files are compatible, the output file is produced by summing all the visibilities in the input files. Two final maps were produced, one with (B) and one without (A) the lead and trail fields subtracted. They are shown in **Figures 3.4** and **3.5** respectively. The



Figure 3.3: The (u,v) coverage of the CBI as used for the observations of W40. Here we have used the data combined over the 6 days of observations and after subtraction of the ground emission has been performed.



Figure 3.4: Final map of W40 produced by combination of the 6-day CBI data using UVCON. Spillover rejection has not been applied. The units are Jy/beam and the size of the synthesized beam is 4.02×3.66 arcmin. The size of the image is $3^{\circ} \times 2^{\circ}$

u-v coverage of the CBI corresponding to the final map without the ground emission is shown in **Figure 3.3**. The plot starts at $0.13 \times (10^3 \lambda)$ and ends at $0.62 \times (10^3 \lambda)$, in wavelength units. The fact that the u-v plot is a complete circle with few, if any, gaps is evidence in support of the CBI's excellent u-v coverage (Section 1.5).

Spillover Rejection in Final Map

The positive effects of spillover rejection are evident when comparing Figures 3.4 and 3.5 where the structure of W40 appears more defined, while it also looks slightly brighter. In particular, the peak flux was 8.2 Jy before the removal of sky/ground emission and 8.6 Jy (slightly larger beam) afterward. In both cases the u-v coverage (seen at the background; also see Figure 3.3) is well-defined but the darker background on the right is indicating lower noise levels. In fact, the estimated noise



Figure 3.5: Final map of W40 produced by combination of the 6-day CBI data using UVCON. Spillover rejection has been applied. The units are Jy/beam and the size of the synthesized beam is 0.073 arcmin × 0.064 arcmin. The size of the image was $3^{\circ} \times 2.5^{\circ}$.

level without spillover rejection is 0.033 Jy/beam while after spillover rejection it is reduced to 0.027 Jy/beam pointing to the significance of the ground emission as a source of error. At the same time, this subtraction could be the cause behind the detection of the larger number of negatives detected in **Figure 3.5**.

Chapter 4

Ancillary Data of the W40 region

In this chapter a comparison between the final image of W40 at 31 GHz, produced by CBI data, with images of W40 from other surveys is made. In this way, a study of the morphology of W40 in radio and IR frequencies was achieved.

In Section 4.1 we outline the procedure followed to download the different survey images, using two electronic services: Skyview¹ and The Survey Sampler². At the same time we assess the quality of each survey and its importance in describing W40. Aligning and smoothing the maps to the same resolution as well as converting them to the same units of flux density is explained in Section 4.2 and Section 4.3 respectively. Completion of these tasks laid the foundations of the production of spectral energy distributions that will be discussed in Section 5.2.

4.1 Description and Acquisition

The different survey images of W40 were acquired through two digital services: NASA's Skyview, and the Survey Sampler of the Max Planck Institute. Skyview holds a wider range of surveys for all the frequency bands however certain important specialized datasets can only be acquired through the Survey Sampler.

Specifically, Skyview was used to download:

• the 12, 25, 60 and 100 μ m Infrared Astronomical Satellite (IRAS) maps,

¹http://skyview.gsfc.nasa.gov/

²http://www.mpifr-bonn.mpg.de/survey.html

- the SFD maps at 100 μ m and E(B-V) Reddening/map extinction
- the Ka, K, Q, V and W-band WMAP 5-year maps,
- the NRAO VLA Sky Survey (NVSS) survey at 1.4 GHz (Section 5.4.1)
- the 408 MHz Haslam map
- the H α full sky map.

The data were downloaded in J2000 equatorial coordinates and the projection was sin-orthographic. Galactic coordinates could have been used but this would only change the alignment of the image with the 2D axis (*i.e.* the x, y axis) so there is not a substantial difference. We requested a pixel size of 0.5 arcmin (Chapter 3) corresponding to an image size of 600×600 pixels (or $5^{\circ} \times 5^{\circ}$). All images were centered on W40.

The Bonn (Max Planck Institute) Survey Sampler was used to download:

- the Parkes 6 cm survey,
- the Effelsberg 11 and 21 cm surveys and
- the Stockert 25 m survey,

where we used a map of size $10^{\circ} \times 10^{\circ}$ and a tabular interval of $[0.5 ' \times 0.5 ']$ (in x and y).

It should be pointed out that in the Parkes 6 cm survey, W40 is not visible because it lies outside the latitude ($b = \pm 2^{\circ}$) and longitude range (238° < l < 5°) covered by the survey. Additional information as well as a summary of the properties of the 9 surveys are shown in **Table 4.1**.

Figure 4.1 shows the images downloaded from all the surveys including the maps we have produced with the CBI.

The situation at radio frequencies appears to be consistent throughout the survey images, with W40 appearing as a very bright source that presents some extension but still retaining a roughly spherical shape. At low frequencies, below 1 GHz, the

Dataset	Frequency	Resol.	Regime	Map Units	Notes
Haslam ^a	408 MHz	0°.85	Radio	K	All-sky map dominated by synchrotron emission
Stockert 25 m b	1.4 GHz	35'.4	Radio	K	Northern-sky map
Effelsberg 21 cm^c	1.4 GHz	9'.4	Radio	$mK(T_b)$	Medium latitude survey. Useful for extended emission studies.
Effelsberg 11 cm^d	2.76 GHz	4'.3	Radio	$mK(T_b)$	Most sensitive galactic plane survey in this range
$GB6/PMN^e$	4850 MHz	3'.5 - 4'	Radio	Jy/beam	Combination of 2 surveys filtered on angular scales $\gtrsim 15$ '.
Parkes 6 cm f	5 GHz	4'.1	Radio	$mK(T_b)$	Galactic plane survey; latitude less than 2 (W40 not covered).
CBI	31 GHz	4 - 5'	Radio	Jy/beam	Survey of the flux densities of 1.4 GHz sources
WMAP K ^g	22.8 GHz	0°.88	Radio	mK(CMB)	5-year Galactic plane survey reproduced from Healpix maps
WMAP Ka	33 GHz	0°.66	Radio	mK(CMB)	5-year Galactic plane survey data reproduced from Healpix maps
WMAP Q	40.7 GHz	0°.51	Radio	mK(CMB)	5-year Galactic plane survey data reproduced from Healpix maps
WMAP V	60.7 GHz	$0^{\circ}.355$	Radio	mK(CMB)	5-year Galactic plane survey data reproduced from Healpix maps
WMAP W	93.5 GHz	$0^{\circ}.22$	Radio	mK(CMB)	5-year Galactic plane survey data reproduced from Healpix maps
SFD 100 μm^{-h}	3 THz	6'	Infrared	MJy/sr	Combined product of COBE/DIRBE and IRAS/ISSA maps.
SFD Dust Map	3 THz	6'	Infrared	Reddening	Combined product of COBE/DIRBE and IRAS/ISSA maps.
IRAS 100 μm^{-i}	3 THz	$\sim 2'$	Infrared	MJy/sr	Some NULL values in the maps.
IRAS 60 μm	5 THz	$\sim 2'$	Infrared	MJy/sr	Some NULL values in the maps.
IRAS 25 μm	12 THz	$\sim 2'$	Infrared	MJy/sr	Some NULL values in the maps.
IRAS 12 μm	25 THz	$\sim 2'$	Infrared	MJy/sr	Some NULL values in the maps.

Table 4.1: A comparison of the properties of 9 different datasets derived from a different telescope each. Images of the W40 region from each survey are going to be compared against the CBI maps made via the procedure described in Chapter 2. T_b denotes units of brightness temperature and CMB denotes thermodynamic units (Section 4.3.3). References for each survey are shown in the footnotes.

^{*a*}Haslam et al. (1982)

- ^bReich & Reich (1986)
- ^cReich et al. (1990)
- ^dReich et al. (1984)
- ^eCondon et al. (1991)
- f Haynes et al. (1978)
- g For all WMAP surveys: Hinshaw et al. (2009)
- h For the SFD surveys: Schlegel et al. (1998)
- ⁱFor the IRAS surveys: Beichmann C.A., Neugebauer, G., Habing, H.J., Clegg, P.E., Chester, T.J., 1985, IRAS Explanatory Supplement

Haslam map is not as well-pronounced, because of the diffuse synchrotron radiation from the Galactic plane, implying that the spectrum of W40 is not dominated by synchrotron emission. The GB6 survey contains multiple bad data, making the image one of very poor quality. The only conclusion that can be made here is that most of the background emission observed in the other images is resolved out, making W40 look similar to a point source. Some extended emission can be seen encompassing the bright core, as well.

At 1.4 GHz, the Effelsberg 21-cm map is subject to some blending with, possibly, emission, from the Galactic plane while remaining largely similar to the 2.76 GHz map.It also presents the highest value of the peak brightness of W40. Large amount of ionized gas is seen at the South and South-East (lower left) direction linking W40 to the Galactic disk. Low-level synchrotron emission could also account for this extended emission. Data from the Stockert 25 m survey at the same frequency because of the significant blending to be seen between the emission from the Galactic plane and the source. The slightly higher resolution of the Effelsberg 11-cm map allows us to discern more details of the background structure and W40's morphology at the edges. Part of the gaseous structure described above is not visible anymore in the South indicating that the gas/dust that form it emit more strongly at lower frequencies. It is also possible that this could just be due to filtering in the survey. It is noteworthy, that the coverage of both the 21- and the 11-cm maps is cut unexpectedly in the N-W (upper right) region of the image at a distance of 16 and 90 arcmin respectively, as shown in Figure 4.1.

