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HELIOSHEATH MAGNETIC FIELD AND PLASMA OBSERVED BY VOYAGER 2 DURING 2012 IN THE RISING PHASE OF SOLAR CYCLE 24

L. F. Burlaga1, N. F. Ness2, J. D. Richardson4, R. B. Decker3, and S. M. Krimigis4

1 NASA Goddard Space Flight Center, Greenbelt, Maryland, USA
2 Institute for Astrophysics and Computational Sciences, Catholic University of America, Washington, District of Columbia, USA
3 Kavli Center for Astrophysics and Space Research, Massachusetts Institute of Technology, Cambridge, Massachusetts, USA
4 Johns Hopkins University, Applied Physics Laboratory, Laurel, Maryland, USA

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ABSTRACT

We discuss magnetic field and plasma observations of the heliosheath made by Voyager 2 (V2) during 2012, when V2 was observing the effects of increasing solar activity following the solar minimum in 2009. The average magnetic field strength $B$ was 0.14 nT and $B$ reached 0.29 nT on day 249. V2 was in a unipolar region in which the magnetic polarity was directed away from the Sun along the Parker spiral 88% of the time, indicating that V2 was poleward of the heliospheric current sheet throughout most of 2012. The magnetic flux at V2 during 2012 was constant. A merged interaction region (MIR) was observed, and the flow speed increased as the MIR moved past V2. The MIR caused a decrease in the $>70$ MeV nuc$^{-1}$ cosmic-ray intensity. The increments of $B$ can be described by a $q$-Gaussian distribution with $q = 1.2 \pm 0.1$ for daily averages and $q = 1.82 \pm 0.03$ for hour averages. Eight isolated current sheets ("PBLs") and four closely spaced pairs of current sheets were observed. The average change of $B$ across the current sheets was a factor of $\approx 2$, and $B$ increased or decreased with equal probability. Magnetic holes and magnetic humps were also observed. The characteristic size of the PBLs was $\approx 6 R_L$, where $R_L$ is the Larmor radius of protons, and the characteristic sizes of the magnetic holes and humps were $\approx 38 R_L$ and $\approx 11 R_L$, respectively.

Key words: magnetic fields – plasmas – Sun: heliosphere

1. INTRODUCTION

Voyager 2 (V2) has been in the heliosheath since it crossed the termination shock (TS) in 2007 August (Burlaga et al. 2008; Decker et al. 2008; Richardson et al. 2008; Stone et al. 2008) moving $\approx 30^\circ$ below the solar equatorial plane. Voyager 1 (V1) moved through the entire heliosheath $\approx 34^\circ$ above the solar equatorial plane, beginning when it crossed the TS in 2005 December during a time of decreasing solar activity and finally exiting the heliosheath and encountering interstellar plasma and magnetic fields on 2012 August 15. Solar cycle variations of the heliospheric magnetic field strength $B$ and flow speed $V$ have to be considered when trying to understand the variations of $B$ with distance in the heliosheath, as shown by Burlaga et al. (2009) and Bogorelov et al. (2009). The observations of $B$ at 1 au (http://cdaweb.gsfc.nasa.gov/) show that $B$ reached the minimum value of 4 nT in mid-2009. This minimum was associated with a historically low minimum of solar activity (McDonald et al. 2010; Mewaldt et al. 2010; Ahluwalia & Ygubits 2011; Ahluwalia & Jackiewicz 2012; McComas et al. 2013). Solar activity during solar cycle 23 decreased to minimum value in August 2009, when the monthly mean sunspot number was zero. Burlaga et al. (2014) found that the minimum latitudinal extent of the heliospheric current sheet (HCS) from the solar equatorial plane occurred near solar minimum, leaving V2 in a unipolar region sampling magnetic fields from the southern coronal hole during 2011, consistent with the propagation time of the solar wind from 1 to $\approx 90$ au being approximately 1 year.

V2 experienced solar minimum conditions until the beginning of 2011, when the density and temperature increased rapidly at the beginning of solar cycle 24. The density ($N$) and temperature ($T$) reached a plateau during mid-2011 (Richardson & Wang 2010, 2012). The heliosheath plasma observations from 2007.7 to mid-2010 and the magnetic field observations from 2007.7 to 2009.4 were reviewed by Richardson & Burlaga (2013). The galactic cosmic-ray modulation for cycle 24 began to increase in 2010 January. Burlaga et al. (2014) found that $B$ measured by V2 increased in an oscillatory pattern with a period of 86 days during 2011.

This paper discusses the observations of the magnetic field and plasma made by V2 during 2012, when solar activity continued to increase, the distance from the Sun increased from 106.8 to 109.80 au, and the latitude relative to the equator increased from $-30.79^\circ$ to $-31.73^\circ$. Daily observations of the magnetic field $B$ and their relationship to the solar wind plasma throughout the year are described in Section 2. Section 3 describes a merged interaction region (MIR). The directions and the polarity of magnetic field $B$ are discussed in Section 4. The distribution of $B$ and the distributions of daily and hourly increments of $B$ are presented in Section 5. At the smallest scales, V2 observed small-scale structures such as current sheets, magnetic humps, and magnetic holes, which are discussed in Section 6.

2. LARGE-SCALE OBSERVATIONS OF THE MAGNETIC FIELD AND PLASMA

2.1. Magnetic Field Observations

The inboard and outboard magnetometers on V2 measure the three components of the magnetic field $B$ as a function of time. Although the magnetometer no longer functions as a "dual-magnetometer" (Behannon et al. 1977), the independent observations made by the inboard and outboard magnetometers are used in determining the best estimate of the components of $B$ for each 48 s interval (Berdichevsky 2009, http://vgrmag.gsfc.nasa.gov/Berdichevsky-VOY_sensor_opu090518.pdf)
Observations of the corresponding daily averages of the components $BT$, $BR$, and $BN$ in RTN coordinates are shown in Figures 1(b), (c), and (d), respectively. The dominant uncertainty in these measurements, a systematic uncertainty related to the spacecraft magnetic field, is approximately $\pm 0.1$ to $\pm 0.2 \, \text{nT}$. The uncertainty can vary in many ways, ranging from discontinuous changes to linear trends in various intervals. The $BT$ and $BN$ components of $B$ were calibrated using the magrols which occur during a $\approx 5.5$-hr interval every two or three months. The uncertainty in $BT$ and $BN$ is $\pm 0.03 \, \text{nT}$. The $BR$ component cannot be calibrated with the magrols, since the spacecraft rolls about the spacecraft $Z$-axis which is very nearly along the $-BR$ direction. Instead, the $BR$ component is calibrated using the “magcals,” which generally give larger uncertainties than the magrols and occur every $\approx 29$ days. Except for the interval discussed in the next paragraph, the uncertainty in $BR$ is $\pm 0.055 \, \text{nT}$. The uncertainty in the corresponding daily averages of $B$ (the square root of the sum of the squares of the uncertainties in the components of $B$) is $\pm 0.07 \, \text{nT}$.

