MAXIMA: Observations of CMB Anisotropy

by

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Abstract

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This document describes the Millimeter Anisotropy eXperiment IMaging Array (MAXIMA), a balloon-borne experiment measuring the temperature anisotropy of the Cosmic Microwave Background (CMB) on angular scales of 10' to 5°. MAXIMA data are used to discriminate between cosmological models and to determine cosmological parameters.

MAXIMA maps the CMB using 16 bolometric detectors observing in spectral bands centered at 150 GHz, 230 GHz, and 410 GHz, with 10' resolution at all frequencies. The combined receiver sensitivity to CMB anisotropy is \( \sim 40 \mu K \sqrt{sec} \), the best reported by any CMB experiment. Systematic errors are rejected by using four uncorrelated spatial modulations, multiple independent CMB observations, heavily baffled optics, and strong spectral discrimination. Observation patterns are well cross-linked and optimized for the extraction of cosmological information. Pointing is reconstructed to an accuracy of 1'. Absolute calibration uncertainty of 3-4% is the best achieved by any sub-orbital CMB experiment.

Two MAXIMA flights were launched from the National Scientific Balloon Facility in Palestine Texas in 1998 and 1999. During a total of 8.5 hours of CMB observations, 300 deg\(^2\) of the sky were mapped, with \( \sim 50 \) deg\(^2\) overlap between the two flights. The observed region was selected for low foreground emission and post-flight data analysis confirms that foreground contamination is negligible.

Cosmological results are presented from the 1998 flight, MAXIMA-1, in which 122 deg\(^2\) of sky were mapped over 3 hours. A maximum likelihood map with 3' pix-
elization is obtained from the three most sensitive and best tested detectors. The angular power spectrum derived from this map shows a narrow peak near $\ell = 200$, and is consistent with inflationary Big Bang models. Within these models, cosmological parameters are estimated, including total density $\Omega_{\text{tot}} = 0.9_{-0.16}^{+0.18}$, baryon density $\Omega_b h^2 = 0.033 \pm 0.013$, and power spectrum normalization $C_{10} = 690_{-123}^{+200} \mu K^2$. In combination with recent supernova observations, we obtain additional constrains on the matter density $\Omega_m = 0.32_{-0.11}^{+0.14}$ and the dark energy density $\Omega_\Lambda = 0.65_{-0.16}^{+0.15}$. All parameter estimates are presented at 95% confidence.

The final chapter is a discussion CMB polarization anisotropy, including an overview of MAXIPOL, the polarization sensitive follow-up to MAXIMA. Measurements of CMB polarization are an essential complement to those of temperature anisotropy.

---

Professor George F. Smoot  
Dissertation Committee Chair
To Andrew P. Kitchen.
Preface

MAXIMA is one of the first experiments to map the Cosmic Microwave Background on sub-degree angular scales. My work on the project began in the spring of 1997. I have been involved with preparation and operations for both flights, data analysis, and design and modification of hardware.

This document is intended as a general overview of the MAXIMA experiment, including goals, hardware, data reduction and analysis, and results. Emphasis is placed on experimental techniques. Chapters 4 and 5 are treatments of areas in which I have been particularly active: pointing and responsivity calibration. Later chapters present data analysis, results, and systematic error tests. The final chapter deals with future work on CMB polarization.

Another MAXIMA PhD dissertation, Winant (2003), is in preparation. Though some of the general information overlaps with that in this document, Winant (2003) includes a detailed look at the detector system and optics, that are only summarized here, and treats pointing, observations, and calibration in less detail. In this sense, the two dissertations are complementary.

— Bahman Rabii

Berkeley
September, 2002
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Chapter 1

Motivation and Background

The first section of this chapter is a brief review of the Cosmic Microwave Background and its use as a probe of cosmology, especially inflationary Big Bang models. Theoretical details have been thoroughly explored in published literature (e.g. Hu & Dodelson (2002), Kosowsky (2002)) and are not repeated here. Section 1.2 presents the history of CMB observation. Section 1.3 deals with technical considerations of measuring the temperature anisotropy.

1.1 The Cosmic Microwave Background

The Cosmic Microwave Background (CMB) is nearly uniform blackbody radiation at a temperature of 2.725±0.002 K. It is believed to be of cosmological origin: the heavily redshifted emission from a hot, optically thick period in the early universe. The existence of low temperature background radiation in a Big Bang universe was proposed in 1948 by George Gamow and further explored in 1950 by his colleagues Ralph Alpher and Robert Herman. The CMB was detected by Arno Penzias and Robert Wilson in 1964. Its existence remains one of the strongest pieces of evidence for the Big Bang. Further measurements of the CMB - its spectrum, temperature anisotropy, and polarization - provide information about the structure and evolution of the universe (Smoot & Scott (2000)). MAXIMA and other experiments of the late 1990’s have measured the temperature anisotropy on sub-degree angular scales to test cosmological models and obtain estimates of cosmological parameters.
1.1.1 Fundamental Implications of the CMB

By the early 1990's, two extremely powerful statements could be made of the CMB. First, its spectrum is that of an astoundingly precise blackbody. Second, its temperature anisotropy is extremely small. These two facts have become cornerstones of modern cosmology.

![Figure 1.1: The spectrum of the CMB as measured by the COBE FIRAS instrument. Data are plotted in units of blackbody equivalent temperature. The vertical zero is suppressed to make the error flags visible. This plot is taken from Mather et al. (1994); subsequent improvements in calibration changed the best fit temperature from 2.726 K to 2.725 K (Mather et al. (1999)).](image)

The CMB spectrum, best measured by the FIRAS instrument of the the COBE satellite (Mather et al. (1999)), shows no statistically significant deviations from a thermal spectrum and has a mean Compton $\gamma$-parameter of $<10^{-5}$. Subsequent experiments have measured spectral distortions near galaxy clusters due to the Sunyaev-Zeldovich effect. This spectrum has only been explained by Big Bang cosmology. The universe began in a hot, dense, expanding state. At $\sim$300,000 years after the Big Bang, the universe consisted mostly of baryons, photons, and dark matter, with the baryons and photons in thermal equilibrium. As it expanded, adiabatic cooling allowed electrons and protons to combine, breaking the thermal equilibrium and suddenly increasing the mean free path of photons to greater than the present horizon size of the universe. This transition from thermal equilibrium to photon free
streaming, referred to as recombination or last scattering, occurred at a redshift of \( \sim 1100 \). Since the time of last scattering, CMB photons have cooled with the cosmological redshift to their present temperature of 2.725 K. Because Doppler shifts introduce no spectral distortions, the CMB spectrum today is nearly identical to that of the thermalized universe just before last scattering.

While the almost perfect spectrum of the CMB answers fundamental questions about the evolution of the universe, the homogeneity of the CMB temperature has introduced new mysteries. The largest temperature variation of the CMB is a dipole of about \( 10^{-3} \) K caused by the peculiar velocity of the earth relative to the CMB rest frame. Apart from this dipole, temperature anisotropies are only about one part in \( 10^5 \). These small fluctuations are the seeds needed for gravitational condensation to produce the structures observed in the modern universe. But how were these fluctuations generated? And why is the temperature so homogeneous on large scales? Answers to both of these questions are provided by the concept of inflation (Liddle & Lyth (2000), Garcia-Bellido (2003)).

Large scale homogeneity suggests that the universe at the time of last scattering was at a constant temperature over what is now the observable universe. This region was not in causal contact at the time. According to the dynamics of simplistic Big Bang models, it would never have been in causal contact and could not have come to thermal equilibrium. This “horizon problem” is the most compelling evidence that such models are inadequate. Inflation proposes that a period of rapid acceleration increased the scale factor of the universe by \( \sim 50 \) orders of magnitude during the first \( 10^{-32} \) seconds. Regions were causally connected (and in thermal equilibrium) before they were separated by inflation, and would remain at the same temperature at the time of last scattering and beyond.

Inflation is also a solution to two further mysteries, the “flatness problem” and the “defect problem”. The flatness problem is one of coincidence; the present density of space is very close to the critical value required for a spatially flat universe. However, the growth of the universe through radiation and matter dominated phases causes divergence from flatness by more than 50 orders of magnitude. The present situation is difficult to explain without a mechanism to have forced extraordinary flatness in the early universe. The strong, accelerating expansion of inflation provides such a mechanism. Inflation can also explain the lack of observed topo-
logical defects, such as magnetic monopoles, resulting from phase transitions in the very early universe. It occurred after these transitions and spread the defects apart, reducing their number density to roughly one per present horizon volume.

Inflation provides a mechanism for generating the primordial fluctuations needed to seed later structure formation (Liddle (1995)). This is an important prediction and has been the key to observational exploration of inflationary models. The signature of inflationary structure generation in CMB anisotropy is discussed in the following section.

1.1.2 CMB Temperature Anisotropy

![Graph](image_url)

Figure 1.2: An example of a CMB power spectrum, including measured data points and a model curve calculated for an inflationary Big Bang cosmology with a dark energy component. Data shown are the results of the MAXIMA-I experiment from Lee et al. (2001).

The study of CMB temperature anisotropy is useful in two ways. First,
it can distinguish between cosmological models, especially between those dominated by primordial density fluctuations and those dominated by primordial stress fluctuations. Second, it can be used to determine cosmological parameters, especially the total density of the universe.

Temperature anisotropy is often quantified with an angular power spectrum, as in Figure 1.2, with CMB power in units of $\ell(\ell + 1)C_\ell/(2\pi)$ (or equivalently $\Delta T_\ell^2$) on the y-axis and $\ell$ on the x-axis. Here, $\ell$ is the angular multipole number, inversely proportional to the angular scale, and $C_\ell \equiv \langle a_{lm} \rangle_m$ is the mean spherical harmonic coefficient at a given $\ell$.

**Adiabatic vs Isocurvature Fluctuations**

CMB temperature anisotropy is the result of fluctuations in the baryon-photon fluid at the time of recombination. Generically, there are two orthogonal types of fluctuation: adiabatic (density) and isocurvature (stress) fluctuations. In adiabatic fluctuations, all species (baryons, photons, CDM, etc.) have fixed ratios, but overall density varies spatially. Adiabatic fluctuations directly seed the gravitational growth of structure. For isocurvature fluctuations, the overall energy density is uniform, but there are variations in the number densities of various species. Isocurvature fluctuations causally relax into density fluctuations, indirectly seeding gravitation growth. Density variations grow through gravitational collapse into the structures of the present universe.

The angular power spectrum of the CMB is very different for primarily adiabatic and primarily isocurvature models. In adiabatic models, the power spectrum shows the familiar pattern of several narrow peaks (see the following section and Figure 1.2). In isocurvature models, these peaks are replaced by at most one broad hump.

Only inflationary models produce significant adiabatic fluctuations on all scales (Liddle (1995)), while models with significant numbers of topological defects provide the stresses needed for isocurvature models. For this reason, the concepts of inflation and adiabatic fluctuations have often been paired together as competing with defects and isocurvature fluctuations. Current data (e.g. Chapter 7) are consistent with adiabatic fluctuations, though the possibility remains of a hybrid uni-
verse with subdominant isocurvature fluctuations. This possibility is best explored through the study of CMB polarization (Chapter 9).

**Adiabatic CMB Anisotropy**

In standard inflationary cosmologies, the universe prior to recombination consists primarily of photons, baryons, and collisionless dark matter. Photons and baryons are tightly coupled, while the dark matter is not. The initial spatial spectrum of density fluctuations is close to the scale-free Harrison-Zeldovich spectrum,

\[ |\delta_k|^2 \equiv \left| \frac{\delta \rho}{\rho} \right|^2 = Ak, \quad (1.1) \]

where \( \rho \) is the average density, \( k \) is a Fourier wavenumber, \( \delta \rho \) is the density fluctuation on that scale, and \( A \) is a constant. Under linear gravitational collapse \( |\delta_k|^2 \) grows with the square of the scale factor. This is the state before recombination, outside the sound horizon. As the modes fall within the sound horizon, photon pressure counters the effects of gravity, causing harmonic acoustic oscillations of the photon/baryon fluid. In contrast, the dark matter is coupled to the photons only by gravity via perturbations in the spacetime metric and collapses monotonically. After recombination, baryons are decoupled from photons and quickly couple to the larger density variations of the dark matter and purely gravitational collapse resumes.

Density fluctuations at the time of recombination are imprinted on the CMB; the phase of a given mode is determined by the time between the start of its acoustic oscillations and recombination. This phase sets the observed CMB power at a given scale. The smooth variation of phase with scale leads to evenly spaced “acoustic peaks” in the angular power spectrum.

Density fluctuations cause CMB temperature fluctuations through three mechanisms. First, density variations in the photon/baryon fluid cause adiabatic heating and cooling; denser regions emit hotter photons. This is the dominant mechanism at \( \ell > 100 \) and is the source of the acoustic peaks. Second, the photons in potential wells are gravitationally redshifted as they climb out; the mechanism causes denser regions emit cooler photons. This is referred to as the Sachs-Wolfe effect and is the dominant mechanism at large angular scales, i.e. those that did not fall within the sound horizon and oscillate before recombination. Third, the local velocity of
the photon/baryon fluid at the time of recombination imparts a Doppler shift to the CMB photons. Velocity extrema are 90° out of phase with density extrema, causing another set of peaks between the acoustic peaks. The Doppler shifts are always sub-dominant and these out of phase peaks do not appear distinctly in the power spectrum.

Two other effects are also considered primary anisotropies: the integrated Sachs-Wolfe effect, and photon diffusion damping.

The integrated Sachs-Wolfe (ISW) effect is the net gravitational redshift as photons pass in and out of potential wells after last scattering. In a static universe, the infall blueshift and the exit redshift would cancel exactly, but if potential wells grow or decay as photons pass through them, there is a net redshift or blueshift. The ISW effect may occur after recombination if the universe is not fully matter dominated (“Early” ISW) or, in an open or dark energy universe after matter domination has ended (“Late” ISW). The ISW effect contributes to CMB anisotropy at large and moderate scales (ℓ up to 300).

Photon diffusion damping causes an exponential decay in the power spectrum at small angular scales. Because recombination is not instantaneous, the CMB does not provide a snapshot of the an arbitrarily thin surface in the early universe. The finite thickness of the surface emitting CMB photons suppresses CMB structure at ℓ≈1000 and higher.

Effects which add or modify CMB anisotropies after recombination (other than ISW) are often called secondary anisotropies. Secondary anisotropies, such as a diffuse Sunyaev-Zeldovich effect or reionization, are most significant at small angular scales (ℓ>2000) and are most likely to be observed by high resolution interferometric experiments.

1.1.3 Cosmological Parameters from Temperature Anisotropy

Sub-degree scale measurements of CMB temperature anisotropy can be used to find cosmological parameters. In particular, the CMB power spectrum is sensitive to Ω_{tot}, the total density of the universe; Ω_b, the baryon density; n_s the primordial spectrum of density fluctuations; and τ_e, the optical depth to reionization. There is a lesser degree of sensitivity to other parameters such as Λ, the vacuum energy
density; and $\Omega_m$ the matter density. The effects of some of these parameters on the CMB power spectrum are illustrated in Figure 1.3.

![Figure 1.3: The dependence of the anisotropy power spectrum on cosmological parameters. Inflationary cosmologies are assumed. Top Left: The power spectrum is most sensitive to total density, or equivalently to spatial curvature. Lower density moves the acoustic peaks to smaller scales (higher $\ell$), and also increases large scale power from the ISW effect. Top Right: For a given curvature, there is some sensitivity to the dark energy density. The shifting of the acoustic peaks is degenerate with the stronger effect of curvature. Bottom Left: Baryon density affects relative peak heights because of its influence on the zero point of the acoustic oscillations. Bottom Right: Matter density affects the total power in the acoustic peaks as well as shifting peaks and influencing their relative heights. Figures by Hu & Dodelson (2002).](image)

The strongest sensitivity is to the total density $\Omega_{tot}$. For $\Omega_{tot} = 1$ (flat space), the power spectrum first peaks at $\ell \simeq 220$. For lower density (positive curvature) the first peak is at higher $\ell$, while for higher density (negative curvature) the first peak is at lower $\ell$. Other peaks are shifted proportionately. This effect is relatively insensitive to the physics at the time of recombination. Because the redshift
to last scattering and the size of the sound horizon are well estimated, the physical scale corresponding to the first peak can be calculated. The apparent angular scale depends primarily upon the curvature of light rays and is a direct measurement of the curvature of space.

Other parameters reflect earlier physics. The change in relative peak amplitudes with baryon density, for example, is a consequence of the gravitational attraction of baryons shifting the zero-point of the acoustic oscillations before recombinations.

The impact of various parameter changes on CMB power spectra has been widely discussed in the literature, and can be calculated using publicly available numerical tools (e.g. CMBFAST Seljak & Zaldarriaga (1996)).

1.2 Observational History

In this section, we outline the history of CMB observations over the last forty years. A more detailed account of observations up to the early 1990’s can be found in Partridge (1995).

The CMB was first observed by Penzias and Wilson at Bell Labs in 1964 as an unknown ‘excess noise’ with a blackbody equivalent temperature of 3.5±1.0 K in a radio telescope observing at a wavelength of 7.35 cm. At the same time, Robert Dicke at Princeton was promoting the construction of a specialized telescope to detect the CMB. It was Dicke who first argued that the excess background observed by Penzias and Wilson was a relic of the Big Bang (Penzias & Wilson (1965), Dicke et al. (1965)). Various theories were proposed in the late 1960’s to explain the measured signal without requiring a Big Bang, though none could account for an isotropic background with a purely thermal spectrum.

1.2.1 Spectral Measurements

Measurements of CMB intensity in other spectral bands quickly followed. A Princeton experiment already in progress measured a blackbody equivalent temperature of 3.0±0.5 K at 3.2 cm (Roll & Wilkinson (1966)). In total, over a dozen consistent measurements of the CMB temperature were published in the late 1960’s over a range of wavelengths from 0.33 cm to 73.5 cm (90 GHz to 0.41 GHz). Ob-
servational efforts were somewhat reduced in the 1970's and rejoined in the 1980's yielding increasingly convincing evidence for a precise blackbody spectrum (Smoot et al. (1987), Sironi et al. (1991)). The limitation of these measurements, made with radio telescopes, was their inability to measure the high frequency Wien region of the CMB spectrum.

Higher frequency measurements were not successful until the end of the 1970's, when bolometric receivers on balloon-borne and rocket-borne platforms were used to overcome the problem of atmospheric emission. The first experiment to confirm the expected reduction of power in the Wien tail of the CMB spectrum was conducted by Paul Richards and Dave Woody of UC Berkeley. The experiment, a balloon-borne Michelson interferometer with a bolometric detector, constrained CMB power over a range of 75 GHz to 720 GHz (4 mm to 0.4 mm) giving clear evidence of a peak in the frequency spectrum (Woody & Richards (1979)). These measurements were further refined over the next decade by several groups (Gush (1981), Peterson et al. (1985), Matsumoto et al. (1988)).

In the early 1990's, the FIRAS instrument on the COBE satellite provided a definitive measurement of the spectrum over the range of 5 GHz to 500 GHz (1.0 cm to 0.1 mm). These data reliably disprove any significant overall distortion of the CMB from a purely thermal spectrum (Mather et al. (1999)).

1.2.2 Anisotropy Measurements

The original Penzias and Wilson measurement sets an upper limit of \( \sim 20\% \) on CMB anisotropy. This was quickly improved to \( \sim 10^{-3} \) by the use of differential instruments sensitive only to the variations in the background (Partridge & Wilkinson (1967)). Continuing improvements led to the detection of the CMB dipole and eventually the intrinsic anisotropy.

In the late 1970's, the dipole was measured using radiometers on balloons and high altitude aircraft (e.g. Smoot et al. (1977)). Over the next decade, several groups obtained increasingly precise measurements; the results from the mid-1980's are quite similar to the current best value, \( 3.358 \pm 0.023 \) mK from the DMR instrument of the COBE satellite (Smoot et al. (1992), Lineweaver et al. (1996)).

Smaller scale anisotropy was not detected until much later. By the mid-
Figure 1.4: **Left:** A composite of CMB anisotropy measurements from over two dozen data sets including MAXIMA, BOOMERANG, DASI, CBI, VSA, and Archeops (Max Tegmark 10/02). **Right:** The measurement from the MAXIMA-I alone, along with the best fit model curve. Points scattered around zero are difference data from subtracted maps of independent detectors.

1970's an upper limit of $\sim 10^{-3}$ was established for anisotropy on angular scales as small as $1'$, primarily by ground based observations. The next generation of experiments included high altitude observations and early bolometric receivers. By the mid-1980's, the upper limit on small scale anisotropy was well below $10^{-4}$.

In the early 1990's, two balloon experiments including MAX (the predecessor to MAXIMA) gave statistical detections of CMB anisotropy at the $10^{-5}$ level (Fisher et al. (1990), Meyer et al. (1991), Alsop et al. (1992)). Shortly thereafter, the COBE DMR provided a completely unambiguous detection of anisotropy at angular scales of $7^\circ$ and higher (Gorski et al. (1996)).

From the early 1990's to the present, anisotropy measurements have pushed to increasingly small angular scales and increasingly large angular dynamic range. Data from the ground-based experiments Saskatoon (Netterfield et al. (1997)) and CAT (Scott et al. (1996)) together provided the first evidence of an acoustic peak near $\ell = 220$ in the angular power spectrum. At present, the balloon experiments MAXIMA (Hanany et al. (2000), Lee et al. (2001)), BOOMERANG (Netterfield et al. (2001)), and Archeops (Benoit et al. (2002)) and ground-based interferometers DASI (Halverson et al. (2002)), CBI (Mason et al. (2002)), and VSA (Scott et al. (2002)) have provided consistent measurements of the CMB power spectrum, including high
signal-to-noise measurements of the first acoustic peak. While no single experiment has so far measured higher acoustic peaks with high signal-to-noise, the combined data strongly suggest the existence of at least two more peaks.

Strong upper limits on polarization anisotropy have been obtained with the PIQUE (Hedman et al. (2001)) and POLAR (Keating et al. (2001)) experiments and, very recently, the DASI experiment is believed to have made a detection (Leitch et al. (2002)). Study of polarization anisotropy is an active field of CMB research, and is the aim of MAXIPOL, the follow-up to MAXIMA (Chapter 9).

1.3 Technical Considerations

Experimental efforts to tap the enormous potential of the CMB have yielded great results since the early 1990's. The recent rapid progress of the field owes both to improved detector technologies and to a strong commitment by the observational community.

The small size of CMB anisotropy compared to astronomical foregrounds, side-lobe sources, atmospheric emission, and the background loading of the CMB itself presents a serious challenge. Time domain noise correlations are problematic for data analysis and require a carefully planned scan strategy. A further complication has been the relatively late development of detector technologies in the optimum frequency range of 20 GHz to 300 GHz.

1.3.1 Optical Signals

CMB anisotropy is much smaller than optical backgrounds and parasitic signals. These unwanted signals are of three general types: constant loading, variable but non-sky stationary parasitics, and sky stationary parasitics (foregrounds). Stable optical backgrounds, referred to as loading, come from the CMB, the atmosphere, and the telescope. For a properly optimized detector system, loading contributes purely statistical photon counting noise (§2.4.1).

Unstable but non-sky stationary signals come from side-lobe sources such as the Sun, Moon, and Earth and from variations in atmospheric or instrumental loading. Small contributions of this type can be acceptable given a number of observations of each sky region; over repeated observations, their effects will tend to
cancel out. Large signals, and those that are not purely uncorrelated with CMB, are a major problem (§8.2).

Sky stationary parasitic signals, referred to as foregrounds, are of astronomical origin (e.g. Galactic dust and radio point sources). Foreground contamination can only be controlled by observing regions of the sky in which it is small and/or well understood (i.e. sky selection), by observing at optical frequencies with lower foreground sensitivity, and by spectral discrimination (§8.1).