Evidently, the CBI cannot detect the faint parts of the emission of W40; as it was proved, the CBI images are limited by dynamic range (Section 3.2.1) and will resolve out large angular scales ($\gtrsim 25$ arcmin). The CBI data can be easily compared with both the GB6 and the Effelsberg 11 cm surveys as they are at a similar resolution; a relative agreement in the morphology of the source is observed. It is not possible to draw any conclusions about the structure of W40 from the WMAP data since the resolution of all the maps is low.

At IR and far-IR, the bigger dynamic range of the IRAS surveys and the possibly

greater dust population in the specific region makes it possible to observe faint and extended emission from W40. This fact, in combination with the high resolution, makes the IRAS surveys very useful. The IRAS 100 μ m image differs significantly from the 12 μ m one, because of its detection of bigger grain sizes. Lastly, the 12 and 25 μ m images are the only IRAS surveys that show some contamination due to spectral lines in their data. The SFD at 100 μ m agrees well with the original IRAS 100 μ m, and the peak brightness derived from both surveys is similar.

Finally, the SFD Dustmap was used to measure the dust absorption of the H α recombination line. This was found to be ~ 50 magnitudes. A correction factor can be derived using the standard definition of the apparent brightness of any celestial body i.e.(Kraus, 1986),

$$m = -2.5 \log_{10} \frac{F}{F_0},\tag{4.1}$$

where F is the observed flux of the source and F_0 is a reference flux. The correction factor was found to be of the order of 10^{20} indicating that, indeed, the H α line cannot be used in any kind of measurements (Section 1.3.1).

4.2 Manipulation of the Ancillary Data

4.2.1 IDL : Alignment and Gaussian Smoothing

The main goal of this project is to derive a plot of the spectrum of W40 *i.e.* derive values of its flux density at different frequencies. The methods employed to achieve this goal (aperture photometry, 2D Gaussian fitting) are discussed in the next chapter. Nevertheless to ensure that the flux densities are on the same scale (*i.e.* sensitive to the same emission) all the images had to be smoothed to the same resolution as well as converted to the same units. We chose brightness units of Jy/pixel.

Both the alignment and smoothing were performed using scripts written in IDL, a programming language used widely in astronomical data analysis. The input files



Figure 4.1: Multi-frequency data of W40 at their original resolution. The name of each survey is shown on the top left corner of the image. All maps are scaled in the same way and are shown in celestial coordinates (Right Ascension, Declination). The size of the region is $1^{\circ}.6 \times 1^{\circ}$. The Parkes 6 m and Haslam 408 MHz surveys are omitted as W40 was not visible in any of them. See text for more details. were in the standard FITS format. The maps were aligned using the CBI image as the reference map.

The idea behind "Gaussian" smoothing, used here, involves convolving the old image with a smoothing "kernel" (*i.e.* a matrix of numbers) to produce a new version of the image at the desired resolution. In other words, the kernel is used to increase the beam of the original map. To find the resolution of the smoothing kernel we had to regard the images we have downloaded (or mapped) as Gaussian functions.

Now, consider two Gaussian functions, g_1 and g_2 , such that:

$$g_1(x,y) = A_1 e^{-(x^2 + y^2)/2\sigma_1^2}$$
(4.2)

and

$$g_2(x,y) = A_2 e^{-(x^2 + y^2)/2\sigma_2^2}$$
(4.3)

where $A_1 = \frac{1}{\sqrt{2\pi\sigma_1}}$ and $A_2 = \frac{1}{\sqrt{2\pi\sigma_2}}$.

The convolution between two Gaussians gives another Gaussian, such that:

$$g_1 * g_2 = \frac{1}{\sqrt{2\pi(\sigma_1^2 + \sigma_2^2)}} \exp\left(\frac{-(x^2 + y^2)}{2(\sigma_1^2 + \sigma_2^2)}\right)$$
(4.4)

and comparing the above to the general form of a Gaussian, that is,

$$g(x,y) = \frac{1}{\sqrt{2\pi\sigma}} e^{-(x^2 + y^2)/2\sigma^2}$$
(4.5)

it is obvious that

$$\sigma = \sqrt{(\sigma_1^2 + \sigma_2^2)} \Rightarrow \sigma_2 = \sqrt{(\sigma^2 + \sigma_1^2)}$$
(4.6)

where σ is the resolution of the final image in pixels, and σ_1 , σ_2 are the resolutions of the initial image and the kernel in the same units, respectively. In other words, the width of the resultant Gaussian is the sum of the 2 Gaussians added in quadrature.

The image is an array of pixels and to convolve an array with a kernel we used

IDL's CONVOLVE function. To define the kernel the function PSF-GAUSSIAN was used. The size of the kernel has to be such that it covers the whole curve up to the tails. This is, typically, equivalent to a pixel number (or "an array size") of five to six times the FWHM. The Gaussian is normalized to a unity area so not to change the brightness scale of the original map.

As stated above, part of the GB6 and Effelsberg maps is missing. This translates to "bad" values in the image arrays *i.e.* values that are too high or too low to be processed by IDL. Usually, these values are equal to -32000 or -1.6375×10^{30} . This can cause errors and prevent the script from running smoothly. To deal with this problem, we set all the bad values equal to standard NaN values, essentially forcing IDL to treat them as missing data. Of course, this affects the smoothing of any map containing bad values. In particular, the output of GB6 will be null for any smoothing at any resolution whilst for the Effelsberg maps (especially the 1.4 GHz one) this discrepancy becomes evident when smoothing the images to lower resolutions such as 1° (**Figure 4.2**). This is because at 1° resolution, the low values are close to the edge of the coverage of the beam and therefore they will also be included in the convolution. As the bad values run from one side to another in the GB6 survey image (**Figure 4.1**), inclusion of them in the convolution with the beam is unavoidable at any resolution.

4.3 Map Units Conversion

Before determining the flux density of the source and plotting its spectrum, all the maps were converted to the same final units of Jy/pixel as they allow easy conversion to an integrated flux density in Jy (Section 5.1). Column 5 of **Table 4.1** presents the units each dataset was initially in; these were Jy/beam, MJy/sr or mK (CMB). Below we describe the mathematical procedure followed to convert from one of the aforementioned units to Jy/pixel. The final conversion factor derived, was scripted in IDL and implemented in the electronic code used for map alignment and smoothing. It must be pointed out that the SFD Dustmap was in units of E(B-V) Reddening



Figure 4.2: The Effelsberg survey images smoothed to 1 degree resolution. The upper right corner is missing because of the incomplete beam coverage. See also **Figure 4.4**

that could not be converted to Jy/pixel.

4.3.1 [Jy/beam] to [Jy/pixel]

To convert from Jy/beam to Jy/pixel, we simply need to multiply by the factor [pixel area/beam area] i.e.,

$$\frac{Jy}{pixel} = \frac{Jy}{beam} \times \frac{A_{pixel}}{A_{beam}} \tag{4.7}$$

where A_{pix} and A_{beam} are the pixel area and beam area, respectively.

Hence we just need to determine the pixel and beam areas. We assume that our beam is a 2D Gaussian peaking at 1 and hence the FWHM along the major axis, θ_{maj} and the minor axis, θ_{min} , will be such that

$$(FWHM)^2 = \theta_{min} \times \theta_{maj} \tag{4.8}$$

where the beam major and minor axis are in units of degrees.

The relation between the FWHM and the standard deviation of the Gaussian, σ will be:

$$FWHM = 2\sqrt{2ln2}\sigma\tag{4.9}$$

and for the standard deviation of the major and minor axis, σ_{maj} and σ_{min} , it will be:

$$FWHM = 2\sqrt{2ln2}\sigma_{min} \tag{4.10}$$

$$FWHM = 2\sqrt{2ln2}\sigma_{maj} \tag{4.11}$$

respectively.