During the interval between $\approx$-day 90 and $\approx$-day 180 it was more difficult than usual to measure $B$, primarily because of the uncertainty in $BR$ which we estimate to be $\pm 0.10 \, \text{nT}$. Assuming that the uncertainty in $BT$ and $BN$ is $\pm 0.03 \, \text{nT}$, the uncertainty in $B$ is $\pm 0.11 \, \text{nT}$. This uncertainty is primarily associated with systematic errors related to the spacecraft field. As a consequence of this uncertainty, the apparently strong $B$ from days 110 to 180 in Figure 1(a) may be deceptive.

The solid curve in Figure 2 is the same as the curve in Figure 1(a), but we added open circles that represent the values of $B$ that would be observed if $BZ$ in the spacecraft coordinate system were $BZ \sim -BR = 0$, which would be the case if $B$ was in the Parker spiral direction. We have found in previous studies that the average $B$ measured by V1 was close to the Parker spiral direction throughout most of the heliosheath. The strongest $B$ observed by V2 during 2012 was $0.30 \, \text{nT}$ on day 249. Strong magnetic fields were observed during the 28 day interval from day 234 to day 262, which is slightly more than one solar rotation, indicating that the strong fields extended azimuthally around the Sun. Strong magnetic fields were also observed from $\approx$-day 210 through day 229, and possibly earlier. These strong fields observed for more than one solar rotation indicate the presence of a global merged interaction region (GMIR). Since the magnetic field was close to the spiral direction during this interval, $BR \approx 0$. Since the uncertainty in $BN$ and $BT$ was $\pm 0.03 \, \text{nT}$ and the uncertainty in $BR$ was $\pm 0.055 \, \text{nT}$, the uncertainty in $B$ was $\pm 0.07 \, \text{nT}$, which is shown by the lower range of the uncertainty bar plotted in Figure 2(a).

On the other hand, between $\approx$-day 90 and day 180, $BR$ had a large uncertainty of $\pm 0.1 \, \text{nT}$ resulting in an uncertainty of $\pm 0.11 \, \text{nT}$ in $B$, which is shown by the larger limit on the error bar in Figure 2(a).

The presence of relatively strong magnetic fields, on average, at V2 during 2012 is expected since the magnetic fields observed during this time correspond to the increasingly strong magnetic fields that were observed at the Earth’s orbit during 2010 and 2011, when solar activity increased during the beginning of solar cycle 24. Figure 3 shows that solar activity and $B$ tended to decrease from 1990 to 2012, as indicated by the straight lines in Figure 3, while still showing the solar cycle variation. This figure shows a minimum in solar activity and $B$ at 1 au during 2009. Solar activity and $B$ at 1 au increased after mid-2009, so one expects to observe increasingly strong magnetic fields at V2 after a propagation time from the Sun of $\approx 1$ year. Relatively strong magnetic fields were observed at V2...
during 2011 (Burlaga et al. 2014) and during 2012 (as shown in Figure 2(a)).

2.2. Plasma Observations

Daily averages of the proton density \( N \), temperature \( T \), speed \( V \), and the radial component of the velocity \( V_R \) measured by the plasma instrument on \( V2 \) are shown in Figure 4. Figures 4(b) and (c) show that the large-scale fluctuations of \( N \) and \( T \), respectively, appear to be correlated. The density fluctuates about \( 0.002 \text{ cm}^{-3} \) and \( T \) fluctuates about \( 60,000 \text{ K} \) at the beginning and end of the year, but they are larger than this average between \( \approx \text{day 205} \) and \( \approx \text{day 290} \), when the total pressure (magnetic pressure plus thermal plasma pressure, \( P_T = B^2 / 8\pi + NkT \)) was higher than average. The speed profile and the radial component of the speed decreased slowly from \text{day 50} \) to \text{day 158} \) and then increased until \text{day 260} \). The density and temperature were enhanced during the interval in which the speed was increasing between \( \approx \text{day 220} \) and \( \approx \text{day 250} \). The increase in speed was associated with a GMIR which is discussed in Section 3.

2.3. Magnetic Flux Conservation

Parker (1958, 1961) predicted that the magnetic flux should be constant for a stationary radial MHD flow. Richardson et al. (2013) showed that the magnetic flux \( BV_\parallel R \) where \( R \) is the distance of \( V2 \) from the Sun, was approximately constant in the heliosheath from 2007 August to 2011, while the magnetic flux observed by \( V1 \) decreased by a factor of 10 with increasing...
time from 2005 to 2011 in the heliosheath. The decrease in magnetic flux at $V_f$ was related to a decrease in $V_R$ with time; $V_R$ reached zero and even had negative values during 2011 and 2012 (Krimigis et al. 2011), while $B_P$ remained relatively constant. Thus, the decrease in magnetic flux could be largely a consequence of the decreasing speed. Pogorelov et al. (2009, 2012) showed that the decrease in the speed to $V_R \approx 0$ could be explained by a nonstationary three-dimensional model that included solar cycle variations. On the other hand, Provonikova et al. (2013, 2014) and Michael et al. (2015), who considered solar cycle variations in an MHD model with boundary conditions near the Sun, predicted that the speed at $V_f$ should be a constant. They interpreted the decrease in magnetic flux at $V_f$ as a result of magnetic reconnection. Opher et al. (2012), assuming that the magnetic flux remains constant, argued that the decrease of the speed to $V_R = 0$ implies a large increase in $B_P$, which was not observed, and they proposed that the excess magnetic energy was dissipated by the process of magnetic reconnection.

Here, we ask whether the magnetic flux observed at $V_2$ remained constant during 2012 as it did from 2007 to 2010. The magnetic flux $BRV_R$ is plotted as a function of day of year during 2012 in Figure 5. A linear fit to these data gives the slope of $(5 \pm 9) \times 10^{-5}$, which is consistent with 0. In other words, the magnetic flux was conserved at $V_2$ during 2012 as $V_2$ moved from 97.32 to 100.48 au, even though the flow in the heliosheath was being diverted toward the “heliotail” direction (Richardson & Decker 2014).

3. A MERGED INTERACTION REGION

3.1. Structure and Formation of the Interaction Region

By definition, an “interaction region” in the solar wind is a region of enhanced total pressure (magnetic pressure plus plasma pressure; Burlaga & Ogilvie 1970; Burlaga 1995). The “total pressure” $PT = B^2/8\pi + NK^2$ observed during 2012 by $V_2$ is plotted in Figure 4(a). Since the average plasma pressure is one third of the average magnetic pressure during 2012, the total pressure variation is very similar to that of the variation of $B$ during 2012. A region of high pressure was observed by $V_2$ from day 234 to day 262, indicating that the strong fields extended azimuthally around the Sun. High pressures were also observed from $\approx$194 through day 261.