1.3.2 Detector Technologies

Despite rapid advances made in the past decade, detectors in the 20 GHz to 300 GHz range are not yet fully developed. Continuing improvements in detector technology have enabled tremendous progress in CMB cosmology and will continue to do so in the future. Presently, two technologies are widely used by CMB experiments: bolometers, which are used in MAXIMA, and high electron mobility transistors (HEMTS).

Bolometers are total power square law detectors, best suited to observations at optical frequencies of 90 GHz and higher. Some bolometers, as in MAXIMA, are coupled to incident light via thin metal films, for large optical bandwidth and polarization independent sensitivity. Antenna coupled bolometers are being developed for faster response times and polarization discrimination. The chief advantage of bolometers is their high sensitivity; single MAXIMA detectors have achieved noise equivalent temperatures of less than 100 μK $\sqrt{sec}$, which is comparable to the photon noise limit. In practice, bolometric experiments have also benefited from the relative small effect of extragalactic point sources at their operating frequencies.

Bolometer technology is in a period of rapid development. In the early 1990's bolometers were often hand made, suffered significant performance variations between devices, and were rarely used in arrays. MAXIMA-era bolometers are made by a combination of photolithography and manual construction, and are typically operated in small arrays. Recent work has focused on the creation of arrays of hundreds of well matched bolometers with little or no manual construction. The success of these efforts is critical for future challenges such as the measurement of CMB polarization anisotropy.
The main disadvantage of bolometers is their operational complexity. Bolometers require extremely low operating temperatures (100 mK for MAXIMA), attainable only with sophisticated cryogenic systems and they are relatively sensitive to microphonic noise. In addition, since bolometers are total power sensors, they are insensitive to the phase of incident radiation and cannot be used for interferometry. Finally, the atmosphere is relatively emissive at bolometric frequencies, making ground-based observations difficult.

HEMT's are coherent amplifiers best suited to observations at frequencies below 90 GHz. They inherently preserve phase and polarization information. Modern HEMT based experiments usually take advantage this phase sensitivity for interferometry. HEMT's operate optimally at temperatures of tens of Kelvins, which can be readily achieved and maintained with simple cryogenic systems. At the lower optical frequencies of HEMT-based systems, atmospheric emission is relatively weak, and most HEMT systems are ground-based.

The main disadvantage of HEMT based experiments is lower sensitivity caused by their quantum noise limit at CMB frequencies and their relatively small optical bandwidth. A secondary disadvantage is the strong influence of radio point sources at their lower operating frequencies.

1.3.3 Noise Correlations

A common feature of all CMB experiments is low time stream signal-to-noise combined with noise correlations. Time domain noise correlations are caused by detector time constants and by drifts in detector temperatures or amplifier gains. Correlations also result from the instrumental high pass filtering often used by CMB experiments to suppress backgrounds. Because of correlations, CMB temperature at a given location on the sky cannot be accurately determined by simply averaging the detector output at that position.

The effects of temporal correlations are minimized by an observation pattern that revisits of each position on the sky over several uncorrelated time scales and along a variety of spatial trajectories (cross-linking) as discussed in Section 4.2.

Even with an excellent scan strategy, noise correlations must be analyzed explicitly (§6.4 to §6.6); this is the reason for the computational intensity of CMB
data analysis.
Chapter 2

The MAXIMA Experiment

This chapter describes the MAXIMA experiment. A brief introduction (§2.1) and an outline of the scientific goals and approach (§2.2) are followed by discussions of the telescope and optics (§2.3), the detectors (§2.4), and the cryogenic receiver (§2.5).

Pointing is discussed in detail in Chapter 4. More detailed discussions of optics and detectors can be found in Winant (2003). Lee et al. (1999) also describes the instrument.

2.1 Introduction

MAXIMA is a balloon-borne telescope designed to measure the anisotropy of the CMB over a wide range of angular scales ($\ell = 35$ to 1500). Over the course of two flights, in 1998 and 1999, high resolution observations have been made of 300 deg$^2$ of the sky. Results have been released (Lee et al. (2001), Hanany et al. (2000)) and cosmological implications have been explored by members of the MAXIMA collaboration (Stompor et al. (2001), Balbi et al. (2000)) as well as third parties.

The experiment is based at the University of California, Berkeley, and includes collaborators from the University of Minnesota and worldwide (Appendix A).
2.2 Goals and Design

The primary scientific objectives of MAXIMA are to distinguish between models of cosmological structure formation, and to measure parameters within these models. The $\ell$-space coverage and resolution of the experiment are well suited for measurement of the first three acoustic peaks of inflationary models. Measurements in this region are a powerful tool for testing the general predictions of inflation and for parameter estimation (Chapter 1, Hu & Dodelson (2002)).

In addition, MAXIMA data have been a realistic test of analysis methods and tools (Stompor et al. (2002a)). Treatments have been developed for problems such as mild beam asymmetry (Wu et al. (2001a)), foreground discrimination, scan synchronous noise, and detection of spatial non-gaussianity (Wu et al. (2001b)).

MAXIMA has been used to test new technologies. In particular, MAXIMA is the first CMB experiment to have used 100-mK spiderweb bolometers, similar to those planned for the Planck Surveyor. The combination of these detectors (§2.4, Bock et al. (1995)) and an adiabatic demagnetization cooling system has provided instrumental sensitivity of $\sim 40 \mu$K $\sqrt{sec}$, the best reported by any CMB experiment.

Experimental Concept

MAXIMA is a bolometric instrument, making CMB observations at relatively high optical frequencies (150 GHz and higher). In order to avoid atmospheric emission, observations are made from an altitude of $\sim 40$ km during multiple balloon flights. The relatively short duration of the balloon-borne observations is offset by the use of a 16-element array of single color photometers with extremely sensitive detectors. Good angular resolution ($10'$) and large sky coverage make the experiment sensitive over a wide range of angular scales. The compact and well cross-linked scan pattern is optimized for extracting the angular power spectrum. The use of three spectral bands allows discrimination between the CMB and foreground sources. MAXIMA benefits from exceptionally precise pointing reconstruction ($1'$) and accurate calibration (4%). The instrument is designed to consistently survive balloon flights, and has been successfully recovered after a test flight and two science flights.

A bolometric receiver, including cooled optics, and an external primary mirror are mounted on a two axis attitude controlled frame. Data are collected
during balloon flights lasting one night. There have been two such flights providing 3 hours and 8 hours respectively at altitudes of 36 km to 39 km. Each flight consists of a pair of cross-linked CMB observations covering areas of 122 deg$^2$ and 225 deg$^2$ with a signal-to-noise of $\sim 5$ at the 10$'$ beam size. Each flight also includes calibrations from the CMB dipole and from a planet. More details about the flights are available in Chapter 3.

The detector system is a high sensitivity array of spiderweb bolometers cooled to 100 mK by means of adiabatic demagnetization (§2.5.1). The array consists of 16 single color pixels, each projected onto the sky with a 10$'$ FWHM beam size. Eight detectors observe at 150 GHz, four at 230 GHz, and four at 410 GHz. Optical bandwidths are 45 GHz, 65 GHz, and 35 GHz respectively. The MAXIMA array provides a combined sensitivity of $\sim 40 \mu$K $\sqrt{sec}$ (See Appendix C for single detector sensitivities). Overall detector responsivity calibration (Chapter 5) is obtained from the CMB dipole with an accuracy about 4%.

The telescope is an off-axis Gregorian system with a 1.3-m diameter primary mirror providing a 10$'$ beam size (FWHM) for all detectors (§2.3). Due to fluctuating atmospheric emission at our observing wavelengths, the telescope is mounted on a high-altitude balloon-borne frame.

The primary attitude sensor is a boresighted CCD star camera. Pointing reconstruction is accurate to $\sim 1$" for CMB observation. Requirements on the selection of observing regions, scan pattern, cross-linking, and scan speed are presented in Section 4.1. A key element of the MAXIMA observation strategy is the active modulation of the primary mirror, which moves the telescope beams $\pm 2^\circ$ at 0.45 Hz.

**MAXIMA Data Sets**

Data were collected in August 1998 (MAXIMA-1) and June 1999 (MAXIMA-II). The MAXIMA-1 data have been analyzed to produce a 40,000-pixel map of 122 deg$^2$ of the cosmic microwave background and this map has been used to estimate the angular power spectrum.

The lowest multipole bin (largest angular scales) measured by MAXIMA spans $\ell=36$ to $\ell=110$. Two factors limit our ability to measure anisotropy at the largest scales: sky coverage and low frequency noise. The highest multipole bin
Figure 2.1: A mechanical drawing of the MAXIMA telescope from an elevated front/side perspective. Rays representing the telescope beam are shown reflecting from the primary mirror into the cryogenic receiver. Electronics housed in the rectangular boxes on the sides of the instrument include the pointing system, data multiplexers and digitizers, and telemetry and command interfaces. Near the top of the telescope are motors controlling azimuthal orientation. The inner assembly consisting of the primary mirror and the receiver is tilted relative to the outer frame to aim the telescope in elevation.

(smallest angular scales) measured spans $\ell=1086$ to $\ell=1235$. Limiting factors in our high $\ell$ measurements are: beam size and characterization, scan speed and detector time constants, pointing reconstruction, and integration time per pixel.

2.3 Telescope and Optical System

2.3.1 Optics

The MAXIMA telescope is a fast (f/1), off-axis Gregorian system consisting of a 300-K primary mirror and two cold reimaging mirrors. The primary mirror, produced by Dornier Satellitensysteme, is an underilluminated 1.3-m off-axis
Figure 2.2: The optical system of the MAXIMA telescope. The telescope is a fast (f/1) Gregorian system, with a prime focus at the window of the cryogenic receiver. The primary mirror, a 1.3-m diameter underilluminated paraboloid at ambient temperature, is modulated about the indicated axis. The secondary and tertiary reimaging mirrors correct aberrations from the primary. A 4-K Lyot stop defines the illumination of the primary. An array of feed horns channels light to the bolometers, which are held in resonant cavities. Optical filters are located at the prime focus, the Lyot stop, and after the feed horns. All detector channels have a beam size of 10$'$ FWHM.
Figure 2.3: Beam maps of the same 150-GHz detector in MAXIMA-I (left) and MAXIMA-II (right). This is flight data from observations of Jupiter (MAXIMA-I) and Mars (MAXIMA-II). Contours shown are 90%, 70%, 50%, 30%, 10%, and 1% of the maximum response. The 1% contour in the MAXIMA-II beam map is noisy because of the relatively low intensity of Mars. Beams were more symmetric and gaussian in MAXIMA-II, due to better telescope focusing. Wu et al. (2001a) shows that the beams in both cases can be approximated as symmetrical.

paraboloid constructed from a lightweight honeycomb material. The mirror is actively modulated during observations. The prime focus is located near the window of the cryogenic receiver. Two cold mirrors inside the receiver are off-axis ellipsoids with corrections to compensate for aberrations from the primary. A cold lyot stop between the tertiary mirror and the focal plane helps to define the beams and strongly suppress telescope side-lobes.

The cold optics (21-cm secondary mirror, 18-cm tertiary mirror, and lyot stop) are contained within a heavily baffled, liquid He-cooled optics box. The optics box maintains a temperature of approximately 3 K during flight. All non-optical surfaces inside the optics box are coated in a combination of Stycast epoxy\(^1\), carbon black powder, and glass beads. The resulting material absorbs stray far infrared radiation with high efficiency (Bock (1994)), further reducing the potential for side-lobe response.

The focal plane array consists of the entrances of 16 copper feedhorns. The feedhorns for the 150-GHz detectors are single-moded and consist of back-to-back straight cones connected by a length of cylindrical waveguide (Figure 2.4, left). The

\(^1\)A filled epoxy produced by Emerson and Cuming.
Figure 2.4: Bolometer Feedhorns. The narrow openings (top) are at the focal plane. The bases are screwed into a liquid $^4\text{He}$-cooled plate. Radiation leaving the horns passes into 100-mK assemblies holding band-defining optical filters and the bolometers. **Left:** A straight-walled, singled-moded 150-GHz feedhorn. A conical section (4.7° flare) leads from the focal plane to a 1.4-mm diameter straight waveguide. A second conical section (3.7° flare) leads to the exit of the feedhorn. **Right:** A multi-moded feedhorn for the 230-GHz and 410-GHz channels. The entrance of the horn consists of back-to-back Winston cones. These are coupled to a conical section (2.4° flare) leading to the exit of the horn.

Feedhorns for the 230-GHz and 410-GHz detectors are multi-moded and consist of back-to-back Winston horns coupled to a straight cone (Figure 2.4, right). Both types of feedhorn end in a straight cylinder, screwed into a liquid $^4\text{He}$-cooled plate. Light exiting a horn passes through a 0.5-mm gap before entering a 100-mK cylindrical waveguide holding optical filters and terminating in the bolometer cavities.

A neutral density filter (NDF) can be inserted into the optical path between the secondary and tertiary filters. The NDF has a transmittance of 1% and is used to simulate in-flight optical loading during ground tests.

The main lobes of all telescope beams were measured in flight by observing the planets Jupiter and Mars. Planets are effectively point sources for our beam size and are detected with signal-to-noise ratios of 100 to 1000 (Chapter 5).

The 1.3-m primary mirror is continuously modulated about the optic axis
(Figure 2.2) in a rounded triangle wave pattern with an amplitude of 4° at a frequency of 0.45 Hz. Section 4.3.3 describes the primary modulation and its role in the MAXIMA scan pattern.

**Far Side Lobes**

Spurious signals from bright sources outside the main lobe of our telescope beams are a potential source of systematic errors. The internal baffling of the cold optics and the beam-defining lyot stop strongly suppresses side-lobe response. In addition, the outside of the telescope is heavily baffle. This baffling is particularly effective at low elevation angles, blocking ground emission. Side lobe measurements are discussed in Section 8.2.1.

**2.3.2 Spectral Bands**

Bolometers are broad-band detectors. Optical filters are used to define spectral bands, which are chosen for atmospheric transparency and foreground discrimination. Band-defining mesh filters for each channel are located in light pipes before the bolometer cavities. These are cooled by the ADR to 100 mK. Each of the 230-GHz and 410-GHz detectors uses two filters for band definition: a capacitive lowpass filter and an inductive highpass filter. The 150-GHz detectors use only a lowpass filter; the lower edge of this band is defined by the size of the feedhorn. Sample spectra from each band are shown in Figure 2.5.

Three lowpass filters are placed before the focal plane. These serve to reduce the optical load on the 100-mK stage and to block high frequency resonant leaks in the band-defining filters. Two of the lowpass filters are located near the prime focus, just inside the cryostat window. The first (closer to the sky) is a reflecting capacitive mesh, with a cutoff at 18 cm\(^{-1}\) (540 GHz). The second is an absorptive alkali halide filter with a cutoff at 55 cm\(^{-1}\) (1650 GHz). A third lowpass filter, located at the lyot stop, is a capacitive mesh with a cutoff at 16 cm\(^{-1}\) (480 GHz). The first reflective filter is cooled to 77 K; the other two filters are cooled to 4 K.

The optical filters used in MAXIMA have been constructed at QMW and Cardiff.
Figure 2.5: The measured spectral response of the MAXIMA detectors. The peaks at 150 GHz, 230 GHz, and 410 GHz are measurements of sample detectors in each of our three observing frequencies. The three spectra are overlaid for comparison. The spectrum of the CMB peaks within the 150-GHz band. The measurement noise at low frequency does not represent real spectral leakage.

**Metal Mesh Filters**

Metallic mesh interference filters provide excellent out of band rejection and in band transmission at far infrared wavelengths. All of the band-defining filters and two of the initial lowpass filters are of this type. The theory of these filters has been widely studied (Ulrich (1967), Irwin et al. (1993)). Mesh filters of the kind used in MAXIMA are described in detail in Lee (1997).

The filters consist of dielectric substrates (1.5-μ taut Mylar) with thin metallic mesh coatings (0.2-μ copper). The metal is etched in a repetitive pattern by photolithography. The periodicity of the mesh is smaller than the radiation wavelength. The mesh thickness is negligible. The spaces between the metal mesh can be treated as transmission lines; different mesh spacings and geometries correspond to different equivalent oscillatory circuits as found by Ulrich (1967). Each filter consists of several layers of substrate and mesh, separated by spacing rings, pressed together, and glued.

Induced currents in the mesh give rise to reflected and transmitted waves. Absorption caused by ohmic losses in the mesh or by dielectric effects in the thin
mylar are of order $10^{-3}$ or less.

### 2.3.3 Multicolor Array

![Multicolor Array Diagram](image)

Figure 2.6: The layout of the MAXIMA focal plane. The arrows indicate the scan direction (azimuth modulation at constant elevation). All 16 channels project onto the sky with a 10' FWHM beam-size.

The MAXIMA focal plane is located at the exit of the optics box and consists of the entrances to the 16 single color feedhorns. Each feedhorn channels light through band-defining optical filters to a bolometer. The focal plane is laid out in four rows of four pixels; each row is projected onto a constant elevation on the sky, and consists of two detectors at 150 GHz, one at 230 GHz, one at 410 GHz (Figure 2.6).

### 2.4 Bolometers

MAXIMA uses an array of high sensitivity bolometers fabricated at JPL (Bock et al. (1995)). Here we present a very brief review of the important concepts in bolometric detection.

Bolometers are incoherent square law detectors most often used in sub-millimeter and far infrared applications. Their main advantages are high optical bandwidth (30% in MAXIMA) and low noise. The primary disadvantages are their
sensitivity to microphonic noise and their low operating temperature; the latter is the reason for the use of an adiabatic demagnetization refrigerator in MAXIMA (§2.5.1). Bolometers are the best detectors for observations of the CMB at frequencies of 90 GHz and higher.

![Diagram of bolometer setup](image)

**Figure 2.7:** A schematic of bolometer power and noise inputs. An absorber coupled bolometer consists of an absorber and an electrically biased thermistor, and a weak thermal link, $G$, to a thermal reservoir at a fixed temperature, $T_o$. Power inputs are shown in blue (solid arrows). $P_{tot}$ is the total optical and electrical power. Noise sources are shown in red (dotted arrows). Amplifier noise is not introduced in the bolometer itself and is not represented here.

In general, bolometers consist of an optical coupling, such as a metal film or antenna, and an electrically biased detector element, such as a semiconductor thermistor or a superconductor near a transition edge. Photons absorbed by the coupling deposit energy in the detector element causing an electrical signal. The bolometer is weakly coupled to a thermal reservoir, $T_o$ by a weak thermal conductance, $G$. The bolometer temperature, $T_{bol}$, is

$$T_{bol} = T_o + \Delta T = T_o + \frac{P_{opt} + P_{bias}}{G},$$

where $P_{opt}$ and $P_{bias}$ are the optical and electrical power inputs of the bolometer (Figure 2.7). In the simplest case, $P_{bias}$ and $G$ are nearly constant, and $\Delta T_{bol}$ varies
linearly with $P_{opt}$.

For a MAXIMA bolometer, the optical coupling is a metallic absorber, impedance matched to free space, in a resonant $(1/4$ wavelength long) cavity. The absorber is a layer of gold coated onto a spiderweb shaped substrate of silicon-nitride. The detector element is a Neutron Transmutation Doped (NTD) germanium thermistor produced at LBNL (Haller (1985)). The NTD is mounted at the center of the spiderweb and is biased with a constant electrical current. The thermal reservoir is the 100-mK ADR. The bolometer operates at $\sim 130$ mK.

MAXIMA is the first experiment to use microlithographed spiderweb absorbers at 100 mK. This design provides fast response time, a small cross-section to cosmic rays, and relatively low microphonic response.

### 2.4.1 Responsivity and Noise

The theory of bolometer noise and responsivity optimization is developed elsewhere (Richards (1994), Grannan et al. (1997)). Spiderweb bolometers of the type used in MAXIMA are described in Bock et al. (1995). MAXIMA bolometer optimization is found in Winant (2003).

Overall sensitivity to CMB temperature fluctuations is characterized by noise equivalent temperature (NET), with units $[CMB\ Temp.]\ [Time^{0.5}]$, given by,

$$NET = \frac{NEP}{S_{CMB}}$$

(2.2)

$S_{CMB}$ is responsivity to the CMB, defined by the change in bolometer voltage per change in CMB temperature $(V \cdot T^{-1})$, and $NEP$ is the noise equivalent power in the detector output voltage $(V \cdot sec^{0.5})$. NET describes sensitivity according to,

$$\Delta T_{cmb} = \frac{NET}{\sqrt{t_{obs}}}$$

(2.3)

where $\Delta T_{cmb}$ is the noise in the measured temperature of an observed region and $t_{obs}$ is the time spent observing that region. The NET generally varies with signal frequency, but single values are often quoted, representing averages over a band of signal frequencies used in the experiment. The NET is often presented in units of $[CMB\ Temp.]\ [Freq.^{-0.5}]$, representing the noise density per unit bandwidth. In this

form, the quantity is larger by a factor $\sqrt{2}$, because 1 second of integration samples a bandwidth of 0.5 Hz.

The white noise NET for a well optimized bolometer is improved by reducing the bolometer operating temperature and reducing the optical background loading. The overall NET is a quadrature sum of NETs from different noise sources: heat sink thermal fluctuation noise, Photon counting noise, thermistor Johnson noise, and amplifier noise.

Thermal fluctuation NET is given by $\sqrt{4kT^2G}$, where $T$ is the bolometer temperature and $G$ is the thermal conductivity of the heat sink. Reduced thermal reservoir temperature ($T_o$) allows a properly optimized bolometer to operate at lower temperature. The dependence of thermal fluctuation noise NET on $T_o$ varies with the type of thermal link between the detector and the thermal reservoir. $T_o^{1.5}$ holds for metals, while $T_o^{2.5}$ holds for insulators and superconductors. In practice, the thermal link consists of multiple materials with different scalings. Reductions in optical loading allow roughly linear reductions in $G$, which is directly proportional to thermal fluctuation NET.

Photon noise is the quantum statistical fluctuation in photon flux. Its importance is minimized by increasing the number of ‘signal’ photons and decreasing the total number of photons absorbed by the detector. For a given signal size (i.e. optical band and optical efficiency), photon NET decreases as at least the square root of optical background loading. Because the loading can not be reduced below the CMB flux, photon NET represents the fundamental limit of bolometer sensitivity in a given optical band.

Johnson noise NET is given by $\sqrt{4kTR/|S|^2}$, where $T$ is the bolometer temperature, $R$ is the electrical resistance of the bolometer, and $S$ is the optical responsivity. Electrical resistance is matched to the input impedance of the amplifier. The Johnson noise NET is reduced by decreasing temperature and increasing responsivity. Assuming properly optimized bolometers, responsivity increases as $P_{tot}^{-n}$, where $P_{tot}$ is the total optical and electrical load and $n$ is at least 1.

Amplifier noise is the quadrature sum of voltage noise and current noise through the thermistor. The noise level (as opposed to NET) is not strongly affected by temperature or optical load in a well optimized system. Amplifier NET is most strongly affected by responsivity, and is minimized by reduced optical background
loading and the use of high quality amplifiers.

Appendix C presents tabulated values of white noise NETs achieved in the MAXIMA flights.

Low frequency noise originates from variations in the bolometer electrical resistance and thermal reservoir temperature. The former term is relatively small for modern, well fabricated bolometers. In MAXIMA, the thermal reservoir fluctuations are substantial, due to the very low thermal mass of the 100-mK stage and the mechanical coupling of telescope modulation to the refrigerators. These fluctuations dominate the bolometer noise at frequencies below $\sim 0.5$ Hz. An additional contribution of low frequency noise from the amplifiers can be avoided by AC-biasing the detectors, as in MAXIMA.

### 2.4.2 Biasing and Readout

The MAXIMA bolometers are AC biased using 10-nA (100-mV) RMS sine waves at $\sim 300$ Hz. AC biasing provides strong rejection of low frequency electronic noise, particularly in the cryogenic preamplifiers. The exact bias frequency is chosen before flight to minimize narrow band microphonic pickup.