Combining Equation 4.8 with Equation 4.10 and Equation 4.11, we have that

$$\sigma_{maj} = \frac{1}{2\sqrt{2ln2}}\theta_{maj} \tag{4.12}$$

and

$$\sigma_{min} = \frac{1}{2\sqrt{2ln2}}\theta_{min}.$$
(4.13)

The area under a 1D Gaussian is given by:

$$A = \int_{-\infty}^{\infty} \exp(-(x - x_0)^2 / 2\sigma^2 dx) = \sqrt{2\pi\sigma}.$$
 (4.14)

Similarly, for a 2D Gaussian with σ_{min} and σ_{maj} as defined previously,

$$A = \int_{-\infty}^{\infty} \int_{-\infty}^{\infty} \exp\left(\frac{-(x-x_0)^2}{2\sigma_{maj}^2} + \frac{-(y-y_0)^2}{2\sigma_{min}^2}\right) dxdy$$
(4.15)

and based on Equation 4.14, the above integral will be equal to:

$$\sqrt{2\pi}\sigma_{min}\sqrt{2\pi}\sigma_{maj} \tag{4.16}$$

or,

$$A = 2\pi\sigma_{min}\sigma_{maj}.\tag{4.17}$$

Combining Equation 4.17 with Equation 4.12 and Equation 4.13 we get:

$$A = 2\pi \frac{1}{2\sqrt{2ln2}} \theta_{maj} \frac{1}{2\sqrt{2ln2}} \theta_{min} \tag{4.18}$$

or,

$$A = \frac{\pi}{4ln2} \theta_{maj} \theta_{min}. \tag{4.19}$$

Therefore, since our beam is assumed to be a 2D Gaussian, then the beam area will just be:

$$\Omega_m = \frac{\pi}{4ln2} \theta_{maj} \theta_{min}. \tag{4.20}$$

Values for the major and minor axis of the beam are usually given in the header of the FITS file of each image, however in a few cases this was not the case and the $(resolution)^2$ was used instead of the product $\theta_{maj}\theta_{min}$. The pixel area is also included in the header and it is equivalent to the product of the values of the keywords CDELT2 and CDELT1 *i.e.* $CDELT = CDELT1 \times CDELT2$. All the values in the header are given in degrees so we have to convert to steradians first:

$$\theta_{pix}(sr) = CDELT^2(deg^2) \times (\frac{\pi}{180}^2).$$
(4.21)

Putting together, the brightness I_{ν} (Equation 1.2) will be:

$$I\left(\frac{Jy}{pixel}\right) = \frac{S}{\Omega}\left(\frac{Jy}{pixel}\right) = \frac{S}{\Omega}\left(\frac{Jy}{beam}\right)\frac{CDELT(sr)}{1.134 \times \theta_{min}\theta_{maj}}$$
(4.22)

where we have substituted the term $\frac{\pi}{4ln^2}$ with its numerical value 1.134.

4.3.2 [MJy/sr] to [Jy/pixel]

Similarly to the method outline in Section 4.3.1, we have to multiply by the factor [*pixel area/steradian*] *i.e.*,

$$\frac{Jy}{pixel} = \frac{Jy}{sr} \times \frac{A_{pixel}}{sr} \tag{4.23}$$

where the pixel area is defined as in Section 4.3.1. Hence:

$$I_{\nu} = \frac{S}{\Omega} \left(\frac{Jy}{pixel} \right) = \frac{S}{\Omega} \left(\frac{Jy}{sr} \right) \frac{\left(CDELT \times 60\right)^2}{\left(\frac{180}{\pi}\right)^2 \times 3600} \times 10^6$$
(4.24)

where the extra factor of 10^6 is because the units are in MJy not Jy.

4.3.3 [mK] to [Jy/pixel]

Re-arranging Equation 1.3 yields

$$T = \frac{\lambda^2}{2 \times 1.381 \times 10^{-23}} \frac{S}{\Omega}$$
(4.25)

where we have substituted the value for the Boltzmann constant.

The above equation can be rewritten as:

$$T(K) = \frac{\lambda^2}{2 \times 1381} \frac{S}{\Omega} \left(\frac{Jy}{sr}\right)$$
(4.26)

where S is in units of Jy, Ω is in units of sr, T is in K and λ is in cm.

We can then use the conversion factor we derived in the previous section to switch to units of [Jy/pixel], *i.e.*

$$T(mK) = \left(\frac{\lambda(m)^2}{2 \times 1381}\right) \frac{\left(\frac{180}{\pi}\right)^2}{\left(CDELT\right)^2} \frac{S}{\Omega} \left(\frac{Jy}{pixel}\right) \times 10^{-3}$$
(4.27)

where we have added an extra factor of 10^{-3} to convert to mK.

For the WMAP surveys, the absolute temperature T is not in units of brightness temperature, K, but in "thermodynamic" units *i.e.* relative to a black-body at the CMB temperature of 2.725 K. To change to Kelvin we need to use the Planck Correction factor (Finkbeiner et al., 1999), *i.e.*

$$T_B = \frac{T_{therm}}{PlanckCorr} \tag{4.28}$$

where T_B denotes the brightness temperature in Kelvin and T_{therm} is the thermodynamic temperature and PlanckCorr denotes the Planck Correction. The PlankCorr can be shown to be (FDS99):

$$\frac{e^x - 1}{x^2 e^x}.\tag{4.29}$$

The Planck correction factor is essentially the derivative of the CMB black-body curve.

4.4 Final Smoothed Maps

A compilation of all the ancillary data of W40, converted to the same units and smoothed to a resolution of 5 ', 9.4 ' and 1 ° are shown in **Figures 4.3**, **4.4** and **4.5** respectively. The SFD Dustmap is only included for completeness and should be ignored in the aforementioned figures as it is not in the correct units. The Parkes survey is excluded from all figures as it does not contain any data of W40.

Because a map can only be smoothed to a lower resolution (the opposite would mean deconvolving the map), maps at a resolution that did not exceed the smoothing resolution (see **Table 4.1** for the initial resolution of each map) are not included in the aforementioned figures. Of course maps that were not smoothed to the same resolution as the other images, were not used in any photometric measurements either (Sections 5.1.1 and 5.1.2).

At 5 arcmin resolution, extended emission was observed only in the IRAS surveys while in the rest of the maps, W40 looked largely like a point source. This extended emission was mostly visible at 3000 and 25000 GHz. The images produced by the CBI and the Effelsberg 11 cm telescope did not differ from those seen in **Figure 4.1**; this was to be expected as the resolution of the surveys did not change radically. In general, a relative agreement was observed between the datasets as far as the morphology of the source was concerned; two exceptions were the IRAS 100 μm and the 12 μm survey, in which W40 seemed to deviate significantly from the spherical shape.

At 9.4 arcmin resolution, the elongation observed at 5 arcmin was still observed. The agreement with the data from the SFD 100 μm map indicated that the elongation is real hence we concluded that W40 presents extended emission in the South-East direction possibly attributing to the effects of an envelope of dust.which surrounds W40. The beam size did not increase greatly from the previous resolution hence no significant changes were observed comparatively.

Lastly, at 1 degree resolution, the larger beam size has allowed more flux to enter from the background (Galactic plane). All images shown, presented a high degree of blending, with the exception of the CBI. This is attributed to the fact that the CBI image was produced by an interferometer which is more effective in resolving out background emission than a radio telescope. The effects of the Galactic plane are clearly seen at the slope formed across the image to the left side of **Figure 4.5**. No conclusions could be made from the Stockert and Haslam survey images as the background emission is so strong that W40 is barely visible.



Figure 4.3: Ancillary data smoothed to a resolution of 5 '. The original map for GB6 is used as smoothing was not possible. Data from the Effelsberg 21 cm, Parkes 6 cm, WMAP, and SFD surveys have been excluded (see text). All maps are scaled in the same way and are shown in celestial coordinates (Right Ascension, Declination). The size of the region was $3^{\circ} \times 1^{\circ}$.



Figure 4.4: Ancillary data smoothed to a resolution of 9.4 '. The original map of GB6 was used. The SFD 100 μm was included. It is reminded that it was not possible to convert the SFD Dustmap in units of Jy/pixel and hence it was excluded. All maps are scaled in the same way and are shown in celestial coordinates (Right Ascension, Declination). The size of the region is $3^{\circ} \times 1^{\circ}$.4.



Figure 4.5: Ancillary data smoothed to a resolution of 1 degree (60 '). The original map of GB6 was used. The WMAP, Stockert and Haslam maps were included. All maps are scaled in the same way and are shown in celestial coordinates (Right Ascension, Declination). The size of the region is $4^{\circ}.2 \times 0^{\circ}.5$.

Chapter 5

Spectral Energy Distributions of W40

In this chapter we describe the procedure to produce the spectrum of W40 (or its spectral energy distribution; SED). The photometry of W40 was derived using two standard photometry methods: 2-D Gaussian fitting (Section 5.1.1) and Aperture Photometry (Section 5.1.2). Limitations in extracting the photometry of W40 from specific datasets were discussed in Section 5.1.3. The assumptions in the error analysis and the final values of the errors are presented in Section 5.1.4.

Simple parametric models were fitted to the data to determine which type of emission best described the observations (Section 5.2.1) and to investigate possible excess emission at ~ 30 GHz. In particular, a free-free model was used for the low frequency data while a thermal dust law was employed at higher frequencies. The models were fitted to data smoothed to 5, 9.4 (Section 5.2.2) and 60 arcmin (Section 5.2.3) resolution. The differences between aperture photometry and Gaussian fitting were recorded for fixed values of the spectral index and emissivity. Special cases, where the the two indices were allowed to remain non-fixed in turn, were also examined.

Hence, an investigation of the emission of W40 at ~ 30 GHz was attempted by comparing the observed values with the theoretical predictions of the model (Section 5.3). Finally in Section 5.4 results related to other properties of W40 are presented and discussed.