Decker et al. (2015) showed that, from the time after $V_2$ crossed the termination shock (TS) in 2007 through the year 2011, the daily averages of the pressure $P$ of protons with energy between 28 keV and 3.5 MeV (“hot ions”) was significantly greater than the pressure of the thermal plasma. The hot ion pressure decreased from $\approx 2.3 \times 10^{-11}$ dynes cm$^{-2}$ in the year 2007.5 to $\approx 2 \times 10^{-13}$ dynes cm$^{-2}$ at the beginning of 2012 and then declined more rapidly to $6 \times 10^{-14}$ dynes cm$^{-2}$ at the end of 2012. They also showed that the ratio of the partial particle pressure $P$ divided by the magnetic pressure $B_{PLS}/B_{MAG}$ significantly exceeded the ratio of the thermal plasma pressure to the magnetic pressure, $\beta = P_{PLS}/P_{MAG}$. This partial particle pressure does not include the bulk of the pressure contained in pickup protons in the energy range not measured by the PLS and LECP instruments, that is, from $\approx 1$ keV to 28 keV. The partial particle pressure quoted here is thus a lower limit to the total particle pressure (Roelof et al. 2010).

Krimigis et al. (2010) note that shortly after the $V_1$ and $V_2$ crossings of the TS the pressure in the high-energy tail of the PUIs in the heliosheath was significantly higher than the magnetic field pressure. In particular, the pressure at $V_2$ for ions with energies at $E > 28$ keV was $2.3 \times 10^{-13}$ dynes cm$^{-2}$, which was $\approx 5$ times the measured magnetic pressure $\approx 4.3 \times 10^{-14}$ dynes cm$^{-2}$ for $B = 0.1$ nT. Thus, the heliosheath consistently had $\beta > 1$, even counting only the partial pressure due to ions $> 28$ keV. At lower energies, the ion PUI spectra imply a pressure over the energy range $8 – 44$ keV of $7.7 \times 10^{-13}$ dynes cm$^{-2}$, giving a partial pressure of $10^{-12}$ dynes cm$^{-2}$ for the energetic protons ($5 – 3500$ keV) just upstream and also downstream of the TS for $V_1$ and $V_2$. The only component of the pressure not measured directly is $E < 6$ keV, i.e., the energy range of the IBEX images. An estimate for this partial pressure provided by a hybrid simulation is $\approx 1.2 \times 10^{-12}$ dynes cm$^{-2}$. Thus, the overall (isotropic) pressure in the heliosheath immediately downstream of the TS was $2.2 \times 10^{-12}$ dynes cm$^{-2}$.

Figure 6(a) shows the pressures $P$ of protons with energy between 28 keV and 3.5 MeV observed by $V_2$ during 2012, and Figure 6(b) shows the corresponding ratio of $\beta$. As during the earlier years, the pressure of the hot plasma greatly exceeds the pressure of the warm plasma. However, unlike any time interval during the previous years in the heliosheath, $V_2$ observed that the magnetic pressure was significantly greater than the pressure of protons ($28$ keV – $3.5$ MeV) when the GMR moved past $V_2$ during 2012. Nevertheless, in view of the estimates given by Krimigis et al. (2010), the magnetic pressure of $2 \times 10^{-11}$ dynes cm$^{-2}$ observed by $V_2$ during 2012 is still smaller than the total isotropic pressure of $10^{-12}$ dynes cm$^{-2}$ of the energetic protons ($5 – 3500$ keV) just downstream of the TS. The pressure of protons ($28$ keV – $3.5$ MeV) decreased by a factor of $\sim 3$ from the TS to the position of $V_2$ during 2012 (Decker et al. 2015). Assuming that the total pressure decreased by the same factor of 3 between the position of the TS and the position of $V_2$ during 2012, the overall pressure at $V_2$ during 2012 was $\approx 7 \times 10^{-13}$ dynes cm$^{-2}$, which is larger than $\approx 2 \times 10^{-12}$ dynes cm$^{-2}$ for the magnetic field pressure with $B = 0.2$ nT in the GMIR during 2012.
We can be confident that the magnetic pressure was indeed high during the 29-day interval from day 234 to day 262, because $B$ was close to the Parker spiral direction, implying that the uncertainties in the $BR$ component were not contributing to $B$. The uncertainty in $B$ was $\pm 0.04$ nT. Since the high-pressure region moved past V2 for more than a solar rotation, it was a GMIR formed by the overlapping of interaction regions beyond 1 au (Burlaga et al. 1983; Burlaga 1995). In view of the results of the previous paragraph, we expect a GMIR in the outer heliosphere and heliosheath to be a region of high “total pressure” where this pressure includes that of the magnetic field, thermal plasma and energetic nonthermal ions. The enhancements in the density and temperature observed during the passage of this region are also consistent with a GMIR, since they contribute to the thermal pressure of the plasma. High pressures were also observed between day 213 and day 234, suggesting that the GMIR might have moved past V2 over an interval of 49 days, but the uncertainties in $B$ were larger during this interval. It is even possible that this GMIR might have arrived at V2 as early as day 180, but there were significant uncertainties in $B$, making the identification of a GMIR extending to these early times ambiguous.

In principle, one can trace the radial evolution of this MIR by determining the appropriate boundary conditions near the Sun and using a realistic physical model to propagate these conditions to the position of V2. In practice, one cannot determine the boundary conditions, although various methods have been used in order to estimate them. Determining the boundary conditions associated with the GMIR is beyond the scope of this work. However, it is clear that one must take into account the fact that solar activity was increasing during 2011.

This point is illustrated in Figure 7, which shows the velocity $V$, and density $N$, temperature $T$ and $B$ observed at 1 au. The observations are taken from the OMNI data set. During 2010, shortly after solar minimum, the structure of the solar wind at 1 au was dominated by a series of quasi-periodic corotating streams with high speeds and temperatures preceded by regions of relatively high density, as shown in Figures 7(a)–(c). During 2011, the solar wind at 1 au was qualitatively different and more complex than it was during 2010. The corotating streams no longer stand out in Figure 7(a) during 2011. Instead, one probably sees a mixture of corotating streams and transient flows that might include coronal mass ejections, magnetic clouds and shocks. This statement needs to be confirmed by a detailed analysis of the plasma and magnetic field observations.
at 1 au. Several spikes of $N$, $T$, and $B$ observed during 2011 are larger than any changes observed during 2010. In summary, these observations show that the structure of the solar wind at 1 au was not uniform, and that there were streams of varying widths at 1 au.

The existence of such closely spaced, relatively narrow streams implies that there would be nonlinear interactions among the flows as they move radially away from the Sun. The interactions would cause the range of flow speeds to decrease with increasing distance from the Sun (Collard et al. 1982). Fast streams would eventually be destroyed as they moved away from the Sun, and nonlinear pressure waves with increased magnetic fields and density would have been created (e.g., Burlaga et al. 1985b; Burlaga 1995). Nonlinear pressure waves would be created by the merging of individual interaction regions, producing MIRs. Shock waves could also be created in this process. Memory of the initial conditions tends to be lost in such an interaction (Burlaga et al. 1983). It is reasonable to expect that similar interactions occurred at the latitude of V2, but models with the appropriate boundary conditions are needed to demonstrate the validity of this expectation for the particular interval under consideration.

There is extensive theoretical literature which predicts that the pressure of pickup ions (primarily protons) is significantly greater than the magnetic pressure and the thermal pressure of the solar wind particles (Zank 2015). There is evidence that pickup ions make the dominant contribution to the pressure in the distant heliosphere (Burlaga et al. 1994) and in the heliosheath (Richardson 2008; Richardson et al. 2008; Krimigis et al. 2010; Decker et al. 2015). Thus, models of the dynamical evolution of the solar wind must include both thermal and nonthermal ions.