All bias signals are phase locked and have fractional amplitude variations of less than $10^{-6}$. The AC signals from the detectors are bandpass filtered and demodulated using a lock-in amplifier referenced to the bias generator. The demodulated signal is processed using a 15-mHz highpass filter to remove 1/f noise not rejected by the AC bias (primarily caused by detector temperature drifts). In addition, a four-pole butterworth lowpass filter with a 19-Hz cutoff is used to eliminate high frequency noise. The phase shifts caused by these filters are measured before flight and are removed in data analysis.

AC Bias generators and lock-in readout electronics are mounted to the outside of the receiver at ambient temperature.

### 2.4.3 Cryogenic Preamplifiers and Microphonics

NTD Bolometers are high impedance devices ($\sim 5$ MOhm for MAXIMA) with correspondingly high sensitivity to microphonic noise in wiring between the detector and the first amplifier. To minimize this sensitivity, preamplifiers are placed within
the cryogenic receiver, very close to the detectors. These differential amplifiers each consist of a matched pair of cryogenic JFETs\textsuperscript{2} operating at 150 K inside a sealed cavity within the liquid $^4$He-cooled portion of the cryostat. Between the detectors and the preamplifiers are $\sim$8 inches of wiring and $\sim$4 inches of circuit board traces. The wiring is potted in epoxy or varnished to a rigid surface over most of its length. The stiffness of the wiring and traces minimizes microphonic response near and below the bolometer bias frequency ($\sim$300 Hz).

The JFET amplifiers typically contribute non-negligible, but subdominant white noise ($\sim$5 nV/$\sqrt{Hz}$). Though the amplifiers are the dominant source of 1/f noise in the bolometer signal (1/f knee $\sim$10 Hz), this low frequency noise is rejected by the AC biasing of the detector. The performance of the JFETs has been acceptable, but with little margin. For slightly lower detector noise, better JFETs would have been needed.

2.4.4 Time Constants

Detector time constants limit telescope scan speed and act as a lowpass filter on the data. MAXIMA detector time constants vary from 1 msec to 10 msec (typically 6 msec to 8 msec). The slowest detectors, at 10 msec, are fast enough for the combination of 4°/sec scanning speed and 10' FWHM beam size, using the $\frac{FWHM}{2\pi}$ criterion of Hanany et al. (1998).

The filtering effects of detector time constants, like those of electronic filters, must be deconvolved during data analysis. Time constants are measured before flight, to an accuracy $\pm$0.5 msec. Flight data from planet observations are used to refine this to $\pm$0.2 msec (Winant (2003)).

2.5 Receiver

The MAXIMA cryostat (Figure 2.8) houses the secondary optics, the bolometer array and preamplifiers, an optical calibration source, and the cryogenic coolers. There are also a number of diagnostic devices including ‘dark’ detector channels not exposed to the CMB and a variety of internal temperature monitors.

\textsuperscript{2}Infrared Laboratories TIA JFETS
Figure 2.8: A mechanical drawing cutaway of the MAXIMA cryogenic receiver. The bottom section of the receiver contains heavily baffled, liquid $^4$He-cooled optics and the internal relative calibration source. The middle holds the bolometers and bolometer feedhorns, cryogenic JFET preamplifiers, and thermal switches. The top section of the receiver contains the cryogenic systems, with the low temperature refrigerators (liquid $^3$He and adiabatic demagnetization) surrounded by the open-cycle liquid $^4$He tank.

### 2.5.1 Cryogenics

The MAXIMA receiver makes use of four cooling systems: open-cycle LN$_2$ and liquid $^4$He tanks, a closed-cycle liquid $^3$He refrigerator, and an adiabatic demagnetization refrigerator.

A 13-liter LN$_2$ tank cools an outer layer of radiation shielding to 77 K. This temperature drops to $\sim$50 K when the LN$_2$ tank is exposed to vacuum, as in flight. The LN$_2$ temperature radiation shields are covered with thin, low emissivity aluminum foil.

Inside the LN$_2$-cooled space is a 21-liter, open-cycle liquid $^4$He tank and an additional layer of shielding at liquid $^4$He temperature. The outer shell of the cold optics box serves as part of this radiation shielding. The outside of these shields is low emissivity aluminum, while the inner surfaces are coated with a blackening mixture. The blackened interior absorbs high temperature radiation that leaks past the shields.
<table>
<thead>
<tr>
<th>Cooler</th>
<th>LN$_2$</th>
<th>$^4$He</th>
<th>$^3$He</th>
<th>ADR</th>
</tr>
</thead>
<tbody>
<tr>
<td>Temperature (Keldins)</td>
<td>50</td>
<td>2-3</td>
<td>0.35</td>
<td>0.1</td>
</tr>
<tr>
<td>Hold Time (Hours)</td>
<td>24</td>
<td>&gt;30</td>
<td>&gt;36</td>
<td>12</td>
</tr>
<tr>
<td>Thermal Cycle</td>
<td>Open</td>
<td>Open</td>
<td>Closed</td>
<td>Closed</td>
</tr>
</tbody>
</table>

Table 2.1: Temperatures, cooling durations, and type of thermal cycle for the MAXIMA cooling systems. Numbers are quoted for flight conditions (high altitude, nighttime). All cryogenic systems have ample capacity for a MAXIMA balloon flight.

Within the liquid $^4$He temperature space are the optics, the detectors, the JFET preamplifiers, the sub-Kelvin coolers and a variety of thermometers. The optics and most electrical components are thermally linked to the coldplate, a 1.0-cm thick copper plate forming the bottom of the liquid $^4$He tank.

Various locations in the liquid $^4$He space range in temperature from 4 K to 6 K, depending on thermal load and proximity to the helium tank. When the liquid $^4$He is exposed to vacuum, for testing or in flight, these temperatures drop to 2 K to 3 K. This causes a significant drop in the background loading of the bolometers.

Inset into the liquid $^4$He tank are the adiabatic demagnetization refrigerator (ADR) and the liquid $^3$He refrigerator. All wiring entering the receiver passes through the liquid $^4$He and LN$_2$ tanks and is made of low thermal conductivity stainless steel leads, ending in cold radio frequency filters (§2.5.3).

**Low Temperature Refrigerators**

Two cooling stages beyond liquid $^4$He are used to reach sub-Kelvin temperatures. The first is a closed-cycle $^3$He refrigerator using 40 liters (stp) of $^3$He which provides a temperature of ~350 mK under flight conditions. The other is an adiabatic demagnetization refrigerator (ADR) (Hagmann et al. (1994)) which provides a temperature of 100 mK. The MAXIMA ADR consists of 40 grams of ferric ammonium alum (FAA), a high permeability ferromagnetic salt, inside a 2.5-Tesla electromagnetic coil.

Mechanical supports for the ADR stage are constructed from thin-walled, low-conductively Vespel$^3$ tubes and taut kevlar string. Liquid $^3$He-cooled copper

$^3$A polymer resin produced by DuPont.
straps are used to intercept all supports for the 100-mK stage, reducing the conductive thermal load from the 4-K stage by more than a factor of ten.

An electrically controlled superconducting solenoid is used as a heat switch, magnetically closing gold plated flanges to create a thermal link between the liquid $^3$He refrigerator and the ADR. It is used to provide the isotemperature phase of the ADR thermal cycle.

The heat load on the ADR - hundreds of nanowatts - causes a temperature increase of $\sim 3$ mK per hour. This is slow enough that continuous temperature control is not needed. Periodic corrections are made via ground-based commanding of the ADR electromagnet current. During the MAXIMA-I flight the ADR field was adjusted twice after the initial cool down ($\sim 2$ hours between adjustments) with temperature drifts of $<5\%$ between adjustments. During the MAXIMA-II flight, a problem with the ADR magnet control electronics made commanding impossible. Over $\sim 6$ hours of MAXIMA-II, the ADR temperature drifted by $\sim 20\%$. During both flights, the effects of temperature drift on bolometer responsivity were monitored using the internal calibration source.

2.5.2 Internal Relative Calibrator

Bolometer responsivity varies over the duration of a flight, primarily because of variations in the temperature of the nominally 100-mK ADR. These variations are monitored using a stable internal calibration source (a stimulator) consisting of a thin nickel-chromium layer ($2$ mm $\times$ $2$ mm) back with a sapphire substrate. The metal layer is impedance matched to radiate efficiently into free space when heated. When a heating current $\sim 1$ mA is applied, the metal warms to $\sim 50$ K with a time constant of $\sim 1$ sec. The heating current is maintained for 10 seconds, and is applied every 20 minutes during flight. The stimulator is mounted inside the cold optics box, and fitted with a light pipe to illuminate the focal plane array from just outside the optical path. The illumination of the array is not uniform, with detectors closer to the calibrator receiving about twice the flux of the more distant detectors.

The source is extremely stable with negligible resistance fluctuations and $<1\%$ current fluctuations. The on-state calibrator temperature is further stabilized by a weak, temperature dependent thermal link to the liquid helium stage.
The absolute flux is not well measured, so the stimulator is used purely for monitoring of responsivity variations over time. Absolute calibration is obtained from celestial sources (the CMB dipole and planets). The use of stimulator data in detector calibration is discussed in Section 5.4.

2.5.3 RFI protection

During flight, several radio transmitters are used for telemetry. Each radiates 15 to 40 Watts at frequencies of 1.5 GHz and higher. Extensive filtering is required to prevent pickup from these sources in the detectors.

The receiver itself consists of three metallic shells which serve as partial Faraday cages. External cabling is also fully enclosed in a metallic shell. All cables between the receiver and the readout electronics and all cables exiting the readout electronics pass through commercial RF filtered connectors (Amphenol FPT02 Series; 60-dB attenuation at 1.0 GHz). Within the receiver, all wiring is potted in 27 cm of Eccosorb CR-124, a metal-filled epoxy\(^4\) that acts as a radio frequency lowpass filter (30-dB attenuation at 1.0 GHz).

Radio frequencies are filtered directly at the detectors using an inductive/capacitive (LC) lowpass filter with a 1.0 GHz cutoff. These filters are fully described in Winant (2003).

\(^4\)Commercially available from Emerson and Cuming
Chapter 3

Observations

Figure 3.1: The MAXIMA telescope on “Tiny Tim”, the launch vehicle, shortly before launch on June 16, 1999

This chapter describes the two MAXIMA flights, including flight conditions, scan durations, flight trajectories, and the positions of the Sun and Moon relative to the scans.

3.1 MAXIMA-I

MAXIMA was first flown on August 2, 1998. Field work began in late April and MAXIMA was flight ready on June 10. Due to unusual weather patterns (El
<table>
<thead>
<tr>
<th>Flight</th>
<th>Hours at Maximum Altitude</th>
<th>1st CMB Scan (hours)</th>
<th>2nd CMB Scan (hours)</th>
<th>CMB Dipole Scan (hours)</th>
<th>Planet Scan (hours)</th>
<th>Daytime Test Data (hours)</th>
</tr>
</thead>
<tbody>
<tr>
<td>MAXIMA-I</td>
<td>3.78</td>
<td>1.63</td>
<td>1.37</td>
<td>0.57</td>
<td>0.57</td>
<td>0.00</td>
</tr>
<tr>
<td>MAXIMA-II</td>
<td>7.78</td>
<td>2.42</td>
<td>2.15</td>
<td>0.62</td>
<td>0.63</td>
<td>1.65</td>
</tr>
</tbody>
</table>

Table 3.1: Scan and flight durations for both MAXIMA flights. Some calibration data were collected before the telescope reached maximum altitude. In the case of MAXIMA-I, this results in a total scan time greater than the time at maximum altitude.

Niño), there were no launch opportunities for nearly two months.

The payload was launched on August 8 at 00:58 UT (19:58 local time) from the National Scientific Balloon Facility (NSBF) in Palestine, TX (latitude 31.8°N, longitude 95.7°W). The maximum float altitude of 37.5 km was reached at 4:35 UT. The telescope traveled 189 km west and less than 1 km south before reaching maximum altitude. At float, the telescope drifted 405 km west and less than 1 km north. Descent began 3.8 hours later at 8:22 UT. Summer flights from the NSBF in Palestine are limited to a range of approximately 600 km. The MAXIMA-I flight was relatively short due to fast high altitude winds.

Four observations were conducted during the flight. First, the CMB dipole was observed in order to calibrate the responsibility of the detectors. The dipole observation was started before reaching float altitude and lasted from 03:37 UT to 04:11 UT. Next, two overlapping, cross-linked scans of CMB anisotropy were conducted over a 122-deg² region in the vicinity of the Draconis constellation. These scans occurred from 04:21 UT to 05:59 UT and 06:02 UT to 07:24 UT. Finally, observations were made of Jupiter to characterize the telescope beams and to calibrate the 410-GHz detectors, which are insensitive to the CMB dipole. Jupiter was observed from 07:30 UT until 08:04 UT.

The sun was at least 20° below the horizon for all observations. The Moon was below 20° elevation during CMB observations, and below the horizon for over an hour of the second observation. While above the horizon, it was below 30° elevation and was at least 70° from the scan region. The relative position of the Moon differed by ~20° azimuth and ~10° elevation between the two CMB scans. During the dipole
observation, the Moon was at \(\sim 30^\circ\) elevation, \(\sim 20^\circ\) below the scan. The Moon was below the horizon during the Jupiter scan. The Moon was 68% full during the flight.

### 3.2 MAXIMA-II

The second flight occurred in June 1999. Field work began in late April, and the instrument was prepared for flight in the first week of June. Footage of the launch and flight preparations can be found in Lucas (2000).

The weather in 1999 was much more favorable for a timely launch. The launch occurred on June 17 at 00:07 UT (19:07 June 16, local time). The telescope traveled 42 km east and 9 km south before reversing direction and reaching maximum altitude 97 km west and 1 km south of the launch position. The maximum float altitude of 38.0 km was reached at 04:34 UT. At float, the telescope drifted 490 km west and 42 km north. Descent began 7.8 hours later at 12:21 UT. The relatively slow high-altitude winds of early summer allowed us a considerably longer flight than MAXIMA-I.

As with MAXIMA-I, two CMB observations and two calibration scans were conducted. The first was an observation of Mars from 03:14 UT to 03:52 UT. Approximately one hour was spent on maintenance tasks. Two overlapping, cross-linked CMB scans were conducted from 05:04 UT to 07:29 UT and 07:31 UT to 09:40 UT. The observed region has an area of 225 deg\(^2\) and overlaps the MAXIMA-I region by 50 deg\(^2\). A calibration scan of the CMB Dipole was conducted from 09:42 UT to 10:19 UT. Further data were recorded from 10:20 UT to 11:59 UT as a test of the daytime performance of the instrument.

The sun rose to -20° elevation at 09:24 UT and to 0° elevation at 11:17 UT. Data collected after 10:20 UT have been used only as test data for future daytime balloon flights. The Moon was 17% full during the flight and was below the horizon during the dipole observation and both CMB observations. During the Mars observation, the Moon was 75° from the scan.
Chapter 4

Pointing and Attitude Reconstruction

Figure 4.1: The full reconstructed pointing for a single detector in both MAXIMA flights. MAXIMA-I is represented in green (light gray) on the right, and MAXIMA-II is in blue (dark gray) on the left. The scan region for each flight is boxed, and the \( \sim 50 \text{-deg}^2 \) overlap region can be seen at Right Ascension \( \sim 15 \) hours.

In this chapter we discuss the pointing and attitude reconstruction of the MAXIMA telescope. There are three main pointing related issues. First is scan strategy; the size and shape of the observation pattern are chosen for high angular power spectrum sensitivity and to minimize the effects of noise correlations and potential
systematic effects (§4.2). Second is pointing control; we have constructed a flexible pointing system, capable of realizing a variety of scan strategies (§4.3). Third is post-flight pointing reconstruction; the orientation of the telescope is determined to 1′ - roughly 10% of the FWHM of the telescope beams (§4.4). Section 4.1 lists specific requirements in all three areas.

A number of MAXIMA collaborators were deeply involved in pointing and attitude reconstruction, including Paolo DeBernardis (planning; scan strategy; flight hardware), Andrea Boscalleri (flight hardware), Barth Netterfield (flight software), Enzo Pascale (flight software), and Amedeo Balbi (pointing reconstruction). In addition, George Smoot, Amedeo Balbi, Andrew Jaffe, and Shaul Hanany developed flight planning software.

4.1 Requirements

In this section we summarize the various requirements for the pointing of the MAXIMA telescope. The implementation and strategies used to meet these requirements are detailed later in this chapter.

4.1.1 Sky Selection

The absence of Galactic dust was the strongest requirement for the observation region. Scan regions in MAXIMA-1 were also chosen to avoid known bright point sources. The expected signal from uncatalogued point sources is very small (Gawiser et al. (1998)), and was not a major concern. Dust, point sources, and other foregrounds are discussed in Chapter 8.

The scan regions for the two flights were chosen to have a modest (~50-deg$^2$) overlap, both as a consistency check and to facilitate the combination of the data sets.

Within these constraints, we selected scan regions offset by roughly 40° azimuth from the north celestial pole. This was useful for technical reasons discussed in Section 4.3.4.
4.1.2 Scan Speed

Scan speed determines the relationship between time domain data and the spatial map of the CMB. In particular, the modulation with the shortest characteristic time scale sets the minimum revisit time for an observed spot on the sky. The corresponding temporal frequency is roughly the minimum at which the data carry significant CMB information.\(^1\)

The MAXIMA detectors have a low frequency knee (1/f or steeper) in their noise profiles between 0.1 Hz and 1.0 Hz. The fastest modulation should be in this range or higher. The fastest MAXIMA modulation (0.45 Hz) is described in Sections 4.2 and 4.3.3.

4.1.3 Depth of Integration

An experiment with fixed integration time and partial sky coverage makes a trade-off between the integration time available for each pixel and the number of pixels observed. Equivalently, the trade-off is between noise and sample variance in power spectrum estimation. The optimal integration for measurement at a given angular scale is one that yields a signal-to-noise of ~1 at that scale (Tegmark (1997)).

A higher SNR with fewer pixels degrades the power spectrum estimation fairly slowly with decreasing pixel count. The opposite case (low SNR, many pixels) causes a much faster degradation in power spectrum estimation as a function of pixel count.

This trade-off is complicated by the fact that we are simultaneously measuring over a large number of angular scales. Optimal integration at large scales leads to too much sky coverage at small scales, while optimal integration at small scales leads to too little sky coverage at large scales. Because of the asymmetry in the optimization, and inherent difficulty of measurements at small scales (CMB anisotropy power drops dramatically at high \(\ell\)), we have optimized our integration time based on beam-sized (10') pixels. We deem this approach optimal for an experiment measuring a large range of angular scales. A final consideration is systematic error testing, which benefits from higher SNR.

Achieved integration time and signal-to-noise per pixel are summarized in

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\(^1\)While not an exact relationship, this is a useful rule of thumb. In general, the exact mapping between temporal and spatial frequencies is complex and best explored through simulations.
Section 4.2.

4.1.4 Scan Pattern

In the case of an ideal CMB experiment, with purely uncorrelated receiver noise, any modulation pattern yielding a constant integration time/pixel would be acceptable. In reality, low frequency detector noise leads to significant noise correlations. Given arbitrary pointing control, the effects of these correlations are best eliminated by a random scan pattern in which pixels in the scan region are measured with uniform probability at each detector sampling. This is obviously unrealistic. In practice, good results are obtained by the use of interlocking observations, with the scan directions tilted to provide cross-linking (Tegmark (1997)).

The overall shape of the scan region is also significant. A more compact scan region (as opposed to a highly elongated region with a large aspect ratio or exotic shape) is desirable, especially for measurement of large spatial modes (low \( \ell \)). Such a region contains more independent modes, particularly large modes, than an elongated region of the same area. In addition, edge effects (i.e. imperfections in the scan pattern near the edges) are minimized for compact scan regions.

Fast elevation modulations should be avoided. The column depth of the atmosphere varies roughly as the cosecant of the elevation and is a significant source of background loading on the bolometers. A modulation of this background loading would cause a modulation synchronous signal in the detectors. In addition, rapid changes in elevation angle disturb liquid cryogens in the receiver, leading to detector instabilities.

A good approach is repeated raster scans, tilted relative to each other for cross-linking. A raster-like pattern can be achieved by a periodic modulation in azimuth alone; the rotation of the sky effectively provides the modulation in the elevation direction. The MAXIMA scan pattern is of this general form. The details are presented in Section 4.2.

\(^2\)An example of cross-linking is shown Figure 4.3.
4.1.5 Precision of Control and Reconstruction

The absolute position of the scan boundaries must be controlled to \( \sim 0.5^\circ \) to maintain the shape of the scan region. Scan periods should be stable to at least 10\%, to ensure reasonable beam overlap from one period to the next. Elevation must be stable to a few arcminutes during a scan period, also to ensure overlap.

For pointing reconstruction, an accuracy of \( \sim 1' \) is desired. This yields negligible pointing-based error in the CMB power spectrum over most of our measured range. Pointing error is most significant at high \( \ell (\sim 10\% \text{ at } \ell = 1000) \), but with 1' accuracy remains subdominant at all angular scales.

4.1.6 Calibration Scan Requirements

In addition to CMB scans, we require observations of both the CMB dipole, and a bright point source (a planet) for calibration.

The dipole is the main responsivity calibrator. The spatial distribution of the signal makes it difficult to observe the dipole on time scales faster than detector drift. We require that the beams be scanned quickly over a large temperature contrast. The observation is repeated continuously for long enough to reduce calibration uncertainty to a few percent. The required pointing reconstruction accuracy of the dipole scans is \( \sim 10' \), corresponding to \( \sim 0.1\% \) calibration bias.

Planet scans are used primarily for beam measurement. They also provide a secondary responsivity calibration. For these observations, we require that a telescope beam cross the planet quickly compared to detector drift, that the beams be well sampled in two dimensions, and that pointing reconstruction be as accurate as for CMB observations.

As with CMB scans, calibration scans must be conducted at constant elevation. Due to the higher signal-to-noise of the calibration sources, cross-linking is not a requirement. Section 4.2 describes the actual observation strategy used for both calibration sources.
4.2 Scan Strategy

4.2.1 Selected Sky

Each MAXIMA flight has consisted of two cross-linked observations of a single patch of the CMB sky. The dust in these scan regions has a predicted in-band equivalent temperature of ~10.0 $\mu$K at 150 GHz with rms fluctuations of ~2.5 $\mu$K in units normalized to the CMB spectrum (Jaffe et al. (1999)). Tests of the spectral and angular profiles of the observed signals, as well as cross correlations with known dust maps, confirm the absence of significant dust contamination in our CMB data ($\S$8.1).

The MAXIMA-I scan region was chosen to contain no detectable point sources. For MAXIMA-II, this requirement was relaxed so that while no point source contribution is expected in the CMB power spectrum, a few bright sources might be detectable in the anisotropy maps, particularly at 410 GHz.

Figure 4.2: **Left:** A simulation of the double modulation in azimuth. The x-axis is the azimuthal position of the telescope beams, while the y-axis is time. The slower modulation is caused by the motion of the entire telescope, while the faster modulation is caused by the rotation of the primary mirror about the optic axis. **Right:** The scan pattern formed in RA and declination, combining the azimuth modulations with the rotation of the sky (data are shown from a MAXIMA-II scan). Note that lines of constant elevation move with the rotation of the sky, in this case spanning the plot in a diagonal arc from the lower left to the upper right. The gaps seen in this scan pattern are less than half the telescope beam size, so a continuous CMB map is obtained.
4.2.2 Modulation Pattern

Each CMB observation is conducted at a fixed elevation, while the telescope beams are moved actively in azimuth. The azimuth modulation defines one dimension of the roughly rectangular scan region. The rotation of the sky over the duration of the observation defines the other dimension.

The azimuth motion consists of two independent modulations. The primary mirror rotation provides a relatively fast modulation with a frequency of 0.45 Hz and an amplitude of 4.0° peak-to-peak. The motion of the entire telescope provides a slower modulation at a frequency of 12 mHz to 25 mHz with an amplitude of 4.5° to 9.0° peak-to-peak. The fast modulation prevents our data from being significantly corrupted by low frequency noise. Taken together, the two make our data relatively robust against modulation synchronous parasitic signals.