5.1 Photometry Methods

The term "photometry" refers to any process that involves the extraction of the integrated flux density of a source (measured in Jy) from an image (in which the image will typically be in brightness units *e.g.* Jy/pixel, Jy/beam, MJy/sr or K). This can be done either by fitting a model and extracting information from the best-fit parameters (Gaussian Fitting) or by adding up the pixel values on the source and subtracting the values in the background (aperture photometry). It becomes clear, therefore, that the key elements in performing successful photometric measurements are to determine as precisely as possible the apparent brightness of the source and the background contribution. The latter is usually light from the Interstellar Medium (ISM), other galaxies or nearby stars, scattered sunlight or even terrestrial sources.

5.1.1 Gaussian Fitting Photometry

In Gaussian fitting, a 2-D Gaussian distribution such as the one defined by Equation 4.5 is fitted to the image data. Then the integrated flux density of the source will be equal to the area under the best-fit model which is going to be another Gaussian.

Fitting a Gaussian model seemed to be a logical conclusion for a variety of reasons. The beam is a roughly Gaussian distribution resulting in images that are composed of a very bright source in the middle (peak), and a significantly fainter background (tails). Furthermore, this photometry method is more suitable for extended sources the size of which is not exactly known. In that sense, W40 is an ideal candidate.

In practice, the Gaussian function used was equivalent to the following equation:

$$A[0] + A[1] \exp\left(\frac{-u}{2}\right),\tag{5.1}$$

where

$$u = \left(\left(x - \frac{A[4]}{A[2]} \right)^2 + \left(y - \frac{A[5]}{A[3]} \right)^2 \right).$$
 (5.2)

However, the above equation is not exactly right since x and y can be rotated by an angle θ (A[6]). For the rest of the symbols: A[0] is the baseline level, A[1] is the amplitude of the Gaussian, A[2] and A[3] the standard deviation of the Gaussian model we use to fit in the x, y axis respectively, while A[4] and A[5] are the peak centroids in x, y.

The IDL routine "MPFIT2DPEAK" was employed for the fitting. This function uses a non-linear least-squares fit method to find the best-fit to the data ¹. At the output, it returns an array of best-fit values stored under the same variable names that were stated above (*e.g.* A[2] is the standard deviation of the best-fit distribution in the x axis).

To extract the integrated flux density the best-fit parameters were converted to units of Jy/pixel. This allows the integrated flux to be calculated in Jy. Hence, reversing the procedure described in Section 4.3.3:

$$S = A[1] \ 2\pi \ A[2] \ A[3] \tag{5.3}$$

where $\sigma^2 = A[2]A[3]$ as yielded by Equation 4.17.

The sky subtraction was accomplished indirectly by fitting over a square of size equivalent to 4 times the FWHM of the beam. This size was big enough to include the entire Gaussian distribution. Practically, this meant that the box size was such that it encompassed all of the source in the image, most of the visible flux and a small amount of the background. It is should be pointed out that several sizes were tested before finally settling to the value stated above. In general, it was found that the results were not sensitive to the alteration of the box size.

For completeness, another method of photometry called PSF photometry, is briefly discussed below. A subdivision of Gaussian fitting, PSF photometry assumes that the PSF of the instrument can be described by a model (Gaussian or Moffat)

 $^{^{1}}http://www.physics.wisc.edu/\sim craigm/idl/down/mpfit2dpeak.pro$
and fits it to the image data to extract the flux density in a manner similar to that described in this Section. As this method works better in cases of point sources of known size, it was decided not to use it.

5.1.2 Aperture Photometry

In aperture photometry, measurement of the flux density is achieved in three steps: firstly, measurement of the observed flux density by summing the pixel counts inside a circular area centered on the source (the aperture), and secondly, measurement of the sky's contribution by adding the pixel values inside an annular region (*i.e.* a ring) defined around the aperture. Finally, subtraction between the two, to find the source's flux density which will be free from contamination from the background.

In mathematical terms (Costa & Loyola, 1992) the observed flux density from the star can be represented as:

$$F_T = \sum_{i}^{N} I_i \tag{5.4}$$

where I_i is the intensity at a pixel I and N is the total number of pixels in the aperture.

Then the true flux density of the object, without the contribution from the sky will be:

$$F = \sum_{i}^{N} I_i - NB \tag{5.5}$$

where B is the estimated background level and the rest of the symbols are as defined above.

In principle, one first needs to move the source to the centre of the image. However, centroiding of the source has already been performed by aligning the ancillary maps with the 31 GHz map (Section 4.2.1). The latter was centered in DIFMAP (Section 3.1).

Determining the aperture size proved to be a challenging task. Since it was found that the total flux increased linearly with the aperture radius (also verified by Wells et al. (1994)) making the aperture too large would have led in overestimation of the noise levels as well as the background errors in the image. On the other hand, a small aperture size could have led to underestimation of the true flux density of W40. Howell (1989) states that the maximum signal-to-noise (S/N) ratio of the measurement is reached at an intermediate radius while other authors (Wells et al., 1994; Costa & Loyola, 1992) assign a range of values from 1.5 to 5 times the FWHM, depending on the situation.

The value finally adopted was twice the FWHM. This choice was based on similar criteria as when determining the size of the box used in Gaussian fitting photometry *i.e.* it has to encompass all the visible flux and a small portion of the background. It should be noted that for a diffraction-limited telescope, where the resolution is given by the Rayleigh criterion, one would expect the aperture size to change for different frequencies. We decided against this course of action for two reasons: firstly, low level extended emission will not be seen in some cases due to the varying PSF. Secondly, the aperture corrections that will need to be applied for the beam slightly spilling over the edge of the aperture will over-complicate the analysis. Therefore, to account for this inconsistency we chose to smooth to the same resolution and use the same size aperture for all frequencies.

Care was taken when deciding for an annulus size and thickness (Wells et al., 1994). Ideally, the annulus size had to be located far enough away that it remained free from the source's flux (similarly to the method of determining the RMS noise in Section 3.2). However, placement of the annulus too far away meant measurements of a background which was not local to the source. At the same time its thickness had to be such that it did not include too much of the sky as that might have led to an overestimation of the noise levels and hence could have reduced the flux density of the source, on the moment of subtraction, beyond its true value (Wells et al., 1994). Based on these considerations, the annulus size was decided to be 1.3 times the aperture radius. This meant that the annulus was placed right outside the aperture.

The Galactic plane affects the photometric measurements, as it lies near W40,

by adding more flux on the side of the annulus which is closest to (see for example, Figure 4.5). As a result, a skewed distribution of intensities is formed and oversubtraction may occur between the noise and the observed flux density. This effect was accounted for by calculating the median value of the pixels in the annulus aperture rather than the mean value; the use of the latter would have been more appropriate if no significant background contribution was observed.

In principle, aperture photometry is a technique that works well for bright sources in a relatively "empty" environment where the aperture and annulus are not expected to include flux from other sources. As discussed above and as it can be seen in the Figures of Section 4.3.1, these criteria were not satisfied by many of the survey images. This point is revisited in Section 5.2.3.

5.1.3 Survey Data Limitations in Photometry

Below we discuss problems that have arisen during the extraction of the flux density from specific survey data and how these were dealt with.

PMN/GB6

The missing data in this survey (Figure 4.1) make it very hard to use in any photometric measurement. Smoothing is not applicable as it is difficult to incorporate all the NaN values (Section 4.2.1). Therefore, alternatives were considered.

To begin with, one could substitute the missing values with the minimum values in the map however this introduces the possibility of underestimating the value of the true flux density of the source because the missing values may be larger in number than the the useful data. Interpolation over the good data was also considered but it was almost immediately abandoned as it is a complicated procedure that requires a small number of pixels to work successfully. This requirement is not fulfilled by the maps we are using (512 \times 512 pixels).

Consequently, the last remaining option was used: fitting to the original data after assigning a zero weight to the NaN values. This method is mostly suitable when dealing with relatively unresolved sources where an elliptical Gaussian is a good model *i.e.*, smoothing returns another Gaussian with the same integrated flux. As such is the case for W40 (Sections 4.2.1 and 5.1.1) the photometric measurements were expected to be at least partially successful.

To perform Gaussian fitting photometry, a standard box size of 50 arcmin/side was used. This box size was defined independently from the beam size as GB6 cannot be smoothed to any different resolution.

It is important to point out that in aperture photometry this method produced unsatisfying results leaving no other option than the complete removal of GB6 from the analysis.

Effelsberg 1.4/2.7 GHz

As explained in Section 4.2.1, the Effelsberg maps contain missing values (**Figure 4.2**. As a result, at lower resolutions (above 40 arcmin), the beam size was large enough to include the missing values and hence affected the photometry of the region even after assigning them a zero weighting .

Fortunately, the effect became prominent only in the 1.4 GHz survey (since the missing values were closer to the source) and only for Gaussian fitting photometry, possibly because of the size of the box used for the fitting. This might explain why the aforementioned effect did not have a significant contribution in aperture photometry; the aperture size was not large enough to include the NaNs. For the reasons stated, it was decided to remove the data of the Effelsberg 1.4 GHz from the Gaussian fitting photometry analysis. For the 2.7 GHz map all the NaN values were set to zero. This produced a reduction in the flux as measured by both of the photometry methods for all resolutions.