Burlaga et al. (2005) discussed a case in which a single fast stream overtook several slower streams at 1 au and produced a broad GMIR in the distant heliosphere where it was observed by V2. The observation that the plasma speed at V2 was increasing from day 194 to day 263, 2012 (Figure 4) is consistent with the idea that the GMIR observed by V2 during 2012 was produced in this way. In order to verify this hypothesis, one needs a model that can follow the formation and evolution of the GMIR in the distant supersonic solar wind, its passage through the TS, and its evolution within the heliosheath.

A one-dimensional MHD model with pickup ions was used by Burlaga et al. (2007b) to follow the radial evolution of the
supersonic solar wind flows out to 90 au. The passage of the supersonic solar wind to the subsonic heliosheath by means of a TS was modeled by Borovikov et al. (2011). A variety of models (e.g., Pogorelov et al. 2009, 2012, 2013; Opher et al. 2011, 2012; Opher & Drake 2013; Washimi et al. 2011; Borovikov et al. 2012; Fisk & Gloeckler 2013) describe the structure and dynamics in the heliosheath. Recent models by Provornikova et al. (2013, 2014) and by Michael et al. (2015) predict the observed B and radial speeds at V1 and V2. Whang & Burlaga (1994; Whang et al. 1995) followed the evolution of a GMIR through the supersonic wind and heliosheath and its interaction with the heliopause. Liu et al. (2014) calculate the propagation of a transient from 1 to 120 au; the MHD model does not include the crossing of the TS and heliosheath. A recent model follows the evolution of an MIR as it moves past the TS, through the heliosheath, and generates a pressure wave in the interstellar medium. Models such as these, with the appropriate boundary conditions, should be applied to observations discussed in this paper, but that is beyond the scope of this paper.

3.2. Effects of the MIR on Cosmic Rays

Burlaga et al. (1985a) observed that strong magnetic fields were associated with a large decrease in >70 MeV per nucleon cosmic-ray intensity (CRI) in the V1 data obtained at 11 au. More specifically, they observed that when B is greater than the average B during a given year (⟨B⟩), the CRI decreases with time at a rate proportional to (B/⟨B⟩ − 1) dCRI/dt = D × (B/⟨B⟩ − 1) and when B > ⟨B⟩ the CRI increases at a constant rate (dCRI/dt = R). This useful empirical relationship, called the “CR–B relation,” describes the qualitative behavior the cosmic rays during one year throughout supersonic wind and within the heliosheath (Burlaga et al. 2011), except near the TS and heliopause.

If the strongest magnetic fields observed by V2 during 2012 were related to a MIR or GMIR, one would expect them to produce a decrease in the CRI. Figure 8(a) shows B during 2012 and Figure 8(b) shows the CRI during 2012. In the first half of 2012 the CRI was relatively constant. From approximately day ≈150 to day ≈210, the CRI increased. As discussed above, there were large uncertainties in the V2 measurements of B. The nominal values, shown by the symbols “x” in Figure 8(a), are significantly greater than the average during this interval, which would imply a decrease in the CRI, in contrast to the observed increase. If one sets the BR component of B equal to 0, as allowed within the error bars, giving magnetic fields close to the Parker spiral direction, one obtains the B(t) shown by the curve bounding the shaded area in Figure 8(a). Between day 150 and day 200, B(t) along this curve is less than the average value for 2012, which would be consistent with the observed increase in the CRI. The strong magnetic fields observed after day 200 produced a CRI decrease from ≈day 213 to day 262 during the passage of the MIR, as expected from the CR–B relation.

The solid curve in Figure 8(b) shows the CR–B relation computed from the shaded curve B(t) in Figure 8(a) using the parameters indicated in the figure. This curve shows the qualitative features of the CRI profile, namely, the rapid rise early in the interval, a much slower increase of CRI in the middle of the interval, and the large decrease in CRI toward the end of the interval. However, there are differences between the profile given by the CR–B relationship and the observations, indicating the limitations of the empirical CR–B relationship. A physical model is needed to explain the details of the CRI observations.

The important point is that strong magnetic fields indicative of a GMIR were observed at V2 during 2012, a period when the Sun became increasingly active, and the GMIR produced a substantial decrease in the CRI at V2, as expected from previous studies.

4. THE DIRECTION AND POLARITY OF THE MAGNETIC FIELD

The azimuthal angle (λ) and elevation angle (δ) of the magnetic field direction in the spacecraft centered RTN coordinate system during 2012 are plotted in Figures 2(a) and (b), respectively. The strongest magnetic fields were observed near days 234–262, with a maximum of 0.30 nT on day 249. The strong magnetic fields were oriented close to the Parker spiral direction λ = 270° and δ = 0° and they were observed at the time of one of the magrols (shown by the symbols “x” at the top of panel (a)), when B could be well-determined.

Figure 2(a) shows that λ was significantly closer to the radial direction than to the Parker spiral direction from day ≈135 to ≈day 232 when B appears to be large. There are several days during this interval in which the angles were nominally >215°. If there were anisotropic energetic particles propagating along this direction they should have been observed by LECP at these times, but such particles were not seen. Therefore, such large deviations from the Parker spiral direction (and consequently a large BR component of the magnetic field) were probably not actually present, consistent with the large of uncertainties in the measurements.

The distributions of the daily averages of λ and δ and B are shown in Figure 9. The distribution of λ shown in Figure 9(b) indicates that 88% of the points are between 180° and 360°, corresponding to a primarily “away” polarity, indicating that V2 was poleward above the HCS most of the time, and measuring mostly magnetic fields from the Sun’s southern hemisphere, as expected for solar cycle 24. A Gaussian distribution of λ = y0 + (A/(w × sqrt(π/2)) × exp(−2 × ((B − Bc)/w)²), provides a good fit (R² = 0.87) to the observations of λ between 180° and 360°. The fit indicates that the width (2σ) of the distribution is 56° ± 8°, and the most probable value of λ is 277° ± 4°, which is consistent with the angle 270° that one expects from a Parker spiral magnetic field. The deviations of many points from the Parker spiral direction discussed above are reflected in the quality of the fit.

The distribution of elevation angles δ shown in Figure 9(a) is Gaussian, as shown by the solid curve in Figure 9(a), which was derived from the fit for which R² = 0.99. The width (2σ) of the distribution of δ is 56° ± 4°. The most probable value of δ is 277° ± 1°, consistent with 0° which is expected for an average Parker spiral magnetic field. Note that the elevation angle is more consistent with the Parker spiral field angle than the azimuthal angle.

5. DISTRIBUTIONS OF B AND INCREMENTS OF B

Previous studies (e.g., Burlaga et al. 2006) have shown that the distribution of B in the heliosheath is often described by a Gaussian distribution when B is unipolar for more than a solar rotation. When B is very disturbed, the distribution of B


described by a Gaussian distribution and the distribution of equally well by a Gaussian or a lognormal distribution. The magnetic field was nearly unipolar. The distribution of magnetic field strength and the cosmic-ray counting rate as a function of time. Panel (a) shows the daily averages of the nominal $B$ plotted in Figure 1 (crosses) and $B$ assuming that $B$ was in the in the Parker spiral direction (shaded region). Panel (b) shows the greater than 70 MeV/nuc cosmic-ray counting rate (line with solid squares) and the rate predicted by the CR–$B$ relation calculated assuming that the magnetic field was in the Parker spiral direction. The empirical CR–$B$ relation describes the general trends cosmic-ray profile. Generally shows a large tail and the distribution can often be approximated by a lognormal distribution.