The two CMB scans of each flight observe the same region of the sky, but at different times and at different elevation angles. The rotation of the sky between these observations causes the two scans to be tilted relative to each other in sky-stationary coordinates. This cross-linking averages 22° in MAXIMA-I and 27° in MAXIMA-II.

The two azimuth modulations, the rotation of the sky, and the cross-linked revisitation are uncorrelated and are on radically different time scales, minimizing the effects of potential scan-correlated systematics.
4.2.3 Depth of Integration

Depth of integration is set by the total width of the combined azimuth modulation and the rotation rate of the sky. The average integration time per beam-size was \( \sim 2.5 \) seconds in MAXIMA-I and \( \sim 2.2 \) seconds in MAXIMA-II. This leads to an average expected noise level of \( \sim 60 \) \( \mu \text{K} \) per beam size area for our best single detector and \( \sim 40 \) \( \mu \text{K} \) for our published combination of four MAXIMA-I detectors. In practice, integration is several times longer in the center of the observed region and shorter near the edges.

4.2.4 Calibration Scans

MAXIMA is calibrated in flight using both planets and the CMB dipole. Planet observations are conducted by pointing the telescope directly above or below it as the planet rises or sets. The telescope tracks the position of the planet in azimuth, while remaining at fixed elevation. The primary mirror modulation moves the beams around the planet in azimuth. The beams are much larger than the planet, so each pass is effectively a one-dimensional map of the telescope response. The planet drifts steadily through the observed elevation giving complete two dimensional beam maps. Each crossing is \( \sim 20 \) bolometer samples (~0.1 sec) and there are several hundred crossings with good signal-to-noise over the course of the observation. The scan pattern is illustrated in Section 5.3.

The dipole is observed by rotating the entire telescope in azimuth at high speed (18°/sec; 3.3 RPM) at fixed elevation. The signal is detected at \( \sim 55 \) MHz. Rotations are conducted for about 30 minutes (~100 rotation). The primary mirror modulation has a small amplitude (4°) and has very little impact on dipole observations; the modulator was on during the MAXIMA-I dipole scan and off during the MAXIMA-II dipole scan.

4.3 The Attitude Control System

The pointing system, or attitude control system (ACS), serves both to control the orientation of the telescope in flight and to acquire the data needed for post-flight pointing reconstruction. The pointing system consists of attitude sensors,
a central feedback loop control computer, and motors. Some of the easily interpreted sensors are used in pointing control, while the most precise sensors, the CCD cameras, are used after the flight for pointing reconstruction.

Figure 4.4: A Schematic of the MAXIMA pointing system. The system acquires data for pointing control and post-flight pointing reconstruction. The central control computer reads all data, commands the motors, and handles remote communications.

4.3.1 Control Electronics

The Feedback Loop Controller (FLC) is the computer that controls the pointing system. The FLC reads data from the various sensors, applies a digital feedback algorithm, and sets the power level for the motors. Each of these tasks is performed once every 96 ms, synchronously with the bolometer data acquisition system.

Signals from most sensors are sampled each cycle by analog-to-digital converters inside the control computer. Data from the CCD cameras are processed by a separate computer, and passed to the control computer every two cycles (192 ms). GPS data (absolute time and position) are updated once per second. Using these data, the computer sets power levels for the motors by generating pulse width modulated (PWM) square-waves.

Flexible commanding and scheduling of the pointing system have been essential to the efficient use of limited observation time. Pointing normally follows one of several preprogrammed flight schedules. Remote (ground-based) commanding is
used to switch between schedules, to make modifications to schedules, or to take complete manual control of the pointing. In addition, remote commanding can be used to modify control loop gains, to make adjustments to sensor calibrations, and to set parameters for the CCD image processing. The FLC also generates a digital data frame for transmission to the ground. This includes sensor data and status information.

4.3.2 Pointing Control

Sensors

Feedback control is based on azimuthal rotation velocity, as measured by a rate gyroscope. Two other gyroscopes measure pitch and roll velocities but are used only for post-flight diagnostics. The gyroscopes are obtained commercially and have an accuracy of $\sim 0.01^\circ$/sec. Though they are very sensitive, the gyroscopes have substantial low frequency drifts, primarily due to ambient temperature fluctuations. Drifts are calibrated once per gondola scan period, and have little impact on pointing control.

Absolute azimuth is measured using a two axis magnetometer. The magnetometer is extremely precise ($<0.5'$) in differential measurement, but is highly non-linear due to the magnetic properties of the telescope. Pre-flight measurements are used to calibrate the magnetometer to an absolute accuracy of $\sim 1^\circ$.

Absolute elevation is measured by an optical angle encoder between the inner assembly (receiver and primary mirror) and the outer frame of the telescope. The accuracy of this measurement depends on the balancing of the telescope ($\sim 0.1^\circ$) and on long time scale pendulum motion ($\sim 0.5^\circ$, varying over tens of minutes). The differential accuracy of this elevation measurement is $\sim 1'$.

The CCD star cameras, described in Section 4.3.4, provide the most accurate measurement of telescope orientation. They are used for post-flight reconstruction rather than pointing control.
Figure 4.5: A mechanical drawing showing the layout of the telescope motors. Two motors near the top of the telescope control azimuthal orientation by driving against a reaction wheel and the cables from the balloon respectively. A linear actuator/servo-arm tilts the inner assembly, pointing the telescope in elevation. A motor below the primary mirror modulates it at relatively high speed (0.45 Hz, ±2° amplitude) in azimuth.

Motors

Three motors are used to point the MAXIMA telescope and a fourth, described in Section 4.3.3, is used to modulate the primary mirror. Two motors, located near the top of the telescope frame, are used for pointing in azimuth. One of these drives a reaction wheel with a moment of inertia of 10 kg·m² (~0.5% that of the telescope). The other torques against suspension cables connected to the balloon, which has a much greater moment of inertia than the telescope. Both motors are direct drive (ungeared) and have a torque of ~35 N·m with our maximum power of 12 Watts. The light reaction wheel provides fast response, while the other motor keeps the speed of the reaction wheel low by dumping angular momentum into the balloon. The rotational velocities of these motors are monitored by tachometers.

Elevation control is provided by a geared motor connected to a linear actuator arm. The arm is fixed between the outer assembly of the telescope frame

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\(^{a}\)BAE Systems Vibrating Structure Gyro
and the inner assembly of the receiver and primary mirror. The inner assembly is balanced about the rotation axis, so the load on this motor is very small.

The Control Loop

Motor power is determined in a digital control loop. Pointing control in azimuth and elevation are not strongly coupled and may be considered separately.

In azimuth, we use a feedback system based on the rotational velocity measured by a rate gyroscope. In CMB scans, the target velocity is constant, except during turnarounds in the scan direction. During turnarounds the target velocity varies linearly with time. The absolute position is not used directly in the control loop; it is instead used to trigger these turnarounds. This is a reasonable approach for our experiment - as long as the telescope is scanning steadily, in the correct scan region, the exact position at any given moment is unimportant.

The azimuth control algorithm determines power for two motors from three inputs, for a total of six gains. The power to the flywheel motor, $P_{fly}$, and the power to the upper motor, $P_{up}$, are given by,

$$
\begin{bmatrix}
P_{fly} \\
P_{up}
\end{bmatrix} =
\begin{bmatrix}
G_{fly,v_\omega z} & G_{fly,\dot{v}_\omega z} & G_{fly,v_{\omega fl}} \\
G_{up,v_\omega z} & G_{up,\dot{v}_\omega z} & G_{up,v_{\omega fl}}
\end{bmatrix}
\begin{bmatrix}
\dot{v}_{\omega \text{err}} \\
\frac{d}{dt} v_{\omega \text{err}} \\
v_{\text{fly}}
\end{bmatrix},
$$

(4.1)

where $v_{\omega \text{err}}$ is the target rotational velocity minus the measured rotational velocity, and $v_{\text{fly}}$ is the rotation rate of the flywheel relative to the telescope frame. $v_{\omega \text{err}}$ and $\frac{d}{dt} v_{\omega \text{err}}$ are the usual error terms in a PD control loop. $v_{\text{fly}}$ is used for two related but distinct purposes. In flywheel control, it serves as a correction for the back EMF generated by the flywheel motion. In upper motor control, it is used.
to transfer the angular momentum of the flywheel to the balloon. There is no term for upper motor velocity, because it is essentially the same as the rotational velocity of the entire telescope. At our scan speeds (up 0.3°/sec during CMB scans), back EMF in the upper motor is negligible.

In practice, several of the gain terms are not needed. Velocity based feedback systems are not stabilized by a derivative term, so both $G_{fly, \text{dv}_\text{az}}$ and $G_{up, \text{dv}_\text{az}}$ are set to zero. In addition, it is possible to set either of $G_{fly, v_{\text{az}}}$ or $G_{up, v_{\text{az}}}$ zero. For MAXIMA-I, $G_{up, v_{\text{az}}}$ was set to zero. In this configuration, the pointing is directly controlled by the flywheel motor, while the upper motor acts only to reduce the angular momentum of the flywheel. For MAXIMA-II both of these gains were finite, though $G_{fly, v_{\text{az}}}$ was relatively small. In this configuration the upper motor controlled the pointing directly, with some contribution from the flywheel. The pointing accuracy of both flights is summarized in Table 4.1. The MAXIMA-I gains are somewhat better in terms of pointing performance, though both sets of gains were adequate for our needs.\footnote{Before the MAXIMA-II flight, there was some evidence that the receiver was more susceptible than usual to vibrational noise pickup. As a safety measure, greater use was made of the slower moving upper motor.}

The elevation control formula is based on the measured angle of the telescope inner assembly relative to the outer frame. In this case, the power to the elevation drive, $P_{el}$, follows a typical PD scheme,

$$P_{el} = \left( G_{el, p_{el}} \ G_{el, \text{dp}_{el}} \right) \cdot \left( \begin{array}{c} p_{\text{el, lerr}} \\ \frac{4}{3} p_{\text{el, lerr}} \end{array} \right),$$

(4.2)

where $p_{\text{el, lerr}}$ is the difference between the target and measured elevation. A derivative term is useful for stability, so both gains are finite.

4.3.3 Primary Mirror Modulation

The primary mirror is continuously rotated from side to side about the axis indicated in Figure 2.2. The motion is a rounded triangle scan with an amplitude of $\pm 2^\circ$ and a frequency of 0.45 Hz. This modulation superimposed on that of the entire telescope yields the scan pattern in Figure 4.2 (left panel).

The mirror motion is the highest frequency modulation in the experiment.

This sets the limit below which the data are much less susceptible to 1/f noise;
because a location on the sky is revisited on time scales of 2.2 seconds, noise on longer time scales has little impact on the data. Considering sky rotation and the motion of the telescope, the actual beam overlap between consecutive mirror scans is \( \sim 99\% \).

Because the mirror rotation axis is not vertical, there is also a small modulation in elevation - a bowed pattern in which the extremes of the mirror modulation rise slightly in elevation. The elevation motion could in principle lead to a scan synchronous signal in the bolometers due to atmospheric emission. In practice, the elevation motion (<2') is not large enough to generate a detectable signal.

The mirror is actuated by a DC motor with solid state PID control electronics. The motor was obtained commercially and the control electronics were built at Berkeley. Mirror position is controlled to an accuracy of 1'.

### 4.3.4 CCD Cameras and Image Processing

We use two CCD star cameras as our main absolute pointing sensors. CCD data are not used in pointing control, but are used for post-flight pointing reconstruction. One camera, referred to as the primary or boresighted camera, is mounted on the inner telescope assembly and is boresighted with the telescope beams. The second camera, referred to as the secondary or offset camera, is positioned on the outer frame so that north celestial pole star (Polaris) is in the camera's field of view during CMB observations - approximately 40° to the right of the boresight, at a fixed elevation angle of about 31°. Camera data are processed by a devoted computer with a video digital signal processor (DSP).

CCD cameras provide an accurate and reliable direct measurement of the telescope orientation and have a number of advantages over other instruments such as magnetometers, rate gyroscopes, or differential GPS systems. Camera data are not susceptible to drifts caused by temperature fluctuations, as gyroscopes and magnetometers are. CCD cameras are more easily and precisely calibrated than magnetometers. A differential GPS receiver would have many of the advantages of a star camera, but with lower precision (>3'). While CCD data are affected by a rotational degeneracy, the small offset from the telescope beams to the boresighted camera's field of view minimizes this effect.
The boresighted camera data are used for the final pointing reconstruction and are accurate to \( \sim 0.5' \). The secondary camera is only accurate to \( \sim 15' \), due to the offset of the measurement from the telescope boresight. It is used to identify stars in the primary camera.

The main disadvantage of CCD cameras is that the limited field of view must contain a sufficiently bright star. Camera optics allow a trade-off between field of view and angular resolution. The boresighted camera has a field of view of \( 7.17^\circ \) by \( 5.50^\circ \) with a pixel size of \( 0.84' \) by \( 0.69' \). The resolution is further improved to \( \sim 0.5' \) in flight by software interpolation. The secondary camera has a larger field of view (\( 14.34^\circ \) by \( 11.00^\circ \)) and lower resolution (\( \sim 1.0' \)).

The boresighted camera reliably detects stars of \( V \) magnitude 5.0 or brighter. Stars of \( V \) magnitude 5.0 to 6.0 are detected intermittently, and stars dimmer than \( V \) magnitude 6.0 are rarely detected. This sensitivity is sufficient for all of \textsc{maxima-i} and most of \textsc{maxima-ii}. Source brightness is not an issue for the secondary camera, which always views Polaris (\( V \) magnitude = 2.0).

A small area near the center of the \textsc{maxima-ii} scan region contains no bright stars. Here we rely on heavy interpolation using data from the rate gyroscope.

**Image Rate and Phase Lag**

The DSP processes an image once every two cycles of the FLC. During CMB observations, the DSP alternately processes data from each of the two cameras (1 image per 384 ms from each camera). During planet observations, the DSP only processes data from the boresighted camera (1 image per 192 ms).

The processing time of CCD data in the DSP causes a 200-ms delay between pointing data and bolometer data. The cameras internally sample the CCD chips at 30 Hz, asynchronous to the rest of the system. This causes a jitter up to 33 nsec in the image processing delay, for an overall delay of 200±16 ms. At our gondola scan speeds, 200 ms translates to 2' to 5' of gondola rotation, which is a significant fraction of our 10' beam-size and must be taken into account. The ±16-ms jitter translates to \( \sim 0.15' \) RMS pointing uncertainty.
4.4 Pointing Reconstruction

Pointing reconstruction is based primarily on data from the CCD cameras. First, data from the secondary CCD camera are used to identify sources observed with the boresighted camera. Next, boresighted camera data are used to find the pointing solution. Where camera data are unavailable, a combination of rate gyroscope data and numerical interpolation is applied. Finally, the effect of the primary mirror modulation is included. These four steps, and their contributions to the final pointing uncertainty, are described here.

4.4.1 CMB Scan Reconstruction

Figure 4.6: Pointing Reconstruction for the four MAXIMA CMB observations, two in each of MAXIMA-I and MAXIMA-II. Plots A, B, C, D show the scan pattern of the telescope, without the effect of the primary mirror modulation. Plots E, F, G, H show the same scan patterns with the addition of the 4°, 0.45-Hz primary mirror modulation. Plots E, F, G, H include small gaps, corresponding to subdivisions in the data that coincide with internal calibration events.
Step 1: Identification of Guide Stars

The first stage in pointing reconstruction is the identification of the stars observed in the boresighted CCD camera. The camera acquires data from the two brightest stars in its field of view at any given time. During a CMB scan, several thousand detections are made of 16 to 40 different stars of magnitude 6.0 or brighter (Table 4.2).

In principle, guide stars can be identified from their relative positions, combined with data from the magnetometer and the telescope elevation angle. The accuracy of this method is limited by the magnetometer calibration, the lack of simultaneous measurements of many guide stars, and the slow pendulum motion of the telescope. Instead, we identify stars using data from the secondary CCD camera. The secondary camera is offset from the boresight so that a known star, Polaris, is constantly in its field of view during CMB observations. This provides boresight pointing information with an accuracy of ~15' which is sufficient to identify >95% of the stars detected in the boresighted camera. The majority of unidentified “stars” are noise in the CCD/DSP system.

Step 2: Reconstruction Based on Known Guide Stars

Up to two stars are located from each image processed by the boresighted CCD camera. These data are the basis of the pointing reconstruction. Pointing is determined at each time sample from each known star using a series of standard coordinate transformations.

Often, pointing can be determined simultaneously from two different stars in the same field of view. In these cases the discrepancy between the two pointing solutions is used to estimate the overall error in the CCD measurement. The typical discrepancy is ~0.5' (Table 4.3).

Step 3: Interpolation

Data are gathered from the boresighted camera once every 384 ms during CMB scans, during which time the telescope moves ~8'. Star data are obtained from most, but not all, CCD images. In rare cases, no stars are detected by the camera for up to eight seconds at a time, making accurate interpolation more difficult. Near
<table>
<thead>
<tr>
<th>CMB Observation</th>
<th>MAXIMA-I Scan 1</th>
<th>MAXIMA-I Scan 2</th>
<th>MAXIMA-II Scan 1</th>
<th>MAXIMA-II Scan 2</th>
</tr>
</thead>
<tbody>
<tr>
<td>Number of Stars</td>
<td>16</td>
<td>16</td>
<td>39</td>
<td>30</td>
</tr>
<tr>
<td>Frames w/ 2 Stars</td>
<td>6929</td>
<td>6185</td>
<td>10627</td>
<td>6077</td>
</tr>
<tr>
<td>Frames w/ 1 Stars</td>
<td>4977</td>
<td>4879</td>
<td>7504</td>
<td>7854</td>
</tr>
<tr>
<td>Frames w/ 0 Stars</td>
<td>3601</td>
<td>2490</td>
<td>4821</td>
<td>6339</td>
</tr>
</tbody>
</table>

Table 4.2: The top row is the number of guide stars used for pointing reconstruction of each MAXIMA CMB scan. The next three rows give the number of times 0, 1, or 2 guide stars were found in a CCD camera image.

the center of the MAXIMA-II scan, there is a region with very few guide stars where interpolation must be used more often over longer time scales.

The initial interpolation is from intermittent CCD data (384 ms or longer between updates) to the consistent 96 ms period of the control electronics. In the azimuth, data from the rate gyroscope are integrated to find position. Because the gyroscope is calibrated continuously from the azimuthal CMB scan, this process is very accurate. In the elevation direction, the rate gyroscopes are more difficult to calibrate. However, the motion of the telescope in elevation is extremely slow and small, so these data are safely interpolated numerically.

A second interpolation from the 96-ms period to the 4.8-ms period of the bolometer data is purely numerical. On these shorter time scales, the telescope beams move much less (<2' during CMB scans) and there is no danger of introducing significant pointing error.

As a test of accuracy, the azimuth data are reinterpolated numerically and compared to the gyroscope-based interpolation. The difference between the two is used to estimate the pointing error introduced by interpolation. Though the RMS discrepancy between these two methods is very small, the distribution has extreme outliers corresponding to regions of the sky with few bright stars. Interpolated regions with a difference of greater than 3.3' are not used in data analysis.
Step 4: Primary Mirror Modulation

In the final stage in pointing reconstruction we add the effect of the primary mirror modulation. The angle of the primary mirror is measured to several arcseconds by a linear variable differential transformer (LVDT). The LVDT data are calibrated both before flight and during flight using data from the planet observation.

The motion of the primary mirror moves the telescope beams primarily in azimuth. However, there is a small motion in elevation which is much more difficult to calibrate and is a source of pointing uncertainty. The elevation motion depends upon the zero position of the mirror. This is measured to $\sim 1\degree$, which leads to a conservative pointing uncertainty of about $0.8\arcmin$ RMS.

The absolute boresight offset of each detector, as measured from planet observations, is included in the pointing reconstruction at this stage.

4.4.2 Pointing Uncertainty

The $0.8\arcmin$ uncertainty of the primary mirror modulation is the largest error term in the MAXIMA pointing solution. Though purely systematic, the scan pattern and cross-linking tend to blur out the effect. In addition, other sources of pointing uncertainty further randomize the total error. The overall pointing error is approximated as a $1\arcmin$ gaussian blur. Simulations show that the effect of such a pointing error on the angular power spectrum is a 10% reduction at $\ell = 1000$ and that this reduction scales roughly as $\ell^2$.

Pointing uncertainty is a subdominant source of CMB power spectrum error at all values of $\ell$. While it is possible to compensate the power spectrum for the reduction caused by pointing error, we have not done so because it is relatively small, and because our model of the pointing error as gaussian is not exact.

4.4.3 Planet Scan Reconstruction

Reconstruction of planet scans is very similar to that of CMB scans with several simplifications. Because there is always a known bright source, the planet itself, in the telescope boresight, source identification is not an issue. Data are acquired every 192 ms, and a source is found in every image, allowing easy and accurate interpolation.
<table>
<thead>
<tr>
<th>Random Errors</th>
<th>Systematic Errors</th>
</tr>
</thead>
<tbody>
<tr>
<td>CCD Camera</td>
<td>Interpolation</td>
</tr>
<tr>
<td>Camera Timing</td>
<td>(see caption)</td>
</tr>
</tbody>
</table>

| MAXIMA-I | 0.46' | 0.15' | <0.001' | 0.25' | 0.81' |
| MAXIMA-II | 0.58' | " | (see caption) | " | " |

Table 4.3: Sources of error in pointing reconstruction. “CCD Camera” is the error in determining the coordinates of the guide star in the CCD image. “Camera Timing” is the effect of the timing uncertainty of image acquisition. “Interpolation” is the error caused by interpolation over periods without camera data. This is negligible for all of MAXIMA-I and most of MAXIMA-II. However, for ~20% of the MAXIMA-II scan region there are very few stars, increasing the use of interpolation and raising the rms error to ~1'. “Detector offset” is the zero-position uncertainty of the telescope beams in azimuth and elevation. “Primary Modulation” is the uncertainty from the rotation of the primary mirror.

The addition of primary mirror modulation and channel specific boresight offsets is handled somewhat differently than in CMB scans. The calibration of the primary mirror modulator and the offset of each beam position are taken as parameters. These parameters, as well as bolometer time constants, are fit to the bolometer data using the pointing solution and the known planet position, calibrating both the primary mirror motion and the detectors’ spatial offsets (Winant (2003)).

### 4.4.4 CMB Dipole Scan Reconstruction

During CMB dipole scans, the telescope is fully rotated in azimuth at 20°/sec (18 second period). At this speed the CCD cameras provide no meaningful information. Data from the rate gyroscopes and the magnetometer, as well as CCD data before and after the dipole observation, are used to reconstruct pointing.

These data yield a pointing reconstruction roughly one order of magnitude less accurate than that of the CMB and planet observations (<10' error). This is sufficient for the dipole measurement, causing an error of ~0.1% in the absolute calibration. The pointing is confirmed using the strong, localized signal of the Galactic plane in the bolometers.
Chapter 5

Calibration

In this chapter we discuss the calibration of detector responsivity. Section 5.1 defines responsivity and the maxima calibration strategy. Sections 5.2 and 5.3 cover the absolute calibrations from the CMB dipole and planets. Section 5.4 deals with the time dependence of the calibration. Section 5.5 discusses the use of all these data to determine the overall calibration. Section 5.6 describes pre-flight responsivity testing.

See Appendix D for a complete summary of calibration parameters and uncertainties for both maxima flights.

5.1 Definition and Overview

Detector data are recorded in units of bolometer voltage. The ratio \( \frac{\Delta v_{\text{detector}}}{\Delta v_{\text{cm8}}} \) is referred to as the responsivity. Measurement of this value using known signals is referred to as responsivity calibration.