Haslam 408 MHz & Stockert 21 cm

The 408 MHz map (Section 4.1) was initially downloaded to add an additional flux density measurement at lower frequencies so that the fit would be better constrained. However, the contribution from synchrotron emission, that dominates at those frequencies (Section 1.3.2), and the Galactic plane blended with the emission from the

source made it hard for an accurate estimate of the flux density to be measured by both photometries. Hence it was removed from the final analysis.

Similarly for the 21 cm Stockert survey, the contribution from the Galactic plane resulted in overestimation of the flux density and hence the data could not be used.

5.1.4 Error Consideration

In general, the basic source of error is the background noise. As new telescopes can operate at increasingly high resolutions, it becomes more likely to include in the observations of the object of interest other sources as well. These sources affect the target's brightness and contaminate the signal. Hence, to determine the errors we assumed that the total error in the flux density is equivalent, to a first approximation, to the average fluctuations in the map (*i.e.* the standard deviation).

The fluctuations in the image were measured by first evaluating the local noise in areas of the map that are sufficiently free from signal and then averaging the value found over the whole image using the STDDEV function in IDL. We used boxes of 32 arcmin on a side and placed them at the two top corners of the map.

It is noted that despite this method being similar to the way the RMS noise was determined in Section 3.2 here only the two top corners were used instead of the ones at the bottom of the image or, even, all four. Such an action was performed to avoid inclusion of the flux from the Galactic plane through the bottom corners (as the Galactic plane lies in the South-East (lower left) direction of W40; Figure 4.5). This is mostly the case for total-power maps like the Effelsberg surveys. Since the Galactic plane is out of the CBI's u-v range and outside its primary beam, leakage by it is not a consideration thus allowing the use of all four corners to estimate the noise.

At low resolutions (≈ 60 arcmin), however, a noise estimation through the method described above would not have been as successful since three out of the four corners in the image were not free of signal (see **Figure 4.5** and the relevant discussion in the same section for more details) *i.e.* the number of independent samples was too low to make a definitive measurement. It follows that the main

measurements would be biased. To solve this problem, strips across the 3 sides of the map were used instead of boxes. The thickness of the strips was equal to the side of the box used when smoothing to high resolutions. To avoid the contribution from the Galactic plane the entire side at the bottom of the map was not included.

An additional calibration error was added in the final error; its value was instrumentbased and thus varied with each dataset. However, in almost all cases, the absolute calibration error was assumed to be 10%. This value is within limits for most of the radio surveys (*e.g.* GB6) and F-IR surveys (*e.g.* IRAS). It is worth pointing out that the same error was used for WMAP data as well, where the error is reportedly less than 1% (Bennett et al., 2003). This was because such a low value for the error caused a discrepancy in the best-line fit between the free-free and the thermal dust model (Section 5.2.1). The only exception were the CBI data where the error was assumed to be 5% (as mentioned in Section 2.2.1).

Using a standard propagation of errors method based on Equation 5.3, we have :

$$\sigma_S = S \sqrt{\left(\frac{\sigma_{A[1]}}{A[1]}\right)^2 + \left(\frac{\sigma_{A[2]}}{A[2]}\right)^2 \left(\frac{\sigma_{A[3]}}{A[3]}\right)^2} \tag{5.6}$$

where σ_S is the error in the flux density, $\sigma_{A[1]}$ are the fluctuations in the image and $\sigma_{A[2]}$, $\sigma_{A[3]}$ are the errors in the standard deviation of the best fit model. (Section 5.1.1).

Finally, the absolute calibration error, σ_{calib} was added in quadrature, *i.e.*

$$\sigma_{final} = \sqrt{\sigma_S^2 + \sigma_{calib}^2} \tag{5.7}$$

to give the final error in the flux density, σ_{final} .

5.2 Spectral Energy Distributions

5.2.1 Theoretical Models and Considerations

In the spectral energy distributions (SEDs) of W40 (see Figures in Section 5.2.2 and Section 5.2.3) the low frequency data points were fitted to an optically thin free-free power law described by the equation:

$$S = S_0 \left(\frac{\nu}{31GHz}\right)^{\alpha} \tag{5.8}$$

where the symbols are defined in the same way as in Section 1.2. Fitting was performed by the IDL function "MPFITFUN". Whenever the free-free index was kept fixed, it was kept at the theoretical value of -0.12 (Section 1.3.1). The validity of this choice was accessed by crosschecking with data by the Parkes telescope at 15 arcmin resolution. Additionally, for fitting at higher resolutions (5, 9.4 arcmin) a computer-generated value was also used *i.e.* the α was left as a free parameter in the fit (see Section 5.2.2 for more details). The latter was calculated by MPFITFUN and was returned to the user as one of the best-fit model parameters.

The value of the index used, corresponds to an electron temperature, $T_e \approx 8000 \ K$ (Figure 1.2). Because of the variation of the spectral index with T_e (Figure 1.2; Equation 1.12) it is important to verify whether the electron temperature of W40 is indeed near 8000 K. This would make our choice of α valid.

For this purpose, the following formula, in units of K, was used (Alves et al., 2010):

$$T_c = \frac{0.08235 \times \alpha(T_e)}{1.4^{2.1} \times 1920} \times 1.08 \times \int T_L dv \ T_e^{1.15}$$
(5.9)

where T_L is the line temperature integrated over the velocity, T_c is (in this case) the free-free continuum temperature as we are assuming that W40 is an HII region dominated by free-free emission, T_e is the electron temperature and the $\alpha(T_e)$ is a slowly varying function of frequency (Alves et al., 2010). Values for the variables were taken from the Parkes HI All-Sky Survey (HIPASS) survey (Staveley-Smith et al., 1996). Re-arranging and evaluating the product of the constants and the integral (Marta Alves; private communication) the value derived for the electron temperature was $T_e = 9100 \pm 300$ K.

The above value of the electron temperature is in agreement (within the errors) with values from other surveys. In particular, Wink et al. (1983) finds $T_e = 8000 \pm 1400$ while Quireza et al. (2006) finds $T_e = 8450 \pm 70$. Hence the value of the spectral index that was used in the free-free model was consistent with the electron temperature of W40.

Data at higher frequencies (>100 GHz) were fitted using a simple thermal dust model described by the equation of the intensity of a modified B-B (or "greybody") (Dupac et al., 2003):

$$S = \epsilon_0 B_{\nu}(\nu, T_{dust}) \left(\frac{\nu}{\nu_0}\right)^{\beta+2}$$
(5.10)

where $B_{\nu}(\nu, T_{dust})$ is the Planck function, T_{dust} is the dust temperature and β is the dust emissivity index. The use of the above equation is appropriate here as the presence of vibrational dust is more noticeable at sub-mm/IR wavelengths (Section 1.2 & Section 1.3.3).

The emissivity index was kept fixed for fitting at higher resolutions (5, 9.4 arcmin) at the value of $\beta = 1.7$ as it is typically the case for HII regions (Dupac et al., 2003). At lower resolutions (60 arcmin) MPFITFUN was allowed to generate a value of β based on the best-fit model (see Section 5.2.3).

For both the free-free and the thermal dust model, the use of indices was based purely on the data enabled by the evaluation of the difference between theory and observations. This allowed the investigation of the possibility of excess emission at any frequency, though the efforts were concentrated at ~ 30 GHz (Section 5.3). For the same reason the data close to those frequencies, namely the CBI 31 GHz and the 33 GHz WMAP K-band, were not included in the fit..



Figure 5.1: Integrated flux density spectrum of W40 at 5 arcmin resolution derived using Gaussian fitting photometry. The filled circles represent data fitted with a simple powerlaw with a fixed spectral index and a modified B-B model with fixed emissivity at higher frequencies. The dotted line represents the line of best-fit. The error bars at 31 GHz are not visible because the symbol used is bigger and hides them.

5.2.2 Model Fitting at 5 & 9.4 Arcminute Resolution

The spectral energy distributions (SEDs) of W40 at 5 and 9.4 arcmin resolutions as derived by Gaussian fitting photometry and aperture photometry are shown in Figures 5.1 through 5.4. The CBI is not included in the fit and is only shown for reference. The spectral index was kept fixed at -0.12 (Section 5.2.1).

In general both photometry methods produced similar results at these resolutions. At 5 arcmin, the values of the integrated flux densities were nearly identical for the low frequency data while at high frequencies small-scale deviations were consistently recorded. It is important to note that the observational value at 31 GHz agreed well with the one predicted by the best fit model. A quantification of this result is presented in Section 5.3.

These deviations became more prominent at the 9.4 arcmin resolution possibly because of the larger beam size which resulted in the acceptance of more flux. In



Figure 5.2: Integrated flux density spectrum of W40 at 5 arcmin resolution derived using aperture photometry. Both the free-free spectral index and the dust emissivity were kept fixed. See **Figure 5.1** for more details on the properties of the plot.



Figure 5.3: Integrated flux density spectrum of W40 at 9.4 arcmin resolution derived using Gaussian fitting photometry. Both the free-free spectral index and the dust emissivity were kept fixed. See **Figure 5.1** for more details on the properties of the plot.