The distribution of daily averages of $B$ for the V2 2012 data is shown in Figure 9(c). A Gaussian distribution provides a good fit to the observations ($R^2 = 0.93$), with the most probable value 0.132 nT and the width 0.126 nT corresponding to a standard deviation of $SD = 0.063$ nT. A Gaussian distribution is often observed in the heliosheath and heliosphere when the spacecraft observes a unipolar magnetic field. The polarity measured by V2 during 2012 was directed predominantly (88%) “away” from the Sun. However, a lognormal distribution $y = y_0 + A/(\sqrt{2\pi}) \times w \times B \times \exp(-(\ln(B/B_0))^2/(2w^2))$ provides an equally good fit to the V2 observations ($R^2 = 0.93$), with the most probable value being $B_0 = 0.141$ nT and a width of $w = 0.461$ nT. A lognormal distribution is often observed when $B$ is disturbed, as it was at V2 during 2012 as a consequence of solar activity.

In the distant solar wind and heliosheath, the distribution of increments of $B$ ($dB = B(t + \tau) - B(t)$) is generally a $q$-Gaussian distribution function (Burlaga et al. 2006, 2007b) which has the form $[1 + (q - 1)/q - \beta^2]^\alpha$ where $q = (1/\alpha - 1)$ (Tsallis 1988, 2009). This is also true of the V2, 2012 observations of the daily and hourly increments of $B$, as shown in Figure 10.

Figures 10(a) and (c) show the hourly increments of $B$ ($dB_{1h} = B(t + 1 \text{ hr}) - B(t)$ and daily increments of $B$ ($dB_{1d} = B(t + 1 \text{ day}) - B(t)$), respectively, as a function of time. Both signals have the characteristics of a noise signal. The distributions of the hourly and daily increments signals are shown in Figures 10(b) and (d), respectively, where the abscissa has the same range in both figures. The distribution of increments of $B$ for our intervals is narrower and has a more extensive tail than the distribution of $B$ for daily intervals. Both distributions can be accurately described by a $q$-Gaussian distribution, with $R^2 = 0.996$ and $R^2 = 0.990$ for the hourly
and daily increments, respectively. The parameter \( q = 1.82 \pm 0.03 \) describing the hourly increments of \( B \) is typical of heliosheath values, corresponding to a distribution with large non-Gaussian tails. The parameter for daily increments of \( B \), \( q = 1.2 \pm 0.1 \), indicates that the distribution is close to a Gaussian distribution, for which \( q = 1 \). A \( q \)-Gaussian distribution in \( B \) is a signature of intermittent compressive turbulence in the heliosheath.

6. CURRENT SHEETS

6.1. Introduction

The monograph by Parker (1994, p. 3) demonstrates that “…almost all continuous magnetic field configurations develop internal discontinuities as they relax to equilibrium.” By “continuous magnetic field configurations” Parker is referring to large-scale MHD configurations anywhere in the universe, such as the magnetosphere, Sun, stars, and the galaxy. The book begins with the statement “… electric currents associated with magnetic fields are universally partially concentrated into widely separated current sheets.” Parker’s emphasis is on current sheets that are generated by the twisting and interweaving of magnetic fields, but the basic theorem of magnetostatics also leads to the formation of tangential discontinuities across which there is no change in the direction of \( B \) locally.

Four types of current sheets have been observed in the solar wind and the heliosheath: proton boundary layers (“PBLs”) associated with the tangential discontinuities discussed by Parker (1994), magnetic holes, magnetic humps, and sector boundaries. This section discusses PBLs, magnetic holes, and magnetic humps observed in the heliosheath by V2 during 2012, which were relatively abundant compared to previous years. PBLs and sector boundaries are probably equilibrium structures, but magnetic holes and magnetic humps may be related to propagating solitons.

We searched for relationships between the magnetic field and plasma parameters associated with these small-scale

![Figure 10](image-url). Panels on the left show (a) hourly increments of \( B \) as a function of time and (b) the distribution of the hourly increments of \( B \). The panels on the right show daily increments of \( B \) as a function of time and (b) the distribution of the daily increments of \( B \). The distributions are described by the \( q \)-Gaussian distribution, where \( q = 1.82 \pm 0.03 \) for the hourly averages and \( q = 1.2 \pm 0.1 \) for the daily averages.
structures, but no clear relationship was found. This is probably because pickup ions make the dominant contribution to the pressure in the distant heliosphere (Burlaga et al. 1994) and in the heliosheath (Richardson 2008; Richardson et al. 2008; Krimigis et al. 2010; Decker et al. 2015; Zank 2015). Pickup protons must be considered in models of PBLs, magnetic holes, and magnetic humps in these regions.

6.2. Proton Boundary Layers

The current sheets associated with tangential discontinuities were first observed in the solar wind at 1 au by Siscoe et al. (1968) and Burlaga (1969), who found that they have thicknesses \( \lesssim 10 \) proton Larmor radii, \( R_L \). Current sheets associated with tangential discontinuities observed at 1 au were modeled by Lemaire & Burlaga (1976), who modeled them as boundary layers across which the sum of the gas pressure, magnetic pressure, and electric field pressure is constant and whose internal structure is described by the Vlasov equation and Maxwell’s equations. In stable current sheets, the current is carried by the protons, hence the current sheets are called “PBLs.” The thickness of PBLs was predicted to be a few proton gyroradii at 1 au.

Lemaire & Burlaga (1976) suggested that the width of PBLs, measured in units of \( R_L \), would be the same inside and outside 1 au as the width observed at 1 au. This prediction has been verified with data from all regions of the heliosphere that have been explored by suitably instrumented spacecraft. References to these early studies are given by Burlaga & Ness (2011), who analyzed observations of PBLs in the heliosheath by \( V_f \) during 2009 between 108.5 and 111.8 au.

As referred above, pickup protons make the dominant contribution to the pressure in the solar wind beyond approximately 20 au and in the heliosheath. In this case, the thermal pressure of the pickup protons determines the relevant gyroradius, rather than the thermal pressure of the solar wind protons. The gyroradius (or Larmor radius) \( R_L = V_{th}/\omega_L \), where the gyro-frequency \( \omega_L = eB/mc \) and \( V_{th} = \sqrt{2kT/m} \), where we consider only the two components of the velocity perpendicular to the gyroradius. Unfortunately, there are no in situ observations of the pickup protons in the energy range of interest and the pickup proton distribution is non-Maxwellian. Nevertheless, one can estimate, based on observations from the LECP instrument, that the effective temperature of the pickup protons is between \( 4 \times 10^6 \) and \( 10^7 \) K [Krimigis et al. 2010]. We choose the average value \( 7 \times 10^6 \) K, which could be accurate to \( \pm 5000 \) K. In this case, the characteristic thermal speed is \( V_{th} = 315 \pm 75 \) km s\(^{-1}\).