Responsivity is strictly a function of both the angular scale and the temporal frequency of the observed signal. The finite resolution of the experiment leads to an effective reduction of responsivity for features on small angular scales. This reduction is referred to as the beam window function or \( B_{\ell} \). \( B_{\ell} \) is unity at small \( \ell \), as for the dipole calibration, and drops off at high \( \ell \). For the planet calibration, the "beam dilution" factor (§5.3.1) accounts for the beam function. For CMB data analysis, the explicit \( B_{\ell} \) is applied to the angular power spectrum (§6.6).

Similarly, temporal filters (e.g. bolometer time constants, electronic filters)
make responsivity a function of frequency. These filters are deconvolved from the
data in the early stages of analysis; the responsivity of the deconvolved data is not
a function of frequency.

Absolute responsivity calibration uses two known sources during each MAX-
IMA flight: the CMB dipole and a planet. Measurements of the CMB dipole give
the best absolute calibration for the 150-GHz and 230-GHz detectors. Observations
of planets (Jupiter in MAXIMA-I and Mars in MAXIMA-II) are used to calibrate the
410-GHz detectors, and are used as a consistency check for the dipole calibration.
Observations of planets are also used to measure the size and shape of the telescope
beams. This is not discussed here, but may be found in Winant (2003).

Responsivity depends on bolometer properties, optical loading, electrical
bias power, and the temperature of the bolometer thermal reservoir. The ther-
mal reservoir temperature, controlled by the adiabatic demagnetization refrigerator,
varies significantly causing responsivity fluctuations. An internal millimeter wave
source (stimulator) is used to periodically measure responsivity changes.

The MAXIMA-I data have a calibration error of 4%, while the MAXIMA-
II data have a calibration error of 3%. These are the most accurate calibrations
achieved by any sub-orbital CMB experiment.

5.2 CMB Dipole

The CMB dipole is the Doppler shift in the observed CMB temperature
resulting from the motion of the Earth relative to the CMB rest frame. The dipole
amplitude of $3.358 \pm 0.023$ mK has been measured by the COBE DMR (Lineweaver et al.
(1996)). The dipole is the main calibrator for the 150-GHz and 230-GHz detectors.
It has two main advantages over point source calibrators. The first is the optical
spectrum: because the dipole is a small Doppler variation on the 2.725 K of the CMB,
it has exactly the same spectral profile as the CMB anisotropy. Uncertainties in the
detectors' spectral response do not affect the calibration. The second advantage of
dipole calibration is the angular scale of the signal. Because the dipole is $\sim 1000$
times larger than our telescope beams, uncertainties in the beam window function
do not affect the calibration.

The 410-GHz detectors, used to confirm the absence of Galactic dust and
<table>
<thead>
<tr>
<th></th>
<th>Dipole Amplitude</th>
<th>Dipole Elevation</th>
<th>Observation Elevation</th>
<th>Observed Amplitude</th>
</tr>
</thead>
<tbody>
<tr>
<td>MAXIMA-I</td>
<td>3.195 mK</td>
<td>20°</td>
<td>50°</td>
<td>2.04 mK</td>
</tr>
<tr>
<td>MAXIMA-II</td>
<td>3.010 mK</td>
<td>48°</td>
<td>32°</td>
<td>1.15 mK</td>
</tr>
</tbody>
</table>

Table 5.1: The CMB dipole signals measured in each MAXIMA flight. “Dipole Amplitude” includes the effect of the Earth’s motion around the Sun. “Dipole Elevation” gives the angle from the dipole direction to the horizon during the observation. For our observing pattern (azimuthal rotation) this would ideally be 0°. “Observation Elevation” is the constant elevation at which the telescope was rotated to observe the dipole. The “Observed Amplitude” is the amplitude of the dipole over the region of the scan. This varies slightly over the course of the observation, due to the rotation of the sky.

Atmospheric foregrounds in our data, are deliberately much less responsive to the CMB spectrum. Dipole calibrations for these channels have very low signal-to-noise. Planet observations are used to calibrate the 410-GHz detectors.

5.2.1 Dipole Observations

Dipole observations are carried out by rotating the telescope in azimuth with a azimuthal angular velocity of 20°/sec. The observed signal from the dipole is a sine wave at 55 mHz.

The observed signal varies from the canonical dipole amplitude for two reasons. First, the dipole fluctuates annually due to the motion of the Earth around the Sun. Second, the observation is a circular pattern at a fixed elevation, and does not span the full extent of the dipole. The observed signals are listed in Table 5.1.

Parasitic Signals

The rotating scan pattern of the dipole observation is sensitive to parasitic signals from Galactic dust and in some cases the atmosphere. The Galactic dust signal is modeled from frequency extrapolations of published maps (Jaffe et al. (1999), Schlegel et al. (1998)). The dust signal is much smaller than the dipole signal at 150 GHz, except near the Galactic plane. Data within 5° of the Galactic plane are neglected in data analysis. Elsewhere, the dust model is fit to the data along with
the dipole model. Overall normalization is taken as a free parameter to account for uncertainties in the frequency extrapolation of the dust signal. In practice, the dust model does not affect the dipole calibration due to its low amplitude and lack of a dipole-like spatial component.

An additional small signal, believed to be atmospheric, is observed in the beginning of the MAXIMA-I dipole calibration. In MAXIMA-I we began the dipole observation near the beginning of the flight, while the telescope was still ascending from ~21.5 km to the final observing altitude of ~38.5 km. The additional signal was observed during the first third of the observation (altitude <30 km). We believe that this signal is atmospheric for four reasons: 1, it is highly correlated in all the optical bolometers; 2, it is spectrally consistent with atmospheric emission, being larger for the higher frequency detectors; 3, it is spatially stable on the scale of a few minutes, but varies on longer scales; 4, the magnitude of the signal declines steadily with altitude.

The atmospheric signal is corrected using data from the 410-GHz bolometers. These data, which are relatively insensitive to the dipole and sensitive to the parasitic signal, are used as a template for the parasitic signal in the 150-GHz and 230-GHz data. A correction is applied for the CMB sensitivity of the 410-GHz detectors, as calibrated by planet observations. As with Galactic dust, we find that fitting the believed atmospheric signal does not affect our final dipole calibration values. It does slightly increase calibration uncertainty, because of noise in the 410-GHz data used to 'model' the parasitic signal.

5.2.2 Dipole Data Analysis

During each flight, the dipole was observed for ~30 minutes (100 rotations). For each detector, the effects of electronic filters and bolometer time constants are first deconvolved from the entire data stream. Data from each rotation are then fit independently according to the model,

\[ T_{\text{detector}} = (A \ast T_{CMB, Model}) + (B \ast T_{dust}) + (C \ast N_{Drift}) + D, \]  

(5.1)

in which \( T_{\text{detector}} \) is time stream of detector data in voltage units, \( T_{CMB, Model} \) is the CMB dipole model in units of temperature contrast, \( T_{dust} \) is the Galactic dust model, \( N_{Drift} \) is linear drift, and \( A, B, C, D \) are fitting constants. \( A \) is the calibration of
Figure 5.1: MAXIMA-ii 150-GHz Dipole Data and Fit. Top panel: The top trace is the data from a 150-GHz bolometer during observations of the CMB dipole. An overall gradient has been removed and the offset is arbitrary. The sinusoidal signal is the CMB dipole modulated by the rotation of the telescope (~18-second period). The large periodic spikes are caused by intense dust signals near the Galactic plane. The lower trace is a model curve, with amplitude fitted to the data, including the CMB dipole and a Galactic dust map. Bottom panel: The difference between the model and the fit in the top panel are shown. The model deviates from the data near the Galactic plane crossing. These highly localized signals are not well fit with ~10' pointing reconstruction accuracy of the dipole observation.
the detector to CMB signals. Data collected within 5° of the Galactic plane are not used for fitting because the pointing reconstruction of the dipole observation is not accurate enough for localized features.

Figure 5.2: MAXIMA-I 150-GHz Dipole Data and Fit. Similar to the MAXIMA-II data in the top panel of Figure 5.1. The top trace is raw data with arbitrary offset, while the bottom trace is a model including dipole, dust, and a linear drift. In this case, the model curve also includes an additional term based on 410-GHz data to account for the believed atmospheric signal observed at low altitude (the first ~1/3 of the data). This is why the model curve is not noiseless.

For MAXIMA-I the believed atmospheric signal is taken into account by modifying the fit to,

\[ T_{detector} = (A \times T_{CMB,Model}) + (B \times T_{dust}) + (C \times N_{Drift}) + D + (E \times T_{410}). \]  \( (5.2) \)

\( T_{410} \) is time stream data from a 410-GHz detector and \( E \) is an additional fitting parameter. Because the 410-GHz data do have some very small sensitivity to the CMB, \( A \) is no longer an unbiased calibration. To account for this, the 410-GHz data are first calibrated using planet data and the parameter \( A \) is corrected,

\[ A' = A - (E \times Cal_{410}). \]  \( (5.3) \)

\( Cal_{410} \) is the planet-based calibration of the 410-GHz data and \( A' \) is the true calibration of the low frequency channel. In practice the 410-GHz term doesn’t affect calibration values by more than 0.5 \( \sigma \), though it does increase uncertainties. The correction from \( A \) to \( A' \) has a negligible effect on both the calibration and the calibration uncertainty due to the small value of the \( E \) parameter.

Each rotation yields a calibration value (\( A \) or \( A' \) above) and an associated error range. These are combined statistically, with 2-\( \sigma \) outliers excluded. Data from ~80 rotations are analyzed from each flight, with 2 to 8 excluded as outliers for each detector.
5.2.3 Dipole Calibration Error Sources

Detector Noise

Dipole calibration uncertainty (1-4\%) is dominated by detector noise at the dipole observation frequency of 56 mHz. The fitting routine is found numerically to reject noise beyond a fractional bandwidth of \(~0.5\). Noise is effectively reduced by a further factor of \(\sqrt{2}\) by the known phase of the dipole model. Raw detector noise near 56 mHz for a 150-GHz bolometer is typically 150 nV Hz\(^{-0.5}\).\(^1\) Considering bandwidth and phase constraints, this yields an expected \(~20\)-nV noise level for a single dipole fit. For a typical 150-GHz detector the amplitude of the dipole response is \(~70\) nV. We therefore expect a statistical uncertainty of \(~30\)% from a single rotation.

Detector noise is the only source of statistical uncertainty in the calibration, and can be estimated directly from the scatter of the individual, single rotation calibrations. Such analysis yields single rotation statistical uncertainties of 10\% to 30\% for 150-GHz detectors in either flight. These numbers are somewhat lower than predicted due to imperfect understanding of detector noise at very low frequencies.

An integration of 80 to 90 dipole observations per flight provides a total statistical uncertainty of 1.4\% to 4.2\% for 150-GHz detectors in maxima-i and 1.1\% to 2.5\% in maxima-ii. Note that when combining data from multiple detectors, we use the highest statistical uncertainty of the combined channels.

For the beginning of the maxima-i dipole scan, a parasitic signal is seen, and is modeled as described above. This modeling is based on data from a 410-GHz detector that contributes additional detector noise to the calibration uncertainty. Because the coefficient \(E\) in Equation 5.2 is small, the noise contribution from the 410-GHz detector is suppressed and increases the final calibration uncertainty by only about 0.5\%.

Because the CMB responsivity of the 230-GHz detectors is 60-70\% that of the 150-GHz detectors, they have a proportionally higher statistical error.

\(^1\)In fact, detector noise performance at these frequencies below the 1/f knee (~0.5 Hz) is much less consistent than at high frequencies, and can vary by a factor of four or more between detectors.
<table>
<thead>
<tr>
<th>Frequency</th>
<th>Dipole Responsivity Change (MAXIMA-I/2)</th>
<th>Planet Responsivity Change (MAXIMA-I/2)</th>
<th>Stimulator Responsivity Change (MAXIMA-I/2)</th>
</tr>
</thead>
<tbody>
<tr>
<td>150-GHz</td>
<td>0.05%/&lt;0.05%</td>
<td>0.7-2.5%/0.2-0.5%</td>
<td>0.1-0.5%/0.1-1.5%</td>
</tr>
<tr>
<td>230-GHz</td>
<td>0.05%/&lt;0.05%</td>
<td>1.0-6.1%/0.2-0.4%</td>
<td>0.2-4.0%/0.1-0.5%</td>
</tr>
<tr>
<td>410-GHz</td>
<td>0.05%/&lt;0.05%</td>
<td>1.9-7.7%/0.4-1.0%</td>
<td>0.5-3.3%/0.3-2.7%</td>
</tr>
</tbody>
</table>

Table 5.2: A summary of detector response linearity during observations of various calibration sources. “Responsivity Change” is the fractional change in detector responsivity due to the optical load from the calibrator. For the dipole calibration there was no responsivity change within noise limits. For the planet and stimulator calibrations, quoted values are derived from the maximum of the signal. Stimulator power was reduced in MAXIMA-II to increase linearity. The impact of these responsivity changes on calibration accuracy is discussed in the text. These values are derived from measured changes in bolometer resistance, as described in Appendix B.

Other Error Sources

In addition to statistical uncertainty, there are a number of known systematic effects, though none have a significant impact on the calibration. First, dipole pointing reconstruction (§4.4.4) is accurate to ~10' and contributes negligibly (~0.1%) to the calibration uncertainty. Second, the characterization of the high pass filter in the bolometer readout is accurate to ~0.3%. Third, the dipole model derived from the COBE measurement is accurate to 0.68%. Finally, the signal from the dipole is small enough that bolometer saturation is negligible (Table 5.2, Appendix B).

5.3 Planets

In each flight observations are made of a planet: Jupiter in MAXIMA-I and Mars in MAXIMA-II. These data serve several purposes. They are used to measure optical parameters of the telescope, including the size, shape, and absolute position of each beam, and to calibrate the primary mirror modulation. They are also used to measure the electronic filters and bolometer time constants. These measurements are described in Winant (2003).
Figure 5.3: Data from observations of Mars and Jupiter. **Left Top:** Raw data from the entire MAXIMA-1 Jupiter observation for a 150-GHz detector. Each of the closely spaced vertical lines is a single pass of the planet. The modulation of the envelope is caused by the drift of the planet in elevation. The apparently non-gaussian shape of the envelope is caused by known pointing control imperfections. **Right Top:** Raw data from a single pass of Jupiter for a 150-GHz detector. These data are an expanded view of one of the vertical spikes in the plot on the left. The scan speed is determined by the modulation of the primary mirror. Signal-to-noise is ~1000. The solid line is a gaussian fit. The ‘bump’ on the right side of the plot is caused by bolometer time constants and electronic filters. Deconvolution of these effects removes the bump. **Left and Right Bottom:** As above, for a MAXIMA-II Mars observation. The signal-to-noise ratio is ~150.
Here we discuss the use of planet data for responsivity calibration. Planet data are the only absolute calibration source for the 410-GHz detectors. They are also used to confirm the dipole calibration of the 150-GHz and 230-GHz detectors.

5.3.1 Planet Data

During planet observations, the telescope tracks the planet in azimuth while remaining at fixed elevation. As the planet drifts through the elevation of the observation, the modulation of the primary mirror “slices” the beams across the planet many times. As the planet drifts in elevation the spatial response of each beam is measured in two dimensions.

Responsivity calibration is obtained from the maximum voltage response.
Expected signals are derived from published measurements and models of planet temperatures and emissivities (Goldin et al. (1997)), combined with the spectral response of the detectors.

A correction is applied for beam dilution - the fraction of the telescope beam filled by the planet. Jupiter had an angular diameter of 46.5" during MAXIMA-I and Mars had an angular diameter 12.7" during MAXIMA-II. The beam dilution is the integral of the spatial response of the detector over the area of the planet, normalized by the integral of the entire spatial response. The dilution factors vary from $3.4 \times 10^{-3}$ to $4.4 \times 10^{-3}$ for MAXIMA-I and $3.1 \times 10^{-4}$ to $4.4 \times 10^{-4}$ for MAXIMA-II. This correction is the dominant error source for the planet calibration in both flights.

An additional correction of roughly 5% for MAXIMA-I and 1% for MAXIMA-II is applied for the reduction in responsivity caused by the optical load from the planet (Table 5.2). This effect was neglected in the initial MAXIMA-I data analysis and caused an apparent small systematic discrepancy between the dipole and Jupiter calibrations. This discrepancy was within the error range of the Jupiter calibration and does not affect the CMB map or power spectrum.

### 5.3.2 Planet Calibration Error Sources

The dominant error term for the planet calibration is the uncertainty in the beam dilution factor. The uncertainty in the integrated beam response is 5% to 10%. In addition, there is a possibility of small, broad side-lobes that are not measured in the beam maps. We assign an uncertainty of 10% from beam shape errors. Beam shape error, especially that due to broad side-lobes, is partially correlated between detectors because of their shared optics.

Uncertainties in the effective brightness temperature of the planets contribute 5% to calibration error. The brightness temperature of Mars has been modeled to this accuracy, both by extrapolations from high frequency observations (Wright & Odenwald (1980)) and by physical modeling (Rudy (1987)). The atmospheric properties for Jupiter make modeling relatively difficult. Our expected Jupiter signal is based on published brightness ratios between Jupiter and Mars (Goldin et al. (1997)). The planet temperature uncertainty is fully correlated between all the detectors.
Measurements of the detector spectra contribute 1-2% error at 150 GHz, 3-7% at 230 GHz, and 2-3% at 410 GHz. Measurements of the peak planet voltage contribute 1-4% error; one detector in MAXIMA-II was anomalously noisy, increasing this term to ~10%. Uncertainty in the bolometer saturation is negligible (Table 5.2).

5.4 Time Dependent Calibration

![Plot](image)

Figure 5.5: An internal relative calibration event. The voltage response of a 150-GHz detector in MAXIMA-II to the stimulator lamp is shown. Power to the stimulator is constant for 0.2 sec < t < 9.5 sec and is zero elsewhere. The rounding of the response is caused by the stimulator on/off time constant, which is ~100 times longer than the bolometer time constant. The sloping of the signal in the ‘on’ state and the slow settling of the baseline after the event are caused by a weak high pass filter in the bolometer readout electronics.

The responsivity of bolometers varies with their operating temperature. A temperature increase of 1 mK in the thermal reservoir leads to a responsivity reduction of 1-2%. The temperature varied by ~6 mK over the course of data collection in MAXIMA-I and by ~21 mK in MAXIMA-II. The larger variation in MAXIMA-II was due partly to the length of the flight, and partly to technical difficulties with the adiabatic demagnetization refrigerator.

Responsivity variations are monitored using the internal millimeter wave source (stimulator) described in Section 2.5.2. The stimulator is activated for 10-sec
Figure 5.6: Temperature dependence of the responsivity of a 150-GHz detector. These data were collected during maxima-II. The large range of temperature was due to the length of the flight and a partial failure of the ADR. Points at 0.105 K and 0.107 K are omitted due to high noise. The point near 0.127 K, measured shortly after sunrise, shows less responsivity than would be expected from the nighttime data. During maxima-I, the temperature of the thermal reservoir was more stable, varying from 98-104 mK over the course of data collection.

periods, once every 20 minutes during the flight. The signal in a 150-GHz detector from a stimulator event is shown in Figure 5.5. The response to the stimulator signal in the 230-GHz and 410-GHz detectors is more than twice that in the 150-GHz detectors. In addition, the stimulator location is asymmetric with respect to the detector array, with the four closest detectors having twice the response of the four farthest ones.

To obtain relative calibration values from stimulator events, we begin by subtracting an overall gradient from each event to remove the effects of detector drift. We then perform a linear fit between pairs of stimulator events. The slope of this fit is the calibration ratio between the events, while the offset of the fit is simply an offset in the detector data.

Once the relative calibration at each stimulator event is known, we fit the values to a linear function of the temperature of the bolometer thermal reservoir (Figure 5.6). This fit is combined with the absolute calibration to obtain the overall
calibration as a function of time throughout the flight (§5.5).

5.4.1 Relative Calibration Error Analysis

The relative calibration between stimulator events is affected by random variations (detector noise or stimulator instability), but is not affected by systematics that are consistent between stimulator events. Uncertainties in the spectra of the detectors and the beam filling of the stimulator signal are purely stable. The reduction in bolometer responsivity due to the large optical load of the stimulator is nearly stable (Table 5.2). Though it does vary with bolometer temperature, this variation contributes less than 0.1% error to the relative calibration.

Random errors in the comparison of stimulator events are 1-2%. Instabilities in the stimulator current account for <0.5% of this, while detector noise accounts for the rest.

We treat the relative calibration as a linear function of temperature and take the optical load during the flight as constant. These assumptions are supported by bolometer models (Winant (2003), Grannan et al. (1997)) and contribute negligibly to calibration error.

5.5 Combined Calibration

The overall calibration for each detector was obtained by combining an absolute calibrator with the relative calibration. The absolute calibrator is the CMB dipole for 150-GHz and 230-GHz detectors and is the planet scan for 410-GHz detectors. The relative calibration is based on the temperature of the bolometer thermal reservoir and the responsivity-temperature relation obtained in Section 5.4. Temperature is monitored continuously.

The overall calibration error is the combined error from the absolute calibrator and the relative calibration. Relative calibration error varies over the course of the flight. Quoted values are based on averages over the CMB observations. Relative calibration error is subdominant for MAXIMA-II (1% to 2%) and negligible for MAXIMA-I (<0.1%).

The published MAXIMA-I data are conservatively assigned the highest calibration uncertainty of the detectors used, 4%.
5.6 Pre-flight Responsivity Tests

Rough measurements of the bolometer responsivity were made before each MAXIMA flight as a diagnostic of receiver performance. For these measurements, an optical load at 273 K (0° C) is placed at the entrance of the receiver optics. A spinning fan blade coated with millimeter wave emitting material at ~300 K (room temperature) periodically blocks the optics from the colder load. This setup gives a ~28-K chopped signal against a ~300-K background at receiver entrance. An internal neutral density filter reduces chopped signal and external loading by a factor of 100.

In principle, stimulator data taken on the ground can be used to transfer pre-flight responsivity tests to in-flight calibration. In practice, the errors in this procedure are far larger than those for either the dipole or planet calibrations. Uncertainties in the emissivity of the load, the temperature contrast, and the transmittance of the neutral density filter are more than 30%. While a determined effort might reduce this uncertainty somewhat, reaching the accuracy of the in-flight calibrations would be impractical or impossible.
Chapter 6

Data Processing and Analysis

This chapter describes the processing and analysis of MAXIMA data to obtain CMB maps and power spectra. It begins with an overview (§6.1), followed by a look at the key steps (§6.2 - 6.6). Pointing reconstruction and calibration have already been described in Chapters 4 and 5. Results are presented in Chapter 7. Tests of systematic effects at various stages are described in Chapter 8.

References of particular interest are Stompor et al. (2002a), which deals with noise estimation and map making using MAXIMA-I data as an example, and Borrill (1999), which describes MADCAP, a software package used in MAXIMA data analysis.

6.1 Introduction

The process of extracting cosmological information from experimental data consists of two general stages. First is the reduction of raw data into time ordered pointing, calibration, and detector data. During this data reduction, experimental details are accounted for, low quality data are removed, calibration is determined, telemetry glitches and cosmic ray hits are identified, and a pointing solution is found. Data with glitches or poorly determined pointing are flagged. These tasks are described in Chapters 4 (Pointing Reconstruction) and 5 ( Calibration) and in Section 6.2 in this chapter (Detector Data Preparation).

The second general stage of data treatment, converting time-ordered data into CMB maps and power spectra, is less dependent on details of the experiment
Table 6.1: The total number of measurements and the number of time stream subdivisions for each CMB observation.