Figure 5.4: Integrated flux density spectrum of W40 at 9.4 arcmin resolution derived using aperture photometry. Both the free-free spectral index and the dust emissivity were kept fixed. See **Figure 5.1** for more details on the properties of the plot.

particular, for the IRAS 100 μm and 60 μm the transitions between Gaussian fitting and aperture photometry were associated with an increase in the flux density of up to ~ 50 %. The opposite was true at low frequencies where the value at 2.76 GHz was returned higher from Gaussian fitting rather than aperture photometry. This was speculated to be because of the missing values in the Effelsberg 11 cm map (Section 5.1.3 and Section 4.2.1) that were expected to affect the aperture photometry method to a higher degree. Deviations recorded in this case however were only about 10%.

The values of the model χ^2 derived are shown in **Table 5.2.2**. The line of bestfit was highly consistent with the datapoints (the χ^2 values are all near 1; **Table 5.2.2**) especially at higher frequencies in both aperture photometry and Gaussian fitting and for both 5 and 9.4 arcmin. This suggested that the simple dust model used to describe W40 at high frequencies was a very good approximation at high resolutions.

At lower frequencies, the fit was still consistent but within the error bars for

Resolution	Gauss. Fitting	Aper. Photom.
5 arcmin	1.13	NaN
9.4 arcmin	0.98	1.002
60 arcmin	1.71	5.4

Table 5.1: The χ^2 values of the fixed spectral index and fixed dust emissivity model at all three resolutions that the ancillary maps were smoothed.

both photometry methods. At 5 arcmin resolution, the absence of data in aperture photometry caused a perfect fit as the curve had only one point to pass through. This was reflected in the value of χ^2 which was calculated to be NaN (representing infinity) due to the absence of any degree-of-freedom (d.o.f.; we remind the index is also fixed). This implied that no robust conclusion can be made from aperture photometry at 5 arcmin resolution. At 9.4 arcmin, the situation is slightly more complicated as there is inclusion of the Effelsberg 21 cm map. The flux densities derived through Gaussian fitting (**Figure 5.3**) described marginally better the theoretical model than in aperture photometry, where the data are barely consistent within the errors (**Figure 5.4**). This conclusion can be verified by the values of χ^2 that are lower for the Gaussian fitting photometry at both resolutions. It should be noted, however, that exclusion of the GB6 data from any SED based on aperture photometries may have contributed to this situation.

Interestingly, at 9.4 arcmin, the CBI starts to become inconsistent mainly in Gaussian fitting (**Figure 5.3**). It is speculated that this is due to a small amount of "flux loss" *i.e.* some of the emission at $\gtrsim 5$ arcmin scale is resolved out.

Non-fixed Free-Free Spectral Index

Fitting a model with a non-fixed free-free index (Section 5.2.1) was also tested. The resultant SEDs using Gaussian fitting photometry are shown in Figures 5.5 and 5.6 for 5 arcmin and 9.4 arcmin resolution respectively. The results from aperture photometry were similar and are thus omitted.

A discrepancy was observed at 5 arcmin with the free-free spectral index becoming positive and gaining a rising tendency compared to the theoretical free-free index used previously. The value of the free-free index at 9.4 arcmin was found to



Figure 5.5: Integrated flux density spectrum of W40 at 5 arcmin resolution derived using Gaussian fitting. At low frequencies, a power-law with a non-fixed index was used while at high frequencies we used a fixed dust emissivity. The properties of the plot are similar to those of **Figure 5.1**.



Figure 5.6: Integrated flux density spectrum of W40 at 9.4 arcmin resolution derived using Gaussian fitting. At low frequencies, a power-law with a non-fixed index was used while at high frequencies we used a fixed dust emissivity. See **Figure 5.5** for more information on the properties of the plot.

be: 0.16 ± 0.25 . The same cannot be said for the SED at 9.4 arcmin, where it is largely similar to the theoretical curve; the value of the spectral index returned by the best-fit model is -0.09 ± 0.12 which is very close to the assumed theoretical value of -0.12 ± 0.2 used previously (Section 5.2.1). Therefore, since the non-fixed index at 5 arcmin is actually consistent with the fixed index but it is also associated with a large error, it was decided it was best not to use it over the fixed value.

We are led to believe that the absence of the 1.4 GHz Effelsberg map between the SEDs of the two resolutions caused the difference between the fits at 5 and 9.4 arcmin. This argument was based on two reasons: by visually examining the plots, it was obvious that if the value of the flux density at 1.4 GHz were to be excluded from the 9.4 arcmin fit (Figure 5.6) then a similar curve as the one seen at 5 arcmin would have been produced. This is supported by the values of the flux density that were derived at 2.76 GHz; at 5 arcmin resolution, $S_{2.76 \ GHz} = 37.2 \pm 3.7$ Jy while at 9.4 arcmin resolution $S_{2.76 \ GHz} = 37.2 \pm 3.7$ where $S_{2.76 \ GHz}$ is the value of the flux density at 2.76 GHz. Furthermore, the value of GB6 remained the same regardless of the resolution as the original data were used. Consequently, since the flux densities at 2.76 and 4.85 GHz remain the same between the two smoothing resolutions, the addition of a value at 1.4 GHz will be responsible for any change observed.

If such a change in the fit could be observed with just the addition of one datapoint, then this should be seen as evidence of the need for more reliable data at lower frequencies in order to constrain a fit with a non-fixed index in a better way.

Thus, it can be concluded that an optically thin (Section 5.4.1) free-free model of *fixed* spectral index equivalent to -0.12 should be used hereafter in this thesis to describe the flux density of W40 at low frequencies up to ~ 15 GHz.

5.2.3 Model Fitting at 1 Degree

Figures 5.7 and 5.8 present the SED of W40 at 60 arcmin resolution. In Figure 5.7, Gaussian fitting photometry was used while in Figure 5.8 the data point values were found by means of aperture photometry. Both the free-free and thermal dust



Figure 5.7: Integrated flux density spectrum of W40 at 60 arcmin resolution derived using Gaussian fitting photometry. Both the spectral index of the free-free model and the thermal dust emissivity were held constant. See **Figure 5.1** for more details on the properties of the plot.

index were kept fixed.

The general picture remained the same albeit the differences between the two methods described in the previous section became more prominent. The flux densities were vastly greater as determined by aperture photometry. This is an important result as at 60 arcmin the beam size is large enough to allow more flux from both the source and the background. Since the aperture size depends on the beam size, the flux density inside the aperture will also increase. It becomes clear then that the difference between the two methods of photometry depends greatly on the contribution from the background and especially the fact that aperture photometry is affected the more by it. The differences in filtering at 1° scales could also account for part of the discrepancy observed *e.g.* the Effelsberg data are high-pass filtered.

The result from the Gaussian fitting was again highly consistent with the data at intermediate frequencies and high frequencies, just as seen at higher resolutions. It is interesting that at 94 GHz a discrepancy was observed between the theoretical model and the observations, implying that the dust model used at this frequency range may have been too simple. Even though such a model is usually expected to work in small sources, it would appear that W40 is far more complex as it involves a lot of extended emission; part of it may even be spinning dust. We concluded then that a model with a fixed dust spectral index of $\beta = 1.7$, as the one assumed in the fit, is *not* able to completely describe the thermal dust component of the W40 spectrum. At the same time, dust components of different temperatures and emissivities may be able to account for the spectrum of W40 more accurately; for instance, a cold component (T<15 K) would peak at ~ 1000 GHz and even though it would avoid detection by IRAS it would still affect the spectrum at 94 GHz.

At lower frequencies, the discrepancy observed was only a result of our inability to properly incorporate the GB6 survey data in the SED. The problem was not as significant at higher resolutions where the beam size was smaller. However, at 60 arcmin it certainly became a liability as it lies 2.5σ away from the best-fit curve. Setting this limitation aside, the conclusion stated in the end of Section 5.2.2, that the model is well-suited for W40 at low frequencies, is reinforced even further. This is reflected in the value of χ^2 (Table 5.2.2) which is slightly higher than 1 (owing partly to the 94 GHz point, too).

It is important to note that the observed value of the flux density of the CBI is underestimated because of the flux loss due to the limited u-v coverage of the CBI data at 1° angular scales.

On the contrary, the theoretical model fitted to the aperture photometry values of the flux density resulted in a very high χ^2 and hence does not seem to adequately describe the observational data. There is a possibility that the deviation from the best-fit curve at low frequencies (between 1.4 and 2.76 GHz) implies strong excess emission and perhaps an anomalous component. In fact, a spinning dust model does look like an appropriate fit between ~ 10 GHz and ~ 60 GHz. Given, however, that such an emission excess is not observed in the Gaussian fitting SED, implies that it is most likely related to a variant background (as explained above) rather than the source itself. We conclude that this is an aspect of the project that deserves future



Figure 5.8: Integrated flux density spectrum of W40 at 60 arcmin resolution derived using aperture photometry. Both the spectral index of the free-free model and the thermal dust emissivity were held constant. See **Figure 5.1** for more details on the properties of the plot.

work.