Observations show that a PBL generally consists of (1) a region of nearly uniform \( B (B_1) \), (2) a current sheet seen by a spacecraft as a curve \( B(t) \) with an inflection point in the center of the sheet followed by (3) a region of different uniform magnetic field strength (\( B_2 \)). The two uniform regions on either side of the current sheet allow us to determine the size of the current sheets.

An example of a PBL observed by \( V_2 \) on day 84, 2012 is presented in Figure 11 which shows \( B(t) \) decreasing along a smooth curve from a relatively large value \( B_1 \), through an inflection point, to a constant value \( B_2 \). We call the structure a “PBL” to indicate that \( B \) was decreasing with increasing time as the structure moved past \( V_2 \). The curve shown in the data is a fit given by the sigmoid (Boltzman) function,

\[
B(t) = B_2 + (B_1 - B_2)/(1 + \exp[(t - t_0)/\tau])
\]

where \( t \) is in units of days, \( \tau \) is related to the size of the structure, and \( t_0 \) is the time that the inflection point of \( B(t) \) moves past \( V_2 \). The quantity \( w = 4.4 \tau \) gives the width of the sigmoid fit that includes 90\% of the change in \( B(t) \) (Burlaga & Ness 2011). The sigmoid function is proportional to the arc tangent function, which is more familiar but more difficult to use to fit the observations. In this case, the parameters derived from the fit are \( B_1 = 0.128 \) nT, \( B_2 = 0.036 \) nT, \( \tau = 0.00943 \) days (hence \( w = 1.00 \) hr), and the time of the inflection point, which we use to label the PBL, is \( t_0 = \) day 84.184. The quality of the fit, measured by the coefficient of determination is \( R^2 = 0.96 \). This set of numbers, and a similar set of numbers for two other PBL observed by \( V_2 \) during 2012 are given in Table 1.

Figure 11 also shows the direction of \( B \) associated with the PBL observed on day 84.184 as a function of time. At early times, before \( \approx \) day 84.19 when \( B \) is relatively strong, \( B \) is along the Parker spiral direction (\( \delta = 0^\circ, \lambda = 270^\circ \)). The fluctuations of \( B(t) \) and the angles are relatively small. Later, when \( B \) decreased below 0.05 nT, the fluctuations in \( B \) and the angles are significantly larger. These fluctuations are a manifestation of the uncertainties in the measurements, which are increasingly large as \( B \) decreases below \( \approx 0.05 \) nT.

Note that there is no significant difference between the direction of \( B \) before and after the current sheet associated with the PBL on day 84, when one considers that the uncertainties in the angles can be \( \pm 30^\circ \) if \( B \) is less than 0.04 nT. This invariance of the direction of \( B \) across a PBL is a general, but not universal, feature of PBLs.
Five PBLs in which $B$ increased with time were also observed by V2. We refer to these PBLs as PBLi. The inflection points were on day 5,814, day 34,69, 90,102, and day 250,821 during 2012. All of these isolated PBLs can also be described by the sigmoid function. The corresponding parameters of the fits to the observations are shown in Table 1. The bold characters in Tables 1–4 highlight the times of the current sheets and the most important derived parameters describing these current sheets.

A pair of PBLs was observed on days 61,151 and 61,246; a second pair of PBLs was observed on day 270,071 and 270,155, shown in Figure 12; and the third pair was observed on day 71 shown in Figure 13. All of these pairs of PBLs can be described by portions of the sigmoid function. The corresponding parameters of the fits are listed in Table 2. The average characteristic sizes of these of PBLd and PBLi, are $(114,000 \pm 8,000) \text{ km} = (6.5 \pm 0.5) R_L$ and $(93,000 \pm 11,000) \text{ km} = (7 \pm 6) R_L$, respectively. These sizes are somewhat smaller than the sizes $(191,000 \pm 177,000) \text{ km} = (5 \pm 4) R_L$ and $(175,000 \pm 57,000) \text{ km} = (5 \pm 2) R_L$ that we found for the isolated PBLd and PBLi (Table 1), but we cannot say that the differences in sizes are statistically significant. Note that the error bars represent the standard deviation about the average size, not the uncertainties of the measurements themselves. Moreover, the Larmor radius $R_L$ is a characteristic value, rather than the value associated with the parameters, since the thermal speed is not measured for each current sheet.

The average size for both isolated PBLs and pairs of PBLs is $\approx 6 R_L$. Given the uncertainties and limitations in the observations this value compares favorably with the results of the average thickness of $12 R_L$ observed at 1 au by Burlaga et al. (1977), the thicknesses of $2 R_L$ to $10 R_L$ predicted by Lemaire & Burlaga (1976), and the thickness $\approx 15 R_L$ observed by Burlaga & Ness (2011).

### Table 1

<table>
<thead>
<tr>
<th>Event</th>
<th>PBLi</th>
<th>PBLi</th>
<th>PBLi</th>
<th>PBLi</th>
<th>PBLi</th>
<th>Av</th>
<th>SD</th>
<th>PBLd</th>
<th>PBLd</th>
<th>PBLd</th>
<th>Av</th>
<th>SD</th>
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<td>DOYc</td>
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<td>34,69</td>
<td>90,102</td>
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<td>28,534</td>
<td>84,184</td>
<td>111,226</td>
<td>5,814</td>
<td>34,69</td>
<td>90,102</td>
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<td>28,534</td>
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<td>$R^2$</td>
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<td>0.95</td>
<td>0.96</td>
<td>0.95</td>
<td>0.99</td>
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<td>0.042</td>
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<td>0.189</td>
<td>0.07</td>
<td>0.07</td>
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<td>0.116</td>
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<td>0.14</td>
<td>0.10</td>
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<td>$(B_2 - B_1)/\text{nT}$</td>
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<td>0.064</td>
<td>0.06</td>
<td>0.074</td>
<td>0.12</td>
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<td>$(B_2 + B_1)/2 \text{nT}$</td>
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<td>0.056</td>
<td>0.079</td>
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<td>$B_2/B_1$</td>
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<td>0.00532</td>
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<td>0.0017</td>
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<td>$w(\text{hr}) = 4.4 \tau$</td>
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<td>115</td>
<td>111</td>
<td>112</td>
<td>107</td>
<td>107</td>
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<td>104</td>
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<td>3</td>
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<tr>
<td>$L(1000 \text{ km})$</td>
<td>192</td>
<td>127</td>
<td>225</td>
<td>227</td>
<td>103</td>
<td>175</td>
<td>57</td>
<td>68</td>
<td>394</td>
<td>111</td>
<td>191</td>
<td>177</td>
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<tr>
<td>$Vth(\text{km/s})$</td>
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<td>315</td>
<td>315</td>
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<td>315</td>
<td>315</td>
<td>315</td>
<td>315</td>
<td>315</td>
<td>315</td>
<td>75</td>
</tr>
<tr>
<td>$R_L(1000 \text{ km})$</td>
<td>55</td>
<td>44</td>
<td>59</td>
<td>42</td>
<td>13</td>
<td>43</td>
<td>18</td>
<td>23</td>
<td>40</td>
<td>35</td>
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<td>9</td>
</tr>
<tr>
<td>$L/R_L$</td>
<td>3</td>
<td>3</td>
<td>4</td>
<td>5</td>
<td>8</td>
<td>5</td>
<td>2</td>
<td>3</td>
<td>10</td>
<td>3</td>
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</tbody>
</table>