<table>
<thead>
<tr>
<th>Flight</th>
<th>CMB Scan</th>
<th>Total Measurements</th>
<th>Time Stream Segments</th>
</tr>
</thead>
<tbody>
<tr>
<td>MAXIMA-I</td>
<td>1</td>
<td>1,060,050</td>
<td>11</td>
</tr>
<tr>
<td>MAXIMA-I</td>
<td>2</td>
<td>917,900</td>
<td>10</td>
</tr>
<tr>
<td>MAXIMA-II</td>
<td>1</td>
<td>1,477,146</td>
<td>6</td>
</tr>
<tr>
<td>MAXIMA-II</td>
<td>2</td>
<td>1,412,874</td>
<td>6</td>
</tr>
</tbody>
</table>

and is more numerically intensive. First, the detector noise spectrum is estimated, calibration is applied, and glitches in the data are replaced with unbiased noise (§6.4). Then, the time-ordered data along with the time domain noise estimate are used to produce a CMB map and spatial noise correlation matrix (§6.5). Finally, the map and noise correlations are used to estimate the power spectrum of the CMB (§6.6). More detailed discussions, and treatments of concerns beyond the scope of this chapter, can be found in Stompor et al. (2002a), Borrill (1999), and Bond et al. (1998).

6.2 Bolometer Data Preparation

The CMB data consist of approximately $2 \times 10^6$ measurements per detector for MAXIMA-I and $3 \times 10^6$ measurements per detector for MAXIMA-II. The raw output of the experiment was approximately $4 \times 10^6$ measurements for MAXIMA-I and $9 \times 10^6$ measurements for MAXIMA-II, including CMB data, calibration, and test data.

The first stage of data preparation is to isolate the detector time streams corresponding to the CMB observations and to divide them into shorter segments (Table 6.1). Internal relative calibration events cause a very large signal in the optical channels, and are a natural break-point between time stream segments. These events occur at intervals of roughly 250,000 measurements (20 minutes); this is the maximum length of any time stream segment. The time streams are also divided when the scan pattern of the telescope is changed. Further subdivisions are added with the requirement that noise properties be approximately stationary within each
segment. The shortest segments are roughly $3 \times 10^4$ elements in MAXIMA-i and $1 \times 10^5$ elements in MAXIMA-II. Gaps between segments are at least $2 \times 10^4$ elements ($\sim 100$ sec). This gap length, combined with an electronic 15-mHz highpass filter, eliminates significant noise correlations between segments.

Within each time stream segment, overall offset, gradient, and quadratic components are subtracted. Electronic filters and bolometer time constants will be deconvolved in a later stage of data analysis.

Measurements compromised by 'glitches' - short transients such as cosmic ray hits and telemetry drop outs - are flagged and excluded from data analysis. Typically, $\sim 2\%$ of data are flagged as glitches. The gaps left in the time stream are short (typically $\sim 10$ measurements) and frequent; the noise on either side cannot be considered uncorrelated. Treatment of these gaps is discussed in Section 6.4.

6.3 Composition of Bolometer Data

The time ordered detector data are modeled as a sum of sky signal, noise, and a parasitic signal synchronous with the modulation of the primary mirror. Sky signals are filtered by the detector time constants and electronics. Various noise components are subject to some or all of these filters.\(^1\) Mirror scan synchronous noise is assumed to be subject to all filters. Denoting the raw time stream for a single detector for a single segment of a flight as $d_F$, the $i$\textsuperscript{th} time sample $d_F(i)$ within these data is

$$d_F(i) = \sum_j F(i, j)[t_{\text{sky}}(\gamma(j)) + x(\alpha(j))] + n_t(i). \quad (6.1)$$

$F(i, j)$ is the combined electronic and time constant filter, $t_{\text{sky}}$ is the temperature fluctuation of the CMB (and foregrounds) as a function of position, $\gamma$ is the pointing solution as a function time, $x$ is the mirror scan synchronous signal as a function of mirror orientation, $\alpha$ is the orientation of the mirror as a function of time, and $n_t(i)$ is the total time stream noise including applicable filters.

\(^1\)Johnson noise, for example, is subject to electronic filters, but not time constants, while photon noise is subject to all the filters. In practice these difference are not important to data analysis.
6.4 Noise Estimation, Gap Filling, and Filter Deconvolution

Once the deglitched time ordered data are available, the frequency power spectrum of the time domain noise is estimated. This estimate is used for map making (§6.5), and also to restore time stream continuity over the short gaps left by deglitching. After noise estimation, the effects of instrumental filters are removed from the data and the calibration is applied.

6.4.1 Time Ordered Data Power Spectrum

The noise power spectrum $P_n(f)$ is defined as $|\tilde{n}_i|^2$. $P_n(f)$ for MAXIMA consists of approximately white noise above $\sim0.5$ Hz. At lower frequencies the noise power spectrum is described by a power law, $1/f^n$, with $1.0<n<2.5$. Above 20 Hz, the noise power spectrum drops rapidly due to the effects of an electronic filter.

The power spectrum $P(f)$ of the full data (signal + noise) is similar to that of the noise. The sky signal is very small compared to the time domain noise, but in some MAXIMA-I channels and most MAXIMA-II channels, the mirror scan synchronous signal is significant, causing peaks in the power spectrum at the frequency of the scan and its first harmonic (0.45 Hz and 0.90 Hz).

6.4.2 Estimation of Signal Free Noise

Noise estimation is simplest under the assumption of a highly noise dominated time stream, free from both CMB and parasitic signals. In this case, noise estimation consists of five steps.

The first step is prewhitening, which reduces the noise correlation length. The data are convolved with a filter selected to yield constant $P(f)$ for $f \to 0$. Gaps in the convolved data are widened to account for the width of the filter. The second step is the estimation of the power spectrum for continuous blocks of data, using standard methods (Press et al. (1992)). Next, time stream gaps are filled with a constrained realization of the estimated noise based on neighboring valid data (Hoffman & Ribak (1991)). Data within the gaps are still flagged as invalid for purposes of map making. In the fourth step, the now continuous time stream
is used to re-estimate the noise power spectrum. A deviation from the previously estimated power spectrum usually indicates a poorly selected prewhitening filter. In this case the process is repeated with a new filter. Finally, all filters (instrumental and prewhitening) are deconvolved and time stream gaps are widened accordingly.

6.4.3 Treatment of Data with Signal and Noise

In the case of significant signal, as with the mirror scan synchronous signal, the above noise estimation procedure is performed iteratively. On each iteration, the CMB sky and scan synchronous signal are estimated by map making (§6.5). The expected signals are then subtracted from the time stream yielding an estimated noise-only time stream. The signal estimation is then repeated and residual signal is again subtracted from the time stream, yielding increasingly pure noise.

When this process converges and an accurate noise estimate is obtained, we return to the raw time stream of signal and noise. Instrumental filters are deconvolved from this full time stream and the gaps are correspondingly widened. The gaps are filled using the corresponding samples in the best estimate of the noise-only time stream. The calibration obtained in Chapter 5 is then applied. The full time stream with gaps filled and filters deconvolved is an input of the map making procedure ($d_t$ in Section 6.5).

6.4.4 Time Domain Noise Correlations

After an estimate of the noise power spectrum is found, a time-time noise correlation matrix is calculated. This matrix will be used in map making. Under the assumption of stationary, gaussian noise, noise correlation matrix elements are given by the Fourier transform of the noise power spectrum as

$$N_t(i, j) \equiv \tilde{P}(t = (i - j)\Delta),$$

where $\Delta$ is the sample interval. In practice, this calculation is prone to numerical errors. These can be reduced by smoothing the noise power spectrum, $P(f)$, before deconvolving the prewhitening filter.

We also make the approximation that $N_t(i, j)$ is zero for $|i - j| \geq \lambda_c$, where $\lambda_c$ is the time stream correlation length.
6.5 Map Making

Map making is the next stage of data analysis. The time stream data and
the time-time noise correlations are combined with the pointing solution (§4.4). The
products of map making are a vector of sky temperature contrast for each pixel, \(m_p\),
and the pixel-pixel noise correlation matrix, \(N_p\).

The removal of instrumental filters reduces Equation 6.1 to,

\[ d_i = A m_p + n_i + B x_q, \]

where \(d_i\) is the deconvolved time stream of signal and noise, \(m_p\) is a vector of sky
temperature contrast for each pixel, assigning each time sample to exactly one pixel,
and \(n_i\) is the noise-only time stream with correlations given by \(N_i\), and \(x_q\) is a vector
of the effective temperature contrast of the mirror scan synchronous signal. \(A\) is a
pointing matrix, and \(B\) is a pseudo pointing matrix assigning each time sample to a
mirror orientation.

Neglecting the \(B x_q\) term, a closed form solution for sky signal \(m_p\) and the
associated pixel-pixel noise correlation matrix \(N_p\) is given by:

\[ m_p = (A M A^T)^{-1} A M d_i, \]

\[ N_p = (A M A^T)^{-1} (A M N_i M A^T)(A M A^T)^{-1}. \]

A minimum variance solution is obtained if \(M = N_i^{-1}\). Though this matrix inver-
sion is not impractical for maxima data sections (up to 250,000 samples), various
approximations to \(N_i^{-1}\) (Stompor et al. (2002a)) have been used and found to be
consistent with the exact approach.

6.5.1 Removal of Unwanted Signals

The \(B x_q\) term in Equation 6.3 represents the mirror scan synchronous sig-
nal. However, the treatment of this signal applies equally well to other unwanted
signals, such as overall temperature offsets.

Mirror positions can be treated as extra pixels, observed simultaneously
with the real (sky) pixels. Concatenating \(B\) and \(x_q\) with the corresponding matrixes
for the sky data yields

\[ A' = \left[A, B\right], \]
\[ m'_p = \begin{bmatrix} m_p \\ x_p \end{bmatrix}, \tag{6.7} \]

where \( A' \) and \( m'_p \) are effective pointing and signal matrixes including both sky and scan synchronous signals. Unlike the original \( A \) matrix, \( A' \) assigns each time sample to two pixels: one for the sky coordinate and one for the mirror orientation. The solution for \( m'_p \) and the corresponding \( N'_p \) are still given by Equations 6.4 and 6.5. Once the map of sky pixels and extra pixels is calculated, the signal at the extra pixels is marginalized to obtain the sky map.\(^2\) A further refinement is required because the scan synchronous signal is not stable over the entire flight, or even over every data segment. To account for this, a different set of template pixels is used for every several minutes of data. The stability of the signal over these time scales is tested directly.

This procedure relies on the orthogonality of the scan synchronous signal and the sky stationary signal from the CMB. Due to the double modulation of the telescope in azimuth, the sky signal varies between primary mirror scans; subtraction of mirror synchronous signals does not remove sky stationary signals. This is confirmed by a power spectrum analysis of maps made with and without treatment of the scan synchronous signal. The power spectra are consistent, though the maps without the scan synchronous signal treatment appear excessively noisy.

The same approach can be used for a variety of applications. For example, overall offsets of data segments are assigned to fictitious pixels before combining them. Similarly, we use the \( A \) matrix to assign all corrupted (glitch) data to a single fictitious pixel.

### 6.5.2 Combined Map

The final CMB map and pixel-pixel noise correlation matrix includes data from multiple photometers and multiple independent data sections from each photometer. Assuming that these data are uncorrelated, the equations in this section can be applied directly to the combined data set. The time stream data and pointing matrixes are concatenated, and the noise correlation matrixes are combined into a single block diagonal matrix. The numerical cost is not increased substantially given

\(^2\)The extra pixels approach is not the only way to account for unwanted signals. Other methods involve marginalization earlier in map making.
the known sparsity of the resulting noise correlation matrix. Alternatively, it may be
desirable to analyze blocks of data separately, and add (or subtract) them later. This
process involves the noise-weighted addition of the CMB maps, and the summing of
the inverses of the noise correlation matrixes.

Regardless of the method used, it is difficult to combine data with little
spatial overlap. We rely upon the common structure in the overlap region to constrain
the unknown relative offset between data sections. Insufficient overlap will cause
spurious shifts between sections of the map, or, in extreme cases, will lead to a
singularity and cause the map making (or map combination) to fail.

Combining data from different detectors, each with some calibration error,
can make overall error estimation difficult. Due to the high accuracy of the CMB
dipole calibration, we neglect this effect and assign the combined data the highest
calibration uncertainty of the individual channels.

6.6 Angular Power Spectrum Estimation

The final stage of data analysis is estimation of the angular power spectrum
of the CMB, based on the temperature contrast map and pixel-pixel noise correlation
matrix. Because there is no closed form for the most likely angular power spectrum,
it is necessary. For this process we use MADCAP, a parallel (super-
computer) software package for CMB data analysis. MADCAP's CMB power spectrum
estimation is an implementation of Newton-Raphson iteration, an approach discussed
in Bond et al. (1998).\(^3\)

Power spectrum estimation requires a pixel-pixel correlation matrix for the
CMB map, including both signal and noise. Because the CMB signal and the in-
strumental noise are assumed to be realizations of independent random Gaussian
processes, the total pixel-pixel correlation matrix, \(D_p\), is the sum of the signal corre-
lation matrix, \(S_p\), and the noise correlation matrix \(N_p\). The latter has already been
found. A signal correlation matrix element, for an assumed CMB power spectrum,
is given by

\[
S_{pp'} = \sum_l \frac{2l + 1}{4\pi} B_l^2 C_l P_l (X_{pp'}). 
\]  

\(^3\)The MADCAP package also includes an implementation of the map making algorithm - one of
several used in MAXIMA data analysis.
$C_l$ is the CMB power at multipole $l$, $P_l$ is the $l^{th}$ Legendre polynomial, and $X_{pp'}$ is the angle between the pixels. $B_l$ is spherical harmonic decomposition of the telescope beam shape - the angular window function. We assume an effective circularly symmetric beam profile for any given combination of photometers as described in Wu et al. (2001a). If necessary, the angular window function for the map pixel size may be similarly included.

Due to the limited sky coverage of the experiment, each multipole is not treated as an independent variable. Instead they are grouped into bins (typically 8 to 12 bins of width 75 to 150 multipoles). Bins may or may not be weighted and/or overlapping. In practice we use either top hat shaped bins for simplicity, or use overlapping bins weighted to eliminate residual correlations between bins. If bin width is not too narrow, the difference is very small.

Calculating $S_p$, and therefore $D_p$, for a given multipole binning and assumed power level within each bin, allows us to find the probability distribution for the map, $d_p$, using,

$$P(d_p|C) \propto \exp \left\{ -\frac{1}{2} \left( d_p^T D_p^{-1} d_p + Tr[\ln D_p] \right) \right\},$$  

(6.9)

where $C$ denotes the power level in each bin, as used in the calculation of $D_p$.

Newton-Raphson iteration (Bond et al. (1998)) makes the assumption that the logarithm of this function is quadratic and determines the deviation of the assumed power at each bin from the maximum of the quadratic. Any analytic function is increasingly quadratic near a peak, so calculating this correction iteratively will converge upon a peak. The curvature of the probability distribution around the peak is used as a direct measure of statistical uncertainty in each power spectrum bin. Issues of non-convergence and the possibility of convergence on local maxima are explored in the references.

Power spectrum estimation is the most numerically intensive step in the data analysis process. MAXIMA power spectra have been calculated using supercomputers at NERSC\(^4\) and at the University of Minnesota. A power spectrum from the 3\' pixelization MAXIMA-1 map requires about 20 hours on 256 processor nodes on the NERSC T3E supercomputer.

\(^4\)National Energy Research Scientific Computing Center at the Lawrence Berkeley National Laboratory
6.6.1 Power Spectrum Uncertainties

Uncertainties in the angular power spectrum are derived from instrumental noise, limited sky coverage (i.e. sample variance), and uncertainties in calibration, instrumental filters, pointing reconstruction, and beam shape measurements. Most of these effects vary strongly with $\ell$. Foreground signals are a different kind of uncertainty, discussed in Section 8.1.

The quadratic estimator directly assesses the effects of instrumental noise and sample variance. Noise is most important at high multipoles where there are few observations per mode. It is quantified by the pixel-pixel noise correlation matrix $N_p$. Sample variance is most important at low multipoles where the fewest modes are observed; poorly sampled modes do not strongly affect the probability of a given measured map in Equation 6.9. The combination of these two random effects is the dominant error source at all multipoles. The inherent asymmetry of these errors is modeled using the offset log-normal distribution of Bond et al. (2000).

Calibration uncertainty (Chapter 5) is independent of the multipole bin. As such it does not affect the shape of the power spectrum. Calibration errors are most important for combining data sets. Given the high accuracy of the MAXIMA calibration from the CMB dipole ($\leq$3-4%), we treat the calibration as perfect during data analysis, and assign the highest calibration uncertainty of a set of channels to their combination.

Beam shape errors, including errors in the approximation of circular symmetry and variations between channels in a combined data set, affect the spatial window function $B_\ell$. These effects are discussed in great detail in Wu et al. (2001a). Beam shape errors are significant only at high multipole bins, reaching \(\sim 15\%\) at $\ell = 1000$.

Neither instrumental filters nor pointing reconstruction contribute significantly to angular power spectrum errors. Uncertainties in instrumental filters are small at frequencies contributing to CMB observations. Pointing error (Chapter 4) tends to systematically decrease power at high multipoles, effectively blurring out the small scale features in the CMB. This effect becomes $\sim 10\%$ at $\ell = 1000$, but is less important than beam shape error at all angular scales.
Chapter 7

Results

In this chapter we present the results of the maxima experiment. Data were analyzed using the methods described in Chapter 6. This chapter starts with an overview of the maxima data products (§7.1), presents the CMB map and angular power spectra (§7.2, §7.3), and concludes with cosmological interpretations (§7.4). Chapter 8 presents the results of systematic error tests, including difference maps and power spectra.

7.1 Introduction

Science data have been derived from the subset of maxima-i detectors which show the highest sensitivity and which pass all consistency and systematic tests. CMB data are obtained from three 150-GHz detectors, designated as ‘B25,’ ‘B34,’ and ‘B45.’ Data from a 230-GHz detector, ‘B33,’ were initially included, but later failed consistency tests at $\ell > 785$ and were omitted from higher resolution analyses. Data from a 410-GHz detector, ‘B22,’ are used to monitor dust and atmospheric signals (Chapter 8).

The original analysis at $5'$ resolution was published in Hanany et al. (2000). A companion paper, Balbi et al. (2000), used these data for cosmological parameter estimation. Results of a $3'$ resolution analysis have been published in Lee et al. (2001), with cosmological parameters estimated in Stompor et al. (2001) and Abroe et al. (2002). Results from the maxima-ii data are not available at this time.

Publicly available maxima data can be obtained by request or at:
Figure 7.1: The 3' resolution MAXIMA-1 map derived from three 150-GHz detectors. The indicated central region has the highest signal-to-noise and best cross-linking and was used in the Lee et al. (2001) analysis. The color scale covers a temperature contrast of -750 μK (black) to +750 μK (white).

http://cosmology.berkeley.edu/maxima/data_release/

A number of papers have been published comparing MAXIMA-1 results with those of other CMB experiments (e.g. Wang et al. (2002), Jaffe et al. (2001)). Power spectra derived from MAXIMA-1, BOOMERANG, CBI, DASI, and VSA are generally consistent with each other and with the weaker constraints of previous generation experiments. Jaffe et al. (2001) is a combined analysis of the 3' MAXIMA-1 data and the BOOMERANG data, including all noise correlations.

7.2 The 3-arcminute CMB Map

The current best map from MAXIMA-1 uses the data from three 150-GHz detectors, analyzed with 3' square pixels. For ~10' instrumental resolution, the 3' pixel
window function has little effect and can be ignored. The map consists of \( \sim 40,000 \) pixels; a central rectangular region consists of \( \sim 23,000 \) pixels with uniform sampling, good cross-linking, and a signal-to-noise of \( \sim 5 \) per \( 10' \) beam-size. (Figure 7.1)

Foregrounds are negligible over the entire area (§8.1). Observed signals on angular scales from 10' to 5° are consistent with Gaussianity under a variety of tests including the method of moments, cumulants, the Kolmogorov test, the \( \chi^2 \) test, and Minkowski functionals tests (Wu et al. (2001b)).

7.3 Angular Power Spectra

![Figure 7.2: The power spectrum of the CMB using a hybrid analysis of 5' resolution (up to \( \ell = 335 \)) and 3' resolution (over \( \ell = 335 \)) maps. Error bars show the statistical uncertainties from Table 7.1. The solid curve is the power spectrum of the best fit model from Balbi et al. (2000) with \( \Omega_b = 0.1, \Omega_{cdm} = 0.6, \Omega_\Lambda = 0.3, n = 1.08, \) and \( h = 0.53 \). The crosses are the power spectrum of the difference between one detector and the combination of the other two. (Lee et al. (2001))](image)

The Lee et al. (2001) analysis uses only the central \( \sim 23,000 \) pixel region of the 3' map. This greatly reduces the computation time needed for consistency tests
Table 7.1: The uncorrelated angular power spectrum from maxima-1 (Lee et al. (2001)). The first four \( \ell \) bins are derived from the 5' resolution map, while the last nine are derived from the 3' map. \( \ell_{\text{min}} \) and \( \ell_{\text{max}} \) give the dominant range for each of the overlapping bins. Statistical errors are 68% confidence offset-log normal probability distributions with a constant prior (Bond et al. (2000)), and are purely uncorrelated. Beam errors for the effective combined beam of the detectors (Wu et al. (2001a)) are highly correlated between bins. Pointing uncertainty is an upper limit.

Complete results are available at: http://cosmology.berkeley.edu/maxima/data_release/ requiring a large number of power spectrum calculations (§8.3). A power spectrum computed from the full map agrees with that of the central region, but was not published in Lee et al. (2001) due to the lack of systematic error testing at the time.

At high \( \ell \), essentially all information is contained in the high signal-to-noise region and errors are not substantially increased by restricting the map area. At low \( \ell \), the restricted region significantly increases sample variance errors. To avoid a loss of precision, the power spectrum up to \( \ell = 335 \) was calculated from the full scan region using the well tested 5' resolution map of Hanany et al. (2000), while the higher \( \ell \) data were calculated from the restricted region of the 3' map. The combined power spectrum benefits from large sample area at low \( \ell \), and high resolution at high \( \ell \) with relatively low computational cost.

This combined power spectrum is presented in Table 7.1. A constant 8% error due to calibration uncertainty is not listed in the table.
Since publication of the Lee et al. (2001), systematic testing of the full 3' map and associated power spectrum has been completed. This power spectrum (Figure 7.3) is in strong agreement with that of Lee et al. (2001).

![Angular power spectra from the full 3' CMB map. Dashed (red) and starred (blue) data points represent separate interleaved analyses. Either of these sets alone represents the statistical weight of the experiment. These analyses show no substantial deviations from the Lee et al. (2001) data.](image)

7.4 Cosmological Implications

The most obvious feature of the angular power spectrum is a clear peak at $\ell \approx 220$ followed by relatively low power and the suggestion of additional peaks at higher $\ell$. The presence of a sharp peak is consistent with an inflationary Big Bang and rules out the majority of cosmological defect models. The spatial Gaussianity of the CMB map provides further evidence for inflationary models.

Within standard inflationary models, cosmological parameters have been determined by Baysian (Balbi et al. (2000), Stompor et al. (2001)) and frequentist
methods (Abroe et al. (2002)). All of these are found to yield consistent best fit models and error ranges. The results quoted here are from the Stompor et al. (2001) analysis based on the Lee et al. (2001) power spectrum unless otherwise noted.