Summing up, in comparing the two methods it became obvious that aperture photometry was more prone to return inaccurate results; a conclusion largely based on the ambiguity involved when selecting the appropriate aperture and annulus size. Simply put, the Gaussian fitting is more robust for high signal-to-noise data such as those described here. The χ^2 values of the Gaussian fitting photometry were also calculated at 5, 9.4 and 60 arcmin resolution. These were largely found to be ~ 1 revealing that this method of photometry was successful and most suitable for use for this type of data. It was, therefore, decided that the search for the excess emission at frequencies around 30 GHz was to be determined using the results from the Gaussian fitting method. This choice was also supported by the fact that the telescope beams are approximately Gaussian (Section 4.2.1, Section 5.1.1) as well as by W40's properties (small size (Section 5.4.1) and high luminosity).

Non-fixed Dust Emissivity

In a similar manner as described in Section 5.2.2, the thermal dust emissivity index was not fixed but rather it was calculated by MPFITFUN (Section 5.2.1). The reduced value for the dust emissivity was $\approx \beta = 0.1 \pm 0.01$. This is a somewhat lower than what is typically found in the ISM ($\beta \sim 1 - 3$). This is likely to be due to the single component dust model we are trying to fit implying that it is too simple; more complicated models may have worked better (for instance a multiple dust component can result in such a value).

Owing to the absence of the WMAP data that work as a link between low and high frequencies (*i.e.* not enough d.o.f. to constrain the fit; Section 5.2.2), calculating the best-fit emissivity index was only possible at 60 arcmin resolution. A new model was fitted to the data as shown in **Figure 5.9**.

At low frequencies, the fixed free-free index model of Section 5.2.2 seemed to adequately describe the situation with the datapoints being consistent with the model within the error bars.

As already mentioned (Section 5.7) at 31 GHz, the value of the flux density

of W40 seems severely underestimated compared to the line of best-fit because of the CBI's property to resolve extended emission. Hence no excess emission could be observed at this frequency. On the contrary, at 33 GHz, a small "bump" was detected; investigation of whether this was a noise fluctuation or some evidence of true excess emission ensued (Section 5.3). At 94 GHz, contribution from the CMBR causes an great increase in the flux density of W40. In particular, as the spectrum of W40 is mostly optically thin free-free emission (Section 5.4.1), at higher frequencies it is expected to be falling; the CMB radiation is so strong at those frequencies that it illuminates W40 and makes its spectrum rise (n.b. the spectrum of CMBR at those frequencies is rising in flux density). This issue was resolved by subtracting the WMAP W-band data with the WMAP ILC, essentially a CMB map. This decreased the initial value of 66 Jy to 50 Jy and made the data point consistent with the model fit. The rise, as seen in **Figure 5.9**, is attributed to the effects of thermal dust that are slowly introduced at this frequency range.

At higher frequencies (>100 GHz) no disagreement was observed between the theoretical model (Figure 5.9) and the observations. Even though it was initially thought that such a model was too simple to account for W40's SED (Section 5.2.3) it would appear that this is not the case.

We conclude that a modified B-B spectrum model with a thermal dust emissivity $\beta = 0.1$ is capable of describing W40 at high frequencies in a relatively sufficient manner.

5.3 Excess Emission

5.3.1 At 31 GHz

The ultimate goal was to quantify whether excess emission is observed in the spectrum of W40 at frequencies around 30 GHz and especially at 31 GHz. This is an ideal observing frequency to detect emission from spinning dust as the only other dominant foreground is free-free emission (Section 1.3.1).



Figure 5.9: The SED of W40 at 60 arcmin resolution derived using Gaussian fitting photometry. The free-free spectral index was kept fixed but the thermal dust emissivity remained free and the best-fit value calculated by the computer was used. See **Figure 5.1** for more details on the properties of the plot.

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To begin with, the predicted value of the flux density at 31 GHz had to be determined. For this purpose, a modified version of the theoretical model used in Section 5.2.3 was employed, where the contribution from the thermal dust is assumed to be negligible. Hence the model reduced to just a simple power law, with a fixed free-free spectral index, such as the one described by Equation 5.8, *i.e.*:

$$S = A[0] \left(\frac{\nu}{31 \ GHz}\right)^{-0.12} \tag{5.11}$$

where A[0] is as defined previously (Section 5.1.1) and the rest of the symbols have their usual meanings. Substituting the appropriate value for the frequency ν , in this case equal to 31 GHz, yielded the predicted value of the flux density at that frequency *i.e.*, $S_{33GHz} = A[0]$.

The error on S_{33GHz} , σ_S , was found by simple propagation of errors of the quantities defined in Equation 5.11. Formally,

$$\sigma_S^2 = \left(\frac{\partial S}{\partial A[0]}\right)^2 \sigma_{A[0]}^2 + \left(\frac{\partial S}{\partial \nu}\right)^2 \sigma_{\nu}^2 \tag{5.12}$$

where $\sigma_{A[0]}$ is the error on A[0] and σ_{ν} is the error on the frequency.

By normalizing to 31 GHz and assuming that $\sigma_{\nu} = 0$, as it is negligible for most telescopes, it was found that $\sigma_S = \sigma_{A[0]}$ at 31 GHz. It should be noted that the magnitude of $\sigma_{A[0]}$ was returned by the fitting function.

To find the excess emission at 31 GHz, S_{excess} , a comparison between the predicted value (derived above) and the observed value (returned by the Gaussian fitting procedure; Section 5.1.1) had to be made. Mathematically,

$$S_{excess} = S_{obs} - S_{31GHz},\tag{5.13}$$

and the error on S_{excess} , σ_{excess} was found by evaluating the following equation:

$$\sigma_{excess} = \sqrt{\sigma_{obs}^2 + \sigma_{31GHz}^2} \tag{5.14}$$

where all the variables are defined as above.

It is important to quote the result for the observational value of the flux density as that will be the integrated flux density of W40 derived by this analysis. At 5 arcmin, $S_{obs} = 27.7 \pm 1.4$ Jy and at 9.4 arcmin, $S_{obs} = 27 \pm 1.3$ Jy as returned by the IDL routine MPFITFUN.

Substituting for all the unknowns, the value found was $S_{excess} \approx -2.2 \pm 2.5$ Jy at 31 GHz at 5 arcmin resolution. An upper limit of 5 with a 2σ confidence level is hence found. This value is consist ant with zero at the 1σ level which implies that no detection of spinning dust has been recorded. A possible interpretation could be related to the environment that W40 resides. As this is a very hot and active region of star formation it is possible that the small dust grains responsible for this type of emission, are unable to withstand the heat and thus are destroyed.

It is possible to tighten the limit we derived above by considering the same calculation at 9.4 arcmin angular scales where the 1.4 GHz point increases the d.o.f. and hence a slightly better fit is acquired. Substituting the appropriate values we get: $S_{excess} \approx -2.8 \pm 2.2$ which corresponds to an upper limit to the excess emission of ≈ 4.4 Jy with a 2σ confidence level.

Lastly, we can use the IRAS 100 μm survey to calculate the expected spinning dust emission levels at 31 GHz. Davies et al. (2006) reports that typical levels for the spinning dust emission are: ~ 10 $\mu K/(MJy/sr)$ of spinning dust at 100 μm and converting to Jy per Jy we get: ~ 4000 Jy of emission at 100 μm for 1 Jy of spinning dust at ~30 GHz.

In our case, the observed value of the flux density at 100 μm , is $S_{3000GHz} = 1154$ Jy and hence based on the aforementioned levels, it is found that for 1154 Jy flux at 100 μm only ~ 0.3 Jy of spinning dust are expected at ~ 30 GHz.

We conclude, therefore, that our inability to observe spinning dust may be attributed to the relatively small amount of 100 μm at ~ 30 GHz.

5.3.2 At 33 GHz

As mentioned (Section 5.2.3), the inability of the CBI to observe extended emission has caused a great flux loss in the 31 GHz data at low resolutions. As a result, conclusions about the emission of W40 at that frequency could not be made at 60 arcmin scales.

However, the possibility that excess emission peaked at a nearby frequency led to suspicions about the nature of the small "bump" observed at 33 GHz (Figure 5.9). To examine this possibility, a method similar to the one described above (Section 5.3.1) was followed, but at a frequency of 33 GHz.

Substituting for all the unknowns, the value found was $S_{excess} \approx 5.9 \pm 4.7$ Jy at 33 GHz. This value is not large enough (less than a 1.5 σ deviation is present) to be regarded as evidence for excess emission and hence it is interpreted as a noise fluctuation.

5.4 Properties of W40

Below we present some of the properties of W40 as derived by fitting the simple models of Section 5.2.1.