**Figure 12.** Pair of proton boundary layers, in the same format as Figure 11.

### 6.3. Magnetic Humps and Magnetic Holes

A “magnetic hump” (“magnetic hole”) is an isolated enhancement (depression) of $B(t)$ on a scale of the order of an hour with a well-defined maximum (minimum). Observations of magnetic holes in the solar wind were first discussed by Turner et al. (1977) and Fitzenreiter & Burlaga (1978) based on observations made at 1 au. These structures are ubiquitous in the solar wind, and they have been extensively studied (e.g., see the references in Avinash & Zank 2007; Avinash et al. 2009; Burlaga & Ness 2011). Magnetic holes and humps have also been observed in the distant heliosphere and the heliosheath, where the pressure of pickup ions dominates that of the proton pressure as referenced above.

A model by Avinash et al. (2009) that best describes the observations of magnetic holes and humps observed in the distant heliosphere and heliosheath identifies magnetic holes and humps as soliton-like structures. This MHD-Hall model includes pickup protons as well as protons, electrons, and neutral hydrogen. All of these components must be considered in the distant heliosphere and the heliosheath (Zank 2015). Avinash et al. show that magnetic holes and humps can only exist if either the Mach number of the pickup protons $\lesssim 1$ and greater than that of the solar wind ions (which is the case for the heliosheath), as shown by the observations of Richardson et al. (2008) or the Mach number of the pickup ions is less than the Mach number of the solar wind ions which must be $\lesssim 1$. 
The model of Avinash et al. predicts that within the heliosheath, the magnetic holes and magnetic humps are pressure balance structures. However, non-propagating tangential discontinuities are also pressure balance structures, so that even if we could measure the total plasma and magnetic field pressures, and could not distinguish between magnetic holes and magnetic humps and tangential discontinuities. One might expect to observe an anti-correlation between the plasma density and the magnetic field for magnetic humps and magnetic holes. We did search for an anti-correlation between the thermal plasma density and magnetic field, but we found no such relationship in the V2 data for 2012. However, since the thermal plasma density is much smaller than the pickup density, this observation does not exclude the possibility that magnetic holes and magnetic humps are pressure balanced structures. Unlike classical solitons, magnetic holes and magnetic humps are modified after they collide as a result of an exchange between kinetic energy and magnetic energy. Magnetic holes and magnetic humps can become unstable, producing oscillations that resemble those observed by Fitzreiter & Burlaga (1978) at 1 au and by Burlaga et al. (2007a) in the heliosheath. This model is an extension of the soliton models published by Baumgärtel (1999), McKenzie et al. (2001, 2004) and Avinash & Zank (2007).

An example of a magnetic hump in the heliosheath observed by V2 on day 250, 2012 is shown in Figure 14. The solid curve in this figure is a Gaussian fit to the observations that passes through the maximum of B. Here we use the form of the Gaussian distribution given by

$$B = B_o + A \times \exp \left\{-2 \times \left[\left(t - t_0\right)/c\right]^2\right\},$$

(2)

where A is the difference between the maximum or minimum value of B from the base value $B_o$, and $2c = \text{FWHM}/\ln(4)^{1/2}$ with FWHM equal to the full width at half maximum $A/2$.

From the Gaussian fit, we derive $B_o = B_{\text{min}}$, $R^2$, and $t_0$ (the time that V2 observed the maximum or minimum value of $B$ that is used to label a particular magnetic hump or magnetic hole). The beginning and end times of the magnetic hole ($t_1$ and $t_2$) were determined by inspection of the plots, and the difference between these two times were taken as the passage time $w$, which we measure in hours. The characteristic width of the structure is then $w \times V$, where $V$ is the solar wind speed measured by the PLS instrument on V2. Since tails of the magnetic humps are not always apparent in the observations, we cannot use the Gaussian distributions to determine the start and end times of the magnetic hole. Instead, $t_1$ and $t_2$ were estimated by eye, giving days 250.711 and 250.762, respectively.

Table 3 gives parameters describing the magnetic hump in Figure 14 and two other magnetic humps that were identified in the magnetic field data from V2 during 2012 on days 1.492 and 221.870. Using the observed start and stop times of each magnetic hump and the solar wind flow speed $V$ measured by the plasma instrument on V2 and assuming that the magnetic holes are not propagating with respect to the heliosheath plasma, we find that the average size of the magnetic humps was $L = V \times (t_2 - t_1) = (706,000 \pm 338,000) \text{ km} = (38 \pm 30) R_\odot$. The uncertainty expresses the range of values that were observed for the average size of the 3 magnetic humps, namely $10 R_\odot$, $79 R_\odot$, $25 R_\odot$ (see Table 3). The average size of the magnetic holes if they were propagating solitons would be larger than indicated in Table 3.

A “magnetic hole” is defined as a depression of $B(t)$ on a scale of the order of an hour with a well-defined minimum in $B$. A magnetic hole that moved past V2 on day 259.109 is illustrated in Figure 15. An approximate fit to the $B(t)$ for this magnetic hole is given by a segment of the Gaussian distribution. The beginning and end times of the passage of the magnetic hole were chosen by eye to be day 259.098 and 259.135, respectively. The $B$ at these times were 0.272 nT and 0.268 nT, respectively. The minimum magnetic field strength, $B_{\text{min}} = 0.169 \text{ nT}$ occurred on day 259.109. There was a small change in the magnetic field direction across this magnetic hole.

A pair of magnetic holes, shown in Figure 16, moved past V2 on day 81, 2012. Two partial magnetic holes can be identified with well-defined beginning and end points but no inflection points, as illustrated in Figure 16. Again, we find that neither magnetic hole has the form of a complete Gaussian distribution with asymptotic tails, but each of the two profiles can be described by a partial Gaussian distribution, as illustrated by the solid curve in Figure 16. A Gaussian fit to the first magnetic hole gives a good fit with $R^2 = 0.97$ with a minimum $B_{\text{min}} = 0.24 \text{ nT}$ on day 81.065. The beginning and end times of this magnetic hole were chosen as day 81.046 and 81.075, respectively. Similarly, a Gaussian fit to the second magnetic hole in Figure 16 is a good fit ($R^2 = 0.97$) with $B_{\text{min}} = 0.15$ and 0.23 nT on day 81.107 and 81.207, respectively. There was no change in the direction of the across either of these magnetic holes. The collision of two solitons was modeled by Avinash et al. (2009) for one special initial condition. It would be of interest to determine whether the two magnetic holes shown in Figure 16 could be described as a stage in the collision of two solitons using the same sort of model.