Seven independent parameters were varied: $C_{10}$, the amplitude of fluctuations at $\ell = 10$; $\Omega_b h^2$, the physical density of baryons; $\Omega_{cdm} h^2$, the physical density of cold dark matter; $\Omega_\Lambda$, the cosmological constant; $\Omega_{tot}$, the total energy density; $n_s$ the spectral index of primordial fluctuations; and $\tau_c$, the optical depth of reionization. Parameters are sampled over the following ranges:

$$
\begin{align*}
C_{10} & \text{ is continuous} \\
\Omega_b h^2 &= 0.00325, 0.00625, 0.01, 0.015, 0.02, 0.0225, \ldots, 0.04, 0.045, 0.05, 0.075, 0.1 \\
\Omega_{cdm} h^2 &= 0.03, 0.06, 0.12, 0.17, 0.22, 0.27, 0.33, 0.4, 0.55, 0.8 \\
\Omega_\Lambda &= 0.0, 0.1, 0.2, \ldots, 1.0 \\
\Omega_{tot} &= 0.3, 0.5, 0.6, 0.7, 0.75, \ldots, 1.2, 1.3, 1.5 \\
n_s &= 0.6, 0.65, 0.7, 0.75, 0.8, 0.85, 0.875, \ldots 1.2, 1.25, \ldots, 1.5 \\
\tau_c &= 0, 0.025, 0.05, 0.075, 0.1, 0.15, 0.2, 0.3, 0.5
\end{align*}
$$

Likelihoods are interpolated between grid points for additional resolution. The likelihood at each grid point is calculated using an offset log normal approximation (Bond et al. (2000)), including statistical uncertainties and systematic error due to calibration and beam functions. The subdominant systematic effects of pointing uncertainty are neglected. Top hat priors are applied to the Hubble parameter ($0.4 < h < 0.9$), the age of the universe ($t > 10 \ Gyr$), and the matter density ($\Omega_m > 0.1$). Additional low $\ell$ constraints are provided by the COBE results of Gorski et al. (1996).

Constraints on individual parameters and combinations are found by explicit marginalization of all other parameters over the range of our sample grid. We obtain 95% confidence limits on total density $\Omega_{tot} = 0.92^{+0.18}_{-0.16}$, baryon density $\Omega_b h^2 = 0.033 \pm 0.013$, and power spectrum normalization $C_{10} = 690^{+200}_{-125} h K^2$. A constraint on dark matter density $\Omega_{cdm} h^2 = 0.17^{+0.16}_{-0.07}$ is largely based on priors (Jaffe et al. (2001)).

The optical depth to reionization, $\tau_c$, and the index of primordial scalar fluctuations, $n_s$, are degenerate and obey the relationship $n_s = (0.99 \pm 0.14) + 0.46 \tau_c$. 

Figure 7.4: Likelihood function of $\Omega_{\text{tot}}$. The solid line is obtained by maximizing over other parameters while the dashed line adds the constraints that $\Omega_b h^2 = 0.0190 \pm 0.0024$ and $h = 0.65 \pm 0.07$. The horizontal line represent the 95% confidence limits. This plot, based on the 5' analysis is virtually identical to that derived from the 3' analysis. (Balbi et al. (2000))

at 95% confidence for $\tau_c < 0.5$. Setting $n_s$ to 1.0 gives an upper limit of $\tau_c < 0.26$, while setting $\tau_c$ to 0.0 gives $n_s = 0.99 \pm 0.14$ (both at 95% confidence). Regardless of $\tau_c$, we find $n_s > 0.8$ at 99% confidence.

The overall best fit model parameters (Stompor et al. (2001)) are

$$(\Omega_b, \Omega_{\text{cdm}}, \Omega_\Lambda, \tau_c, n_s, h) = (0.07, 0.68, 0.1, 0.0, 1.025, 0.63),$$

with $\chi^2 = 30$ for the 41 MAXIMA-I + COBE power spectrum points and $\chi^2 = 4$ for the 13 MAXIMA-I points. The low estimate of vacuum energy is consistent with recent supernovae results because of the strong degeneracy between matter and vacuum energy. Combining the MAXIMA-I + COBE results with supernovae data (Perlmutter et al. (1999), Reiss et al. (1998)) yields a combined best fit on $(\Omega_m, \Omega_\Lambda)$ of $(0.32^{+0.14}_{-0.11}, 0.65^{+0.15}_{-0.16})$ at 95% confidence (Figure 7.5).
Figure 7.5: Constraints on $\Omega_m$ and $\Omega_\Lambda$ from the combined MAXIMA-1 and COBE DMR data sets as well as those from high redshift supernovae data (Perlmutter et al. (1999), Reiss et al. (1998)). Likelihood contours at 68%, 95%, and 99% confidence are shaded for each. Outlines around $\Omega_m = 0.32$ and $\Omega_\Lambda = 0.65$ are the joint likelihood contours. (Stompor et al. (2001))
Chapter 8

Foregrounds and Systematics

CMB temperature fluctuations are $\sim 10^{-5}$ K. Such a small signal can be obscured by systematic effects including celestial foregrounds, far side-lobe contamination, atmospheric emission, and instrumental instability. Section 8.1 deals with astronomical foregrounds. Section 8.2 summarizes other potential systematic problems. Section 8.3 describes general consistency tests.

8.1 Foregrounds

Foregrounds are the best known source of systematic error for CMB observations and have been widely discussed in the literature (e.g. Bouchet & Gispert (1999)). Jaffe et al. (2002) is a detailed treatment of diffuse Galactic foregrounds in the MAXIMA-I scan region.

MAXIMA deals with foregrounds in three ways. Scan regions are chosen for low foreground emission. Foreground maps and point source catalogues are used to model expected signals. Spectral discrimination provides empirical limits on foreground signals. Subtraction of modeled and spectrally identified foregrounds is viable, but has not been necessary for MAXIMA.

Figure 8.1 is a schematic of the relative importance of foregrounds to different types of CMB experiments. The main foregrounds for MAXIMA are Galactic dust (§8.1.1) and point sources (§8.1.2). Synchrotron and free-free emission and zodiacal dust are secondary concerns (§8.1.3).
Figure 8.1: A schematic of the four main celestial foregrounds for CMB observations. The vertical axis is the optical frequency of the observation, while the horizontal axis is the $\ell$ mode observed. The shaded regions indicate that foreground anisotropy is at least comparable to CMB anisotropy; proper selection of scan regions can greatly reduce the impact of foregrounds. The upper region (red) indicates Galactic dust. The v-shaped region on the right (green) indicates point sources - IR at high frequency and radio at low frequency. The bottom regions represent synchrotron emission (steeper, blue) and free-free emission (shallower, magenta). The black outlines indicate the observed frequencies and angular scales of MAXIMA and several satellite experiments. (Figure by Martin White)

8.1.1 Galactic Dust

Galactic dust is the primary foreground contaminant for the spectral bands and angular range of MAXIMA. Dust emission has a characteristic temperature of 17 K to 21 K and a spectral opacity index of 1.5 to 2.7, with peak emission at 100 $\mu$m to 200 $\mu$m (3 THz to 1.5 THz). Averaging over the entire sky, dust emission is comparable to CMB anisotropy at 150 GHz and much larger at 230 GHz and 410 GHz. The spatial anisotropy of Galactic dust decreases at smaller angular scales as $\ell^{-3}$ (Gautier et al. (1992), Wright (1998)).

Dust models based on COBE/DIRBE, IRAS/ISSA, and COBE/FIRAS data (Schlegel et al. (1998), Finkbeiner et al. (1999)) have been used, both for sky selection and
Figure 8.2: Thermodynamic temperature of dust observed in each of the three maxima-I bands (Jaffe et al. (2002)). Values are derived from cross-correlation of maxima-I and Schlegel et al. (1998). Dust is not strongly detected at 150 GHz or 230 GHz. The curve is the average emission from the Finkbeiner et al. (1999) “Model 8” prediction in the scan region.

for modeling of expected signals. Dust data and models are manipulated via the FORECAST software package (Jaffe et al. (1999)). We have restricted observations to regions of the sky with low dust contrast, >50° from the Galactic plane. The dust in the scan region observed in maxima-I has a predicted in-band equivalent temperature of 10.0 μK (34.5 μK) at 150 GHz (230 GHz) with rms fluctuations of 2.5 μK (8.8 μK) at 150 GHz (230 GHz). For maxima-II, the average equivalent temperature is 9.5 μK (32.8 μK) at 150 GHz (230 GHz) with rms fluctuations of 2.5 μK (8.5 μK) at 150 GHz (230 GHz). These values are normalized to the CMB thermal spectrum. The expected dust signal is ~50 times higher at 410 GHz than 150 GHz.

Correlation with the Schlegel et al. (1998) dust map is used to directly quantify the effects of dust on maxima-I (Jaffe et al. (2002)). The measured dust levels are consistent with zero and are ~1 σ lower than the FORECAST projections
at 150 GHz and 230 GHz (Figure 8.2).

8.1.2 Point Sources

Infrared and radio point sources are most important at small angular scales. A random distribution of unresolved point sources will cause apparent excess power which increases as $\ell^2$.

Catalogued point sources are relatively easy to investigate and can be removed from the map if needed (Sokasian et al. (2001), Gawiser & Smoot (1997)). The MAXIMA-I observing region was selected to be free of known bright sources. The MAXIMA-II observing region is well away from the Galactic plane, but no extra precautions were taken to avoid known point sources. In neither case are known point sources expected to measurably affect the angular power spectra. Power spectra calculated omitting the regions around the brightest known sources are consistent with that of the full map.

Dimmer, uncatalogued point sources are more difficult to handle. For some sources, spectral arguments based on the measured power spectrum at 410 GHz can be used to rule out significant contributions at lower frequencies. However, it is possible to postulate a large number of faint point sources with exotic spectra, the effects of which are highly model dependent. Reasonable estimates (Gawiser et al. (1998)) indicate that uncatalogued point sources are unlikely to be significant for our observations.

8.1.3 Other Foregrounds

Synchrotron and free-free emission are diffuse Galactic foregrounds which peak at radio frequencies and are fairly weak at 150 GHz and higher. MAXIMA scans, which are well away from the Galactic plane, are subject to very little synchrotron and free-free emission. Estimates based on Bouchet & Gispert (1999) yield expected contributions of less than 1 $\mu$K in all of our optical bands. This agrees with upper limits from a correlation analysis of synchrotron and MAXIMA-I maps (Jaffe et al. (2002)).

Zodiacal dust is a diffuse foreground concentrated on the ecliptic plane of the solar system. Zodiacal dust is not a major foreground for CMB anisotropy due
to its low emissivity and smooth spatial distribution. MAXIMA scans are conducted approximately 70° from the ecliptic plane, where the column density and anisotropy of zodiacal dust are negligible.

8.2 Other Systematic Concerns

Though foreground contamination is the most universal systematic error for CMB anisotropy measurement, a variety of other effects must also be considered. Their origins may be optical (e.g. far side-lobe pickup, atmospheric emission) or non-optical (e.g. radio frequency pickup, instrumental noise instability). No list of systematic concerns can be complete; it is possible to postulate any number of instabilities in detector noise and operating temperature, unexpected atmospheric phenomena, or artifacts of readout and data acquisition. Many of the tests described later in this chapter are of a generic nature, sensitive to broad classes of problems.

8.2.1 Far Side-Lobe Contamination

Side-lobe contamination refers to spurious signals from bright sources outside the main lobe of the telescope beams. The temperature contrasts of the balloon, the earth, the Sun, and the Moon are up to ten orders of magnitude greater than that of the CMB.

Pre-flight side-lobe measurements were made using a directional Gunn Oscillator with variable attenuation as a 150-GHz test source (Figure 8.3). Measurement imperfections, such as reflections of the source from the ground, affect these tests; the actual side-lobe sensitivity may be lower than the measurements indicate. Side-lobe response is most overestimated in the lower elevation direction. The measured attenuation at 15° below the beam is sufficient to prevent earth-based side-lobe contamination.

The possibility of side-lobe contamination is minimized by collecting data at night, with the Sun well below the horizon and with the Moon far from the scan region (Chapter 3).

Likely side-lobe contaminants (the Sun, Moon, and features on the Earth) are not oriented constantly with respect to the scan. Differences maps of the two CMB scans for each flight (§8.3.2) are therefore sensitive to side-lobe signals.
Figure 8.3: Data from pre-flight side-lobe tests. The source was roughly 30 m from the telescope. **Left Top**: Test data in the elevation direction for MAXIMA-I. The angle is that of the telescope above the test source. The source is held at fixed elevation while the telescope is aimed at different elevations. **Right Top**: Test data in the azimuth direction for MAXIMA-I. The telescope beam is at fixed elevation (~30°). The source is moved around the telescope at the same elevation. The flat data at roughly -78 dB represent the noise limit of the measurement. The apparent back-lobe is suspected to be an artifact of the measurement technique. If real, such a back-lobe would not affect the CMB data. **Left Bottom**: As above, for MAXIMA-II. Most of these data were collected with the source at higher elevation than the telescope beam (negative angles on the x-axis). The measured response drops more quickly than for MAXIMA-I, probably because of improvements in the test setup. **Right Bottom**: As above, for MAXIMA-II. No apparent back-lobe is observed. This is also likely because of improved measurement techniques.
8.2.2 Scan Synchronous Signals

MAXIMA data show a small spurious signal synchronous with the modulation of the primary mirror, with a CMB temperature amplitude of up to 200 $\mu$K at 150 GHz. It is possible that the scan synchronous signals in MAXIMA-I and MAXIMA-II have different sources.

These signals might be radio frequency pickup (§2.5.3); the radio signal from the on-board telemetry transmitters is partially blocked by the primary mirror mechanism. The modulation of the mirror could vary the degree of pickup by the receiver, leading to a scan synchronous signal. This model is plausible for MAXIMA-I, during which the receiver was relatively susceptible to radio interference. It is less likely for MAXIMA-II because the signals are more prevalent despite additional radio frequency filtering.

A second possibility is variation in atmospheric optical loading as a function of mirror position. This could be explained by a tilt in the mirror motion, causing the telescope beams to change elevation. Under this model, we would expect the synchronous signal to display a strong spectral signature. Atmosphere accounts for less than 1% of the optical load at 150 GHz but about half of the optical load at 410 GHz. The observed scan synchronous signals are similar at all three optical frequencies. In MAXIMA-II the signals vary greatly between detectors, but show no clear correlation with the spectral bands. We conclude that atmospheric loading variations are unlikely to account for the synchronous signals.

Scan synchronous signals may also be caused by electrical or mechanical pickup from the mirror drive. It is difficult to disprove these effects or to estimate their magnitude. However, subsequent tests of the integrated system in the laboratory have failed to reproduce the observed in-flight scan synchronous signals.

Regardless of the source, the map making procedure in Chapter 6 includes a treatment of parameterized parasitic signals (§6.5.1). The assumption that the spurious signal is stable over periods of several minutes has been confirmed in data analysis.

In addition to primary mirror scan synchronous pickup, a signal synchronous to the azimuth modulation of the entire telescope has also been measured. It is relatively small (up to $\sim 50 \mu$K) and much slower than the primary modulation (40 to
70 seconds). This signal is strongly rejected by the faster modulation of the primary mirror; treating it explicitly has no effect on maps and power spectra. It is ignored in the final data analysis.

8.2.3 Telescope Pendulum Motion

Pendulum motion is a well known danger for balloon-borne telescopes. Pendulum motion changes the telescope elevation angle and therefore the observed atmospheric load. The two main pendulum modes of the MAXIMA telescope have frequencies of roughly 0.6 Hz and 0.05 Hz. Of these the mode at 0.6 Hz is relevant to CMB observations. The 0.6-Hz mode is suppressed by a factor of $\sim 10$ using passive pendulation dampers.\footnote{The pendulation dampers, built by Geneva Observatory, are highly damped harmonic oscillators consisting of weighted spheres rolling in oil filled spherical cavities.}

Pendulation modes may be excited by the attitude control motors. The telescope is driven in azimuth using a reaction wheel (§4.3.2) that is symmetric about the rotation axis. However, the telescope is not symmetric and its moment of inertia tensor may have non-zero off-diagonal terms, coupling the azimuth drive to the pendulum modes.

Pendulation amplitude is $<10''$ based on pre-flight tests. No pendulum motion is observed during flight on time scales of $\sim 1$ minute or less within the $1'$ accuracy of the pointing reconstruction. Detector data, including the 410-GHz data most sensitive to atmospheric emission, show no signal at 0.6 Hz or 0.05 Hz.

An irregular variation of up to $20'$ is observed on time scales of $\sim 20$ minutes. This slow pendulation shows a strong correlation with the altitude of the telescope and is believed to result from the coupling of the pendulum modes to vertical motion of the balloon. Regardless of the cause, oscillations on such long time scales do not affect CMB observations.

8.2.4 Secondary Data Errors

Errors in calibration, pointing reconstruction, beam measurement, optical filters, and electronics can affect CMB maps and power spectra. Errors in these data contribute to the error estimates in Chapter 7.
Figure 8.4: 5-arcminute resolution MAXIMA-I maps computed from two single detectors. Neither map is Weiner filtered. **Left:** The map from a 150-GHz detector (B34), the least noisy detector from MAXIMA-I. **Right:** The map from a 230-GHz detector (B33), the noisiest detector used in published results.

### 8.3 Consistency Tests

Systematic errors are examined using general consistency checks that are sensitive to broad classes of effects.

Many systematic tests involve the use of difference maps. Difference maps are generated with the same techniques used for summing maps (§6.5.2). Assuming that the individual maps include only well calibrated sky-stationary signals, difference maps should be consistent with combined detector noise. These maps are subject to a variety of statistical tests, including power spectrum and $\chi^2$ analyses.

The data shown in this chapter are a representative subset of the tests conducted on published MAXIMA-I data. Corresponding systematic tests of MAXIMA-II data are in progress. Further discussions of these and other systematic error checks are found in Stompor et al. (2002b).

#### 8.3.1 Cross-channel consistency

Data from multiple detectors are analyzed separately and in a variety of combinations to test consistency between channels and with the combined map. These tests are sensitive to certain instrumental problems, to spectral consistency,
Figure 8.5: The angular power spectra derived from 5’ resolution MAXIMA-1 maps. Each panel shows the published (Hanany et al. (2000)) 4-channel combined data (filled circles), as well as two single channel power spectra. In each bin, the single channel power spectra have been horizontally offset for readability. **Left:** Triangles (B45) are 150-GHz data and diamonds (B33) are 230-GHz data. **Right:** Triangles (B34) and diamonds (B25) are 150-GHz data.

to the relative spatial offsets of the detectors, and to calibration consistency.

Channels and their combinations are compared in several ways. First, independent maps are compared directly. Figure 8.4 shows the maps from a 150-GHz and a 230-GHz detector. Cross-correlation is used to determine the relative amplitude and position of the measured signals. In all cases, maps are found to be aligned to much better than 1’, with power levels deviating by less than the calibration uncertainty.

Second, angular power spectra from independent maps are compared. Figure 8.5 show the power spectra from four channels, compared to that of the combination of those channels.

Finally, difference maps and angular power spectra are generated. Figure 8.6 shows the difference and sum of a pair of two-channel combinations. Figure 8.7 shows angular power spectra from single channel difference maps.

### 8.3.2 Cross-scan consistency

Each MAXIMA flight includes two largely overlapping CMB scans. The two scans have different scan geometry, telescope orientation, and uncorrelated detector noise.

Cross-scan consistency tests are sensitive to far side-lobe contamination, to
Figure 8.6: 5' resolution MAXIMA-1 maps. **Left**: The combined 4-detector map of Hanany et al. (2000). This map has not been Weiner filtered. **Right**: A difference map from the same four detectors. The detectors are summed pairwise, and these pairs are differentiated to produce this map. The summed pairs are B34 (150 GHz) + B45 (150 GHz) and B25 (150 GHz) + B33 (230 GHz). The difference map is consistent with combined detector noise.

atmospheric pickup, to the time dependence of the responsivity calibration, to time varying instrumental effects, and to pointing reconstruction. Figure 8.8 shows the maps and for the two CMB observations of MAXIMA-1 and Figure 8.9 shows the power spectrum of the their difference.

### 8.3.3 Dark Channel Maps and Power Spectra

The receiver includes three non-optical ‘detectors’. Data from these devices are processed by the same electronics as the optical data, and are sensitive to different subsets of potential instrumental problems. The first is a “dark” bolometer, not coupled to an optical feedhorn. The second is a bolometer-style NTD thermistor, thermally coupled to the ADR. The third is a temperature independent resistor, with impedance similar to that of the bolometers. Four additional sets of dark data are obtained from “bias monitors”, i.e. the locked-in signal from the bolometer AC-bias generators, not connected to the bolometers or cryogenic preamplifiers. Figure 8.10 shows a power spectrum of the dark bolometer in MAXIMA-1.
Figure 8.7: Angular power spectra of difference maps from pairs of single detectors. The four detectors used in Hanany et al. (2000) are differenced in every possible pairing. Each panel shows three of the six combinations. A small amount of residual power is seen at high $\ell$ for some pairings, caused by differences in the beam patterns of the detectors (Wu et al. (2001a)).

8.3.4 Other Consistency Tests

Description of all systematic and consistency tests is beyond the scope of this document; more details are found in Stompor et al. (2002b). In addition to those already discussed, systematic tests have included: selective omission of various subsets of time domain data; variations in acceptance criteria for pointing and detector data; separate analyses of different sections of the CMB maps; variations in map pixelization and resolution; and variations in data analysis.
Figure 8.8: $5'$ resolution 4-channel maps from the two CMB observations of MAXIMA-1. The first scan is shown on the left, and the second scan is shown on the right. Only the overlapping region of the two scans is shown. Each of these maps has a much lower pixel sensitivity than the combined map, due to the lower number of measurements and the lack of cross-linking.

Figure 8.9: An angular power spectrum derived from the difference of the two MAXIMA-1 CMB observations. Based on the Hanany et al. (2000) $5'$ analysis.
Figure 8.10: An angular power spectrum from an optically insensitive detector in MAXIMA-I. Because this “dark” data has no real mapping onto the sky, the pointing solution from an optical bolometer (B34) is used. Similarly, the CMB temperature units on this plot are based on the calibration of an optical bolometer (B34) for purely comparative purposes. The measured power is consistent with detector noise.
Chapter 9

Future Work: Polarization

This chapter discusses CMB polarization anisotropy and the motivation for measuring it (§9.1). Section 9.2 describes MAXIPOL, the polarization sensitive follow-up to MAXIMA.

9.1 Polarization Anisotropy

CMB polarization anisotropy is a fundamental prediction of inflationary Big Bang models. Thomson scattering of CMB photons from free electrons at the surface of last scattering causes a net linear polarization where there is a non-zero local quadrupole (Kosowsky (1999)). The degree of linear polarization is a measure of anisotropy at the time of last scattering; polarization probes the epoch of last scattering directly, unlike temperature fluctuations that can arise in part after last scattering. This “localization in time” makes polarization a strong constraint on the origin of anisotropies, complementary to temperature (Hu & White (1997)).

9.1.1 E-modes and B-modes

Linear polarization of photons is described by two orthogonal components. CMB polarization is naturally discussed in terms of the components $E$ and $B$ (Zaldarriaga (2001)). This form has two advantages over the Stokes parameters $Q$ and $U$. First, the $E$ and $B$ components are independent of coordinate rotation. Second, each has a distinct symmetry under parity inversion. $E$-mode polarization is symmetric under parity, while $B$-mode polarization is anti-symmetric under parity. The
$E$ component is named for its likeness to an electric field (i.e. a pure gradient vector field). The $B$ component is likened to a magnetic field (i.e. a pure curl vector field).$^1$

Because Thomson scattering turns local quadrupoles - which are parity symmetric - into polarization, it can only produce $E$-modes. However, spatial modulation of the perturbations over the surface of last scattering can convert $E$-modes into $B$-modes (Hu & White (1997)). Whether such a conversion actually occurs depends on the source of the quadrupole. Density (scalar) fluctuations do not create $B$-modes, but gravitational waves are tensor fluctuations that do create $B$-modes (Kamionkowski et al. (1997)). $B$-modes are also created by vortical flows, but this mechanism is not expected to contribute significantly. Primordial $B$-mode polarization is expected to be two or more orders of magnitude smaller than $E$-mode polarization.

$E$-modes are also converted into $B$-modes by gravitational lensing (Zaldarriaga & Seljak (1998)). This effect is itself of interest, but for $B$-mode CMB measurement it is significant foreground. It can be subtracted, though only partially, from $B$-mode maps to help search for gravity wave signatures (Hu & Okamoto (2001), Knox & Song (2002)).