5.4.1 Size and Temperature

Gaussian fitting photometry allows the estimation of the size of W40 by making use of the best-fit parameters returned by MPFIT2DPEAK (Section 5.1.1). As previously mentioned, in Gaussian fitting the source is assumed to be a Gaussian peak with the background representing the tails of the distribution. Hence, it can be easily seen that the width of the peak and the position of the centroid in the x- and y-axis will correspond (in physical terms) to the size and position of the source respectively. Specifically, parameters A[2] and A[3] were used to find the size of W40. It should be noted that the values returned by the fitting procedure will include the beam size and hence deconvolution has to be performed before using them in any calculations. Naturally the results were found for images smoothed at 5 arcmin resolution *i.e.* the CBI resolution

In a similar way, the temperature of W40 was determined by MPFITFUN when fitting with the non-fixed dust emissivity model (therefore, at 60 arcmin resolution). In that case, it was one of the best-fit parameters returned.

We are mostly interested in the value of the size at 31 GHz (as suggested by the title of this thesis). At that frequency, W40 was found to be 5.8 ± 0.2 arcmin across. It is worth noting that the size found by CBI agreed to the sizes "viewed" by almost all the other surveys which were included in the fit indicating that at high resolutions, the emission from W40 is dominated by the beam size.

The temperature of W40 was also returned as one of the best-fit parameters when fitting for a non-fixed dust emissivity (Section 5.2.3) and was found to be \approx 57 ± 3.8 K. It would appear that the value derived is quite large making W40 hotter than was initially expected due to heating by O/B stars within the complex.

Ultra-Compact Regions in W40

It is widely accepted that W40 is a large complex of star formation with multiple HII regions (Rodney & Reipurth, 2008) and hence all the components that reside in and around it are of great interest for radio astronomy. Particularly interesting are the ultra-compact (UC) HII regions. These regions have optically thick spectra at low frequencies but as one moves to higher frequencies their spectra fall and they tend to become optically thin. This, in turn, allows us to discern more details in the structure of the components inside a complex, like W40.

We can use the NRAO VLA Sky Survey (NVSS) data at 1.4 GHz to verify whether UC HII regions do exist in W40 and whether their emission spectra contribute to the W40 SED at higher frequencies. It should be noted that, ideally, one would try to detect an UC HII region using a telescope of both high resolution and high frequency. Since such data are not available, the NVSS that has a higher resolution (45 arcsec) than most other telescopes seems to be the best alternative.

Therefore, below, we attempt to constrain the optical depth towards the brightest hotspot in W40 as seen by the NVSS (Figure 5.10). In this way, we can constrain the continuum of the optically thick free-free emission from UC HII.

The apparent size of the hotspot is 1 arcmin (60 arcsec) across but this value encompasses the convolution with the NVSS beam (45 arcsec). Hence by deconvolving,



Figure 5.10: The W40 cloud complex as seen by the NVSS at 1.4 GHz. The resolution is 45 arcsec and the image was downloaded from Skyview (Section 4.1). The size of the image is $4^{\circ} \times 3^{\circ}$ across

the true size was found to be $\sqrt{60^2 - 45^2} = 40$ arcsec.

The observed brightness was also be calculated using Equation 1.1. Then,

$$T_b = \frac{S\lambda^2}{2k\Omega} = 100 \ K \tag{5.15}$$

where the values substituted were S = 0.3 Jy/beam, $\lambda = 0.21$ m and $\Omega = (0'.75)^2 \times \frac{\pi}{180}^2$ steradians.

Substituting in Equation 1.6 an electron temperature of $T_e = 8000$ K and $T_{b,0} = 0$, the value of the optical depth derived is $\tau_{\nu} \approx 0.013$ which is << 1. So at low frequencies (such as the NVSS one), the component found is optically thin. Hence, it could not belong to an UC HII region. Therefore no contribution from UC HII regions was expected at higher frequencies.

Chapter 6

Conclusion

6.1 Results and Summary

This thesis presented microwave observations of the cloud complex W40 with the aim of determining the physical emission processes that contribute to its microwave flux. Located at Galactic coordinates (l,b) = 28.77 + 3.48, W40 is a star forming region that is speculated to be a spinning dust/anomalous emission candidate. We used observations of W40 at 31 GHz taken with the Cosmic Background Imager, an interferometer array located in Chajnantor Observatory, Chile.

Subsequently, the data were calibrated and edited using the software CBICAL. Editing was done both manually, by considering factors such as the number of closure errors produced by a receiver and the phase/amplitude of the calibrator. By carefully editing the data by hand, an improvement of roughly 20% was obtained. Calibration was performed in 3 steps. Firstly, quadrature calibration was applied and the noise and antenna calibration procedures followed. To remove the local correlated signal ("spurious" signal) that was detected on short baselines it was necessary to subtract a lead/trail field.

Using the reduced and calibrated visibilities six images were produced with the imaging software DIFMAP. A uniform weighting was applied to ensure maximum resolution as it was found that CBI is limited by dynamic range not thermal noise. At 31 GHz, W40 appeared as a well-defined, compact (of the order of 6 arcmin), bright

source apart from a hint of extension towards the longer latitudes. An envelope of dust was observed encompassing W40. The noise levels in the images were less than 1% of the total flux. The average peak brightness was found to be 8.6 Jy with a 4.38×3.84 arcmin synthesized beam and an integrated flux densty of 27.7 ± 1.4 Jy. The images produced from each day of observation were combined into one final map using the program UVCON.

A comparison between the 31 GHz data and data at other frequency bands made it obvious that low-level extended ionised gas/dust is located South-East of W40 and connects the cloud complex to the Galactic disk. It became evident that CBI is unable to discern the extended parts of the emission coming from W40. In parallel, the H_{α} dust absorption was found to be very high (~ 50 mag) such that it could not be used to estimate the free-free contribution.

Aligning, smoothing and conversion of the multi-frequency maps allowed the estimation of the flux density of W40 at different frequencies. The flux density was derived using two standard photometry methods: Gaussian fitting and aperture photometry. Both the size of the box used in the former and the size of the aperture used in the latter were twice the size of the FWHM for each particular occasion.

The data were smoothed to three resolutions of 5, 9.4 and 60 arcminutes and plots of the flux density at different frequencies were made to create the SEDs of W40 from ~ 1.4 GHz to ~ 5000 GHz. At the same time, the effects of the Galactic plane (blending with the flux originating from the source, increase in the noise levels in the background) and problematic ancillary maps (missing data, filtering) were the greatest limitations we had to deal with. The errors were assumed to be a combination of the background fluctuations in the image and a standard calibration error of 10% (in most cases). In general, it was found that aperture photometry was returning more biased results than Gaussian fitting mostly due to the effects of the varying background and how easily the aperture was affected.

The last part of the project involved the fitting of different models against the plotted flux densities For the low frequency (1.4-5 GHz) data points it was found that an optically thin free-free power-law was a good fit while at high frequencies

(>100 GHz) a single dust component of a modified blackbody is more appropriate. The optical depth of the brightest hotspot found in the NVSS 1.4 GHz survey was $\tau = 0.013$ indicating that there is no optically thick emission component.

At 9.4 arcmin where two low frequency points (1.4, 2.76 GHz) could be used, a power law with a spectral index $\alpha = 0.16 \pm 0.25$ was found. However, due to the large error in extrapolation, a fixed power law was used throughout at the theoretical value of -0.12 for $T_e \sim 8000$ K. At higher frequencies the model that best described the data was a modified BB spectrum curve with a thermal dust emissivity equal to 0.1 owing mostly to the inclusion of the WMAP data (23-94 GHz). At the same time, the dust temperature T_{dust} , was found to be 57 ± 4 K. This value is typical for diffuse HII regions. We believe this is attributed to the heating from the active star formation that is occurring nearby.

Hence, the calculation of an upper limit on the emission at two different frequencies: 31 and 33 GHz was obtained. In both cases, the flux densities were determined using Gaussian fitting since they were deemed more robust. During the model fitting, the need for more reliable data at lower frequencies (< 15 GHz) and between 100-1000 GHz became evident.

The CBI flux density was found to be completely consistent with the extrapolation at low frequencies. We derived an upper limit on the excess emission of 5.1 with a 2σ confidence level at 31 GHz at 5 arcmin resolution and a tighter limit of ~4.4 at 9.4 arcmin. The contribution from thermal dust was found to be negligible. This result shows that no evidence of excess emission, and therefore spinning dust, is observed from W40. It is speculated that this due to destruction of spinning dust grains from the young stars that are formed in this region.

An upper limit of the excess emission of 5.4 Jy with a 2 σ confidence level at 33 GHz was also determined. This small "bump" was observed at 33 GHz that was simply interpreted as a statistical noise fluctuation. Aperture photometry at 1 degree scales gives up to 50% higher flux density compared to Gaussian fitting. Furthermore, the spectrum at 20-100 GHz is significantly higher than the extrapolation for low (~ 1 GHz) frequencies. It is not certain if this is real of if it is related to the background.

6.2 Future Work

Future work should build upon the results of this thesis and aim into getting a tighter error on the upper limits we have determined. HII regions are interesting for studies of spinning dust since they will typically have a different dust grain distribution from the diffuse ISM away from hot O/B stars. Hence this provides many regions to place tighter limits and to compare it with diffuse ISM environment.

To achieve this, it is essential for more data to be acquired especially at 1-15 GHz (to constrain synchrotron and free-free emission) and above 100 GHz (to constrain thermal dust). PLANCK (30-857 GHz) will definitively be an important dataset for the future.

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