Angular power spectra of $E$-mode and $B$-mode polarization anisotropy are denoted as $C_{l}^{EE}$ and $C_{l}^{BB}$ respectively (in this context the temperature power spectrum is generally denoted as $C_{l}^{TT}$). In addition, correlations between $E$-mode and temperature anisotropy lead to a finite cross power spectrum, $C_{l}^{TE}$. Due to parity invariance, $B$-mode CMB polarization anisotropy is uncorrelated with $E$-mode and temperature anisotropies; $C_{l}^{TB}$ and $C_{l}^{EB}$ are zero in the absence of foreground effects.

### 9.1.2 Cosmology with CMB Polarization

CMB polarization provides cosmological information that cannot be obtained from temperature anisotropy alone. The partial polarization of the CMB is a basic prediction of our model of structure formation. If the present structure in the universe grew through gravitational instability from primordial fluctuations, their existence at the time of last scattering would ensure CMB polarization.

$^1$The $E$ and $B$ components are often called $G$ and $C$ for the same reason.
Polarization allows discrimination between types of primordial fluctuations. Parameter estimation from temperature anisotropy relies on the assumption of purely adiabatic (scalar) primordial fluctuations. Temperature anisotropy alone has ruled out pure isocurvature models, but more general models involving a mix of adiabatic and isocurvature fluctuations are still possible (Bucher et al. (2000b)). While different combinations of perturbations may lead to equivalent temperature anisotropy, they can be distinguished in the polarization anisotropy (Bucher et al. (2000a)).

In addition, polarization measurements constrain cosmological parameters and can break degeneracies in temperature data. A great example of this is the reionization of the universe. There is a strong degeneracy in temperature anisotropy between the optical depth of reionization, $\tau_e$, and the overall amplitude of primordial density fluctuations, $A_s^2$. In the $E$-mode power spectrum, reionization causes a distinct peak at $\ell \approx 20$, which breaks the degeneracy (Zaldarriaga (1997)).

\begin{figure}
\centering
\includegraphics[width=0.6\textwidth]{figure91.png}
\caption{Measurement of reionization using temperature anisotropy (dashed) and both temperature and polarization anisotropy (solid), based on the projected sensitivity of the Planck Surveyor. Temperature anisotropy alone cannot break the degeneracy between the optical depth of reionization, $\tau_e$, and the amplitude of density fluctuations, $A_s^2$. (Figure by M. White)}
\end{figure}

$B$-mode polarization, though smaller than $E$-mode polarization, is especially interesting. $B$-modes are a measure of the long wavelength gravitational waves predicted by inflation. The amplitude of these waves is proportional to the energy
scale of inflation. Thus, the detection of $B$-modes would provide overwhelming evidence for inflation and new information about inflationary physics (Kamionkowski & Kosowsky (1998)).

### 9.1.3 Detection of CMB Polarization

The low level of CMB polarization, and the poorly understood foregrounds and systematic error sources, make measuring polarization an experimental challenge. If models currently favored by CMB temperature anisotropy are accurate, the $E$-mode polarization anisotropy is expected to be about an order of magnitude smaller than the temperature anisotropy. Several experiments (Hedman et al. (2001), Keating et al. (2001)) have already set upper limits near the predicted level, and very recently the DASI experiment is believed to have made a detection (Leitch et al. (2002)).

$B$-mode polarization is probably not detectable by current generation experiments. Large bolometric arrays currently under development promise a large increase in sensitivity. However, even with arbitrary sensitivity, primordial gravity wave signatures may be obscured by the effects of weak lensing and other foregrounds.

### 9.2 MAXIPOL

MAXIPOL is the polarization sensitive follow-up to MAXIMA. The primary goal is a robust detection of $E$-mode polarization near the spectral peak of the CMB. This includes detections of power in both the $C^{TE}_l$ and $C^{EE}_l$ power spectra. MAXIPOL is designed primarily for detection and is not expected to measure the shape of the power spectra to high accuracy. Two or three MAXIPOL flights are planned.

MAXIPOL shares the main advantages of MAXIMA - high sensitivity, well optimized scan strategy, excellent calibration, and precise pointing reconstruction - but differs as follows: the bolometric receiver is retrofitted for polarization sensitivity; the scan strategy is modified and the scan region is much smaller ($\sim 10$ deg$^2$ over 2-3 scan regions per flight); and the flight time is longer and data are taken during both day and night ($\sim 30$ hours per flight, with 15-20 hours of CMB observation). Projected power spectrum sensitivity is shown in Figure 9.4.
9.2.1 Polarimetry

Figure 9.2: The 16 element, two color MAXIPOL focal plane array. The 140-GHz detectors observe near the peak of the CMB signal, while the 420-GHz detectors monitor foregrounds and atmospheric emission. All photometers share a single linear polarization, while the polarization of incident light is rotated by a spinning halfwave plate at the telescope aperture stop. All beams have 10' FWHM.

The polarimeter uses a combination of a rotating halfwave plate (HWP) and a stationary polarizing wire grid. This results in modulated sensitivity to both linear polarizations in each detector. This technique is common in infrared and millimeter astronomy, though it is new in CMB science.

The HWP rotates continuously at a frequency of 2 Hz, turning the incident polarization vector at 8 Hz. The linearly polarizing wire grid is between the HWP and the focal plane. Each detector independently produces maps of CMB temperature, and the Stokes parameters $Q$ and $U$ which are converted to $E$ and $B$ for power spectrum estimation. Because each detector is used to measure all three components, cross-calibration does not bias the polarimeter.

As in MAXIMA, 16 single-color photometers are used, but in MAXIPOL, 12 operate at optical bands around 140 GHz with a bandwidth of 45 GHz and four operate around 420 GHz with a bandwidth of 35 GHz. These bands cover the first and third orders of the HWP, respectively. The actual detector spectra are the same as those of MAXIMA at 150 GHz and 410 GHz.\(^2\) Broad spectral coverage is essential for monitoring atmospheric emission and foreground contributions from Galactic

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\(^2\)The change in nominal center reflects the HWP bands. Given the bandwidth and variation between channels, either convention is reasonable.
dust. All photometers have a beam size of 10′ FWHM.

The fastest optical modulation in MAXIPOL is the 8-Hz rotation of the polarization. This modulation strongly rejects noise at lower frequencies, including scan synchronous signals. The polarization rotation frequency is the fourth harmonic of the physical rotation frequency of the HWP; the phase sensitive detection effectively discriminates against spurious signals at lower or higher harmonics. Though an ‘offset’ signal is observed at the polarization rotation frequency, laboratory measurements have found that it is stable to at least the detector noise level.

The 0.45-Hz primary mirror modulation from MAXIMA is not used in MAXIPOL.

### 9.2.2 Scan Strategy

![Pointing simulation of a 13-hour MAXIPOL observation](image)

Figure 9.3: A pointing simulation of a 13-hour MAXIPOL observation. An area of about 10 deg² is covered by the 12 140-GHz photometers. The color code corresponds to the log of the number of detector samples in each 3.5′ square pixel. This scan is realizable in a Spring MAXIPOL flight; it consists of ~10 hours of nighttime data and ~3 hours of daytime data.
The observing strategy for MAXIPOL provides a very deep integration. For the first flight, we will observe two or three scan regions of 4 to 10 square degrees, for a total of about 400 beam-size pixels with an expected noise of 1.4 $\mu$K in each of $Q$ and $U$ and 0.35 $\mu$K in temperature (at 140 GHz).

In order to focus on such a small region, the scan pattern is modified from that of MAXIMA. The gondola is modulated in azimuth with an amplitude of $\sim 2^\circ$ peak-to-peak and a period of $\sim 15$ seconds. Unlike MAXIMA, the center of the modulation is fixed in right ascension and declination and follows the rotation of the sky in both azimuth and elevation. The rotation of the sky sweeps out a ‘bowtie’ pattern about the center of the scans (Figure 9.3).

### 9.2.3 Flights

Balloon flights of up to 36 hours can be achieved during brief launch windows in Spring and Fall from the National Scientific Balloon Facility in Fort Sumner, New Mexico. Using our scan strategy, a region with very low Galactic dust can be observed continuously for up to 13 hours in the Spring and for up to six hours in the Fall.

Two modifications were required to make CMB observations during daylight hours. A great deal of additional side-lobe shielding was required to prevent solar radiation from heating surfaces near the optical path. Similar shielding was used successfully for the BOOMERANG experiment during daylight observations in 1998. In addition, one of the CCD camera star sensors used for pointing reconstruction was modified to observe stars in the daytime. Development of this sensor has been one of my main contributions to MAXIPOL.

Camera tests during daylight hours at altitudes 20 km to 40 km have confirmed that we reliably detect stars of visible magnitude 2.5 and higher, and detect certain stars of visible magnitude 3.0. These limits are more than adequate for the MAXIPOL scan strategy.

### 9.2.4 Beyond MAXIPOL

Current polarization experiments have reached the sensitivities required to measure $E$-modes. The next wave of data from experiments such as MAXIPOL,
MAP and DASI will begin to characterize the $C_l^{EE}$ and $C_l^{TE}$ power spectra. After this, more experiments devoted to CMB polarization will be required. Of particular promise are large bolometric arrays for use in ground-based and balloon-borne polarimeters.

The Planck satellite will provide a great deal of information about $E$-mode polarization and has the benefit of full sky coverage. However, sub-orbital experiments with deeper integration have the potential to discover more about the CMB, especially low amplitude $B$-mode fluctuations.

$B$-mode polarization may not be measurable due to confusion caused by gravitational lensing. As more is learned about CMB polarimetry, the challenge of measuring $B$-modes must be weighed against this possibility. The additional science goal of studying gravitational lensing provides another motivation to observe CMB polarization beyond the level of $E$-modes.
Figure 9.4: Projected performance of MAXIPOL and MAP in measuring the E-mode polarization. Model curves are based on MAXIMA-I best fit parameters. **Top:** The top curve shows the temperature power spectrum, $C_{l}^{TT}$, while the lower curve shows the E-mode power spectrum, $C_{l}^{EE}$. The horizontal lines represent the error on $C_{l}^{EE}$ for MAP (above) and MAXIPOL (below) assuming a single band power estimate for $200 < \ell < 3000$. The very deep integration of MAXIPOL makes it comparable in power to MAP. **Bottom:** The model curve shows the E-mode and temperature cross power spectrum, $C_{l}^{TE}$. Horizontal lines are the single band error projections for MAXIPOL (above) and MAP (below). Sample variance affects the $C_{l}^{TE}$ and $C_{l}^{EE}$ power spectra differently, so the relative power of the experiments varies for the two measurements.
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Appendices
Appendix A

The MAXIMA Collaboration

The following collaborators have been authors on one or more MAXIMA result papers:

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Peter Ade (2)            Phil D. Mauskopf (2)
Amedeo Balbi (3)         C. Barth Netterfield (14)
Domingos Barbosa (4,5)   Sang Oh (13)
James Bock (6,7)         Enzo Pascale (9)
Julian Borrill (8)       Bahman Rabii (4,13,15)
Andrea Boscaldi (9)      Paul L. Richards (13)
Pedro de Bernardis (10)  George F. Smoot (4,13,15)
Pedro G. Ferreira (11)   Radek Stompor (8,16)
Shaun Hanany (1)         Andrew E. Lange (7)
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Appendix B

Calibration Linearity

To calculate the linearity of the detectors during calibrations, we use a combination of theoretical models, measured bolometer properties, and in-flight measurements of bolometer resistance.

The responsivity of a current biased NTD bolometer (Grannan et al. (1997)) is

\[ S = \frac{\alpha IR}{G_d - \alpha I^2 R} \]  \hspace{1cm} (B.1)

where \( S \) is the responsivity, \( I \) is bias current, \( G_d \) is the differential thermal conductance which varies as \( T^\beta \) with \( \beta \) between 1.0 and 3.0 depending on the heat sinking of the bolometer. In the case of \textsc{maxima} \( \beta = 1.0 \) is a good approximation. \( \alpha \) is

\[ \alpha = \frac{1}{R} \frac{\delta R}{\delta T} \]  \hspace{1cm} (B.2)

\( R \) and \( T \) are the NTD resistance and temperature, related by

\[ R(V, T) = R_0 e^{\exp \left[ \left( \frac{A}{T} \right)^n - \frac{eVL}{dk_b T} \right]} \]  \hspace{1cm} (B.3)

where \( V \) is the voltage across the bolometer, \( d \) is NTD thickness, and \( L \) is the characteristic spacing between NTD impurities. \( A, R_0, \) and \( n \) are constants measured before flight. In \textsc{maxima}, \( n \) is 0.5 and \( eVL \ll dk_b T \), so \( R \) reduces to

\[ R(T) = R_0 e^{\sqrt{T}} \]  \hspace{1cm} (B.4)

Equation B.1, given the relations above for \( \alpha \) and \( R(T) \), reduces to,
\[ S = -\frac{1}{2T} \sqrt{\frac{A}{T}} \frac{RI}{G_d + \frac{1}{2T} \sqrt{T^2 R I^2}}. \]  

We differentiate this with respect to NTD temperature, taking into account the dependencies of \( R \) and \( G_d \) on temperature, with bias current \( I \) held fixed\(^1\),

\[
\frac{dS}{dT} = \frac{1}{4T} \sqrt{\frac{A}{T}} \frac{5G_d + \frac{G_d}{T} \sqrt{\frac{A}{T}}}{(G_d + \frac{1}{2T} \sqrt{T^2 R I^2})^2}.
\]  

Finally, we obtain the fractional first order responsivity change, \( \frac{\Delta S}{S} \), corresponding to a measured temperature change,

\[
\frac{\Delta S}{S} = \frac{1}{S \frac{dS}{dT}} \Delta T = -\frac{\Delta T}{2T} \frac{5 + \sqrt{\frac{A}{T}}}{1 + \frac{1}{2T} \sqrt{\frac{A}{T} R I^2}}.
\]  

where \( \Delta T \) is the change in bolometer temperature.

The bolometer responsivity variations in Table 5.2 are calculated using Equation B.7. \( T \) and \( \Delta T \) are calculated using Equation B.4 from \( R \) and \( \Delta R \) as measured in flight. \( A, R_o \) are constants and \( G_d \) at the bolometer operating temperature is measured before flight.

\(^1\)Strictly speaking, the bias current changes with the NTD resistance. However, this is an extremely small effect with \( \frac{\Delta I}{I} \approx 0.01\% \) during calibrations.
Appendix C

Detector Sensitivities

<table>
<thead>
<tr>
<th>Channel [Freq (GHz)]</th>
<th>MAXIMA-I CMB NET $\mu$K $\sqrt{\text{sec}}$</th>
<th>R-J NET $\mu$K $\sqrt{\text{sec}}$</th>
<th>MAXIMA-II CMB NET $\mu$K $\sqrt{\text{sec}}$</th>
<th>R-J NET $\mu$K $\sqrt{\text{sec}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>B14 [150]</td>
<td>142.3</td>
<td>84.7</td>
<td>257.2</td>
<td>145.3</td>
</tr>
<tr>
<td>B15 [150]</td>
<td>87.6</td>
<td>52.9</td>
<td>62.8</td>
<td>36.3</td>
</tr>
<tr>
<td>B24 [150]</td>
<td>177.0</td>
<td>106.0</td>
<td>94.6</td>
<td>54.1</td>
</tr>
<tr>
<td>B25 [150]</td>
<td>78.4</td>
<td>45.2</td>
<td>101.5</td>
<td>57.6</td>
</tr>
<tr>
<td>B34 [150]</td>
<td>88.6</td>
<td>46.1</td>
<td>118.7</td>
<td>66.7</td>
</tr>
<tr>
<td>B35 [150]</td>
<td>185.6</td>
<td>108.0</td>
<td>90.8</td>
<td>51.6</td>
</tr>
<tr>
<td>B44 [150]</td>
<td>271.1</td>
<td>128.0</td>
<td>321.7</td>
<td>179.7</td>
</tr>
<tr>
<td>B45 [150]</td>
<td>92.1</td>
<td>52.6</td>
<td>73.2</td>
<td>41.1</td>
</tr>
<tr>
<td>B13 [230]</td>
<td>316.6</td>
<td>107.2</td>
<td>*</td>
<td>*</td>
</tr>
<tr>
<td>B23 [230]</td>
<td>142.3</td>
<td>46.4</td>
<td>155.7</td>
<td>45.3</td>
</tr>
<tr>
<td>B33 [230]</td>
<td>123.1</td>
<td>37.6</td>
<td>263.0</td>
<td>76.2</td>
</tr>
<tr>
<td>B43 [230]</td>
<td>270.0</td>
<td>89.0</td>
<td>*</td>
<td>*</td>
</tr>
<tr>
<td>B12 [410]</td>
<td>1701.7</td>
<td>66.7</td>
<td>2656.1</td>
<td>104.2</td>
</tr>
<tr>
<td>B22 [410]</td>
<td>2049.6</td>
<td>80.3</td>
<td>2080.9</td>
<td>81.5</td>
</tr>
<tr>
<td>B32 [410]</td>
<td>3013.1</td>
<td>116.9</td>
<td>1699.6</td>
<td>65.9</td>
</tr>
<tr>
<td>B42 [410]</td>
<td>4416.2</td>
<td>173.3</td>
<td>4102.0</td>
<td>160.9</td>
</tr>
</tbody>
</table>

Table C.1: “CMB NET” is the detector sensitivity (§2.4) to CMB temperature fluctuations. Values are derived from calibrated responsivity (Appendix D) and detector noise over a range of 0.1 Hz to 16 Hz. In the case of MAXIMA-II, for which responsivity varies significantly over the flight, average values are used. “R-J NET” is the sensitivity to temperature fluctuations for sources in the Raleigh-Jeans limit (i.e. at least $\sim$15 K at 410 GHz). R-J NET and CMB NET are related by the measured detector spectra.

* Dead channel.
### Appendix D

**Calibration Parameters**

<table>
<thead>
<tr>
<th>Channel</th>
<th>MAXIMA-I Calibration</th>
<th>MAXIMA-I Jupiter Calibration</th>
<th>MAXIMA-II Calibration</th>
<th>MAXIMA-II Mars Calibration</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>(10⁻⁶ V/K)</td>
<td>(10⁻⁶ V/K)</td>
<td>(10⁻⁶ V/K)</td>
<td>(10⁻⁶ V/K)</td>
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<tr>
<td>B14 [150]</td>
<td>10.90±0.26</td>
<td>9.20±1.16</td>
<td>5.13±0.242</td>
<td>5.73±1.17</td>
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<tr>
<td>B15 [150]</td>
<td>11.56±1.16</td>
<td>10.49±1.27</td>
<td>9.21±0.103</td>
<td>12.29±1.25</td>
</tr>
<tr>
<td>B24 [150]</td>
<td>8.90±0.19</td>
<td>8.17±1.01</td>
<td>6.20±0.145</td>
<td>6.96±0.93</td>
</tr>
<tr>
<td>B25 [150]</td>
<td>10.92±0.16</td>
<td>10.34±1.23</td>
<td>7.88±0.198</td>
<td>9.60±0.98</td>
</tr>
<tr>
<td>B34 [150]</td>
<td>9.02±0.16</td>
<td>9.68±1.13</td>
<td>5.68±0.153</td>
<td>7.39±0.75</td>
</tr>
<tr>
<td>B35 [150]</td>
<td>9.04±0.30</td>
<td>8.41±1.04</td>
<td>6.04±0.078</td>
<td>9.21±0.94</td>
</tr>
<tr>
<td>B44 [150]</td>
<td>8.17±0.34</td>
<td>9.65±1.16</td>
<td>5.30±0.174</td>
<td>7.52±0.81</td>
</tr>
<tr>
<td>B45 [150]</td>
<td>10.10±0.23</td>
<td>9.78±1.15</td>
<td>8.20±0.128</td>
<td>9.39±0.59</td>
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<tr>
<td>B13 [230]</td>
<td>3.46±0.28</td>
<td>2.87±0.36</td>
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<td>*</td>
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<tr>
<td>B23 [230]</td>
<td>6.92±0.21</td>
<td>5.95±0.91</td>
<td>4.06±0.129</td>
<td>5.38±0.76</td>
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<tr>
<td>B33 [230]</td>
<td>6.11±0.19</td>
<td>5.65±0.81</td>
<td>4.61±0.262</td>
<td>2.80±0.41</td>
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<tr>
<td>B43 [230]</td>
<td>6.30±0.24</td>
<td>5.51±0.78</td>
<td>*</td>
<td>*</td>
</tr>
<tr>
<td>B12 [410]</td>
<td>** 0.89±0.10</td>
<td>** 0.81±0.09</td>
<td></td>
<td></td>
</tr>
<tr>
<td>B22 [410]</td>
<td>** 0.76±0.09</td>
<td>** 0.62±0.067</td>
<td></td>
<td></td>
</tr>
<tr>
<td>B32 [410]</td>
<td>** 0.66±0.08</td>
<td>** 0.80±0.122</td>
<td></td>
<td></td>
</tr>
<tr>
<td>B42 [410]</td>
<td>** 0.54±0.06</td>
<td>** 0.44±0.046</td>
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</tr>
</tbody>
</table>

Table D.1: Absolute calibration parameters for both MAXIMA flights. “Calibration Dipole” is the responsivity as measured by the dipole observation. “Calibration Jupiter” or “Calibration Mars” is the responsivity as measured by the planet observation.

* Dead channel.

** 410 GHz detectors are not calibrated from the CMB dipole.
<table>
<thead>
<tr>
<th>Channel [Freq (GHz)]</th>
<th>MAXIMA-I Absolute Calibration Ratio</th>
<th>MAXIMA-II Absolute Calibration Ratio</th>
</tr>
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<tbody>
<tr>
<td></td>
<td>Ratio</td>
<td>Stimulator</td>
</tr>
<tr>
<td>B14 [150]</td>
<td>0.95±0.12</td>
<td>1.04±0.02</td>
</tr>
<tr>
<td>B15 [150]</td>
<td>0.96±0.11</td>
<td>1.05±0.02</td>
</tr>
<tr>
<td>B24 [150]</td>
<td>0.95±0.11</td>
<td>1.04±0.02</td>
</tr>
<tr>
<td>B25 [150]</td>
<td>0.99±0.13</td>
<td>1.03±0.01</td>
</tr>
<tr>
<td>B34 [150]</td>
<td>1.08±0.13</td>
<td>1.01±&lt;0.01</td>
</tr>
<tr>
<td>B35 [150]</td>
<td>0.98±0.12</td>
<td>1.05±0.01</td>
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<tr>
<td>B44 [150]</td>
<td>1.18±0.15</td>
<td>1.00±&lt;0.01</td>
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<td>B45 [150]</td>
<td>0.98±0.12</td>
<td>1.02±&lt;0.01</td>
</tr>
<tr>
<td>B13 [230]</td>
<td>0.86±0.12</td>
<td>1.03±0.02</td>
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<tr>
<td>B33 [230]</td>
<td>0.95±0.14</td>
<td>1.03±&lt;0.01</td>
</tr>
<tr>
<td>B43 [230]</td>
<td>0.89±0.13</td>
<td>1.02±&lt;0.01</td>
</tr>
<tr>
<td>B12 [410]</td>
<td>**</td>
<td>1.02±0.02</td>
</tr>
<tr>
<td>B22 [410]</td>
<td>**</td>
<td>1.06±0.01</td>
</tr>
<tr>
<td>B32 [410]</td>
<td>**</td>
<td>1.01±&lt;0.01</td>
</tr>
<tr>
<td>B42 [410]</td>
<td>**</td>
<td>1.02±0.04</td>
</tr>
</tbody>
</table>

Table D.2: Time dependent calibration. “Absolute Calibration Ratio” is the ratio of the responsivity measured during the planet calibration to that measured during the dipole calibration. “Ratio from Stimulator” is the ratio for these two points in the flight, as determined from the internal relative calibrator.
* Dead channel.
** 410 GHz detectors are not calibrated from the CMB dipole.