Figure 13. Pair of proton boundary layers by means of which $B$ decreased in two steps on day 71, in the same format as Figure 11.
Table 4 contains the information describing seven magnetic holes observed by V2 during 2012. As for the magnetic humps, we find that the average size of the magnetic holes is $L = V \times (t_2 - t_1) = (723,000 \pm 355,000) \text{ km} = (11 \pm 4) R_L$, assuming that the magnetic holes were not propagating relative to the heliosheath plasma. The size of the magnetic holes would be larger for a given observed passage time if the magnetic holes were propagating solitons. There was a significant change in the direction of $B$ across only one of the magnetic holes. There was no change in $B$ in five of the magnetic holes listed in Table 4.

The points in Figure 4(d) show the times that which magnetic holes and magnetic humps were observed. There were two clusters of magnetic holes and magnetic humps. One cluster was observed in a region with low temperature, density, magnetic field strength, and pressure. The common feature of both of the clusters is relatively high speeds, although there are also regions of high speeds without magnetic holes or magnetic humps. The GMIR is a region that is actively being compressed and is associated with increasing speeds. One might suppose that the compression is somehow generating magnetic holes and magnetic humps. It would be interesting to investigate whether the overtaking of systems of slower ejecta and corotating fast streams by a fast flow, which is known to be a source of a MIRs and GMIRs in the solar wind and heliosheath, can also produce magnetic holes and magnetic humps in these regions, which would presumably propagate as waves propagating away from the Sun nearly perpendicular to the magnetic field lines.

7. SUMMARY AND DISCUSSION

This paper presents a comprehensive discussion of the observations by V2 during 2012 when it was observing the effects of increasing solar activity associated with solar cycle 24 and the increasingly strong magnetic fields observed at 1 au. Strong magnetic fields associated with solar cycle 24 were first
strongest magnetic fields, 0.30 nT, were observed on day 249, and strong magnetic fields were also observed from day 234 to day 262. We have confidence in the observed values of the strong magnetic fields, because they occur near a magrol and because the direction of $B$ was close to the spiral field direction, indicating that the $BR$ component of $B$ was small. A long-lasting decrease in the CRI was observed when these strong magnetic fields moved past V2, as expected from the CR–$B$ relationship, giving further evidence of the reality of strong magnetic fields during this interval. Strong magnetic fields are expected in the heliosheath during the rising phase of the solar cycle and strong magnetic fields were also observed at 1 au at this time. The strong magnetic fields observed by V2 during 2012 were among the strongest magnetic fields observed in the heliosheath by either V1 or V2.

Since the total pressure (magnetic pressure plus plasma pressure) was large between day 213 and day 262, a period of nearly 2 solar rotations, we identify the structure associated with these strong magnetic fields as a “MIR.” This MIR could possibly be a GMIR extending around the Sun in the azimuthal direction and to relatively high latitudes, but we have no direct observations to prove this statement. V1 observed strong magnetic fields in 2012, but they may be related to the proximity of the heliopause. In principle, one can use the boundary conditions at 1 au and recent MHD models including pickup protons to calculate the evolution of this MIR through the supersonic solar wind, across the TS and within the heliosheath. Observations by the LECP instrument on V1 show that the total pressure in the GMIR, including the pressure of protons ($28 \text{ KeV} - 3.5 \text{ MeV}$), is less than the magnetic pressure in the GMIR. This is the first time that $\beta < 1$ has been observed in the heliosheath.

It has been observed that the magnetic flux $BVqR$ was conserved at V2 from 2007 to 2011 whereas it decreased with increasing distance at V1 from 2005 to 2011. This paper shows that the magnetic flux at V2 was constant during 2012, as it was earlier in the heliosheath, and there was no decline in the bulk speed during the year.

On smaller scales, it was observed that daily increments of $B$ could be described by a nearly Gaussian distribution (namely, a $q$-Gaussian distribution with $q = 1.2 \pm 0.1$), whereas increments of hourly averages of $B$ were accurately described by a $q$-Gaussian distribution with $q = 1.82 \pm 0.03$. Thus, the fluctuations of $B$ at small scales have the characteristics of weak intermittent compressive turbulence.

On the smallest scales, a relatively large number of current sheets was observed by V2. Eight isolated PBLs were observed, five associated with a decrease in $B$ and three associated with an increase in $B$. In addition, three pairs of PBLs were observed, each pair within a single day. The average size of the isolated PBLs determined from the passage times was $\approx 180,000 \text{ km} = 5 R_L$, where $R_L$ is the Larmor radius measured using the thermal speed of a pickup proton rotating in the magnetic field. Thus the size of the isolated PBLs is comparable to the thickness of the isolated PBLs observed at 1 au by Burlaga et al. (1977). The average size of the pairs of PBLs was $105,000 \text{ km} = 7 R_L$, which is comparable to the size of the isolated PBLs. The pairs of PBLs were embedded in larger structures, and consequently one cannot obtain the asymptotic value for the sigmoid fits. In any case, the size of the PBLs observed by V2 in the heliosheath during 2012 is
comparable to the thickness predicted by Lemaire & Burlaga (1976), namely (2 to 10) $R_L$.

The PBLs were modeled as current sheets across which the pressure is constant, the electric field is important and the currents are generated by drifts of protons. Most of the PBLs observed by V2 show little or no change in the magnetic field direction across the current sheet. Such current sheets cannot undergo magnetic reconnection, but they can remain in equilibrium with a characteristic size. One explanation for the dominance of current sheets with no change in direction is that these current sheets are the survivors of a much larger set of current sheets that were associated with a change in magnetic field direction but were gradually annihilated or converted to tangential discontinuities by magnetic reconnection of opposing components of the inhomogeneous $B$. Alternatively, the current sheets associated with no change in the magnetic field direction could be produced by a compressive effect.

A relatively large number of current sheets was observed by V2 during 2012 when the active Sun was injecting magnetic fields into the heliosphere in the form of ejecta and transient flows. It is known that such activity produces complex magnetic field configurations with strong magnetic fields as a result of dynamical interactions in the heliosphere. It is also known that these configurations relax to less disturbed conditions farther from the Sun, approaching the TS. Although the formation of tangential discontinuities has been modeled in great detail for magnetic fields near the Sun (Parker 1994), relatively little is known about formation and evolution of discontinuities and current sheets in the heliosphere. However, Parker (1994) showed that the magnetic field can relax only by forming tangential discontinuities and the corresponding current sheets, which would correspond to the PBLs that we have described.

V2 also observed magnetic holes and magnetic humps in the heliosheath during 2012. Such structures have been observed throughout the supersonic solar wind, in the heliosheath, and in planetary magnetospheres. We found that many of the magnetic holes and humps observed by V2 during 2012 have sharp boundaries, rather than being described by a pair of sigmoid distributions or a single Gaussian distribution. The average size of the magnetic humps was $(706,000 \pm 338,000) \text{ km} = (38 \pm 12) R_L$, and the average size of the magnetic holes was $(201,000 \pm 99,000) \text{ km} = (11 \pm 4) R_L$. The average size of magnetic holes is twice that of the average size of the PBLs, as has been observed in previous studies in the solar wind and heliosheath. The average size of the three magnetic humps was relatively large in this regard, with a thickness of $79 R_L$. The average size of magnetic holes and magnetic humps quoted above is based on the assumption that they are not propagating with respect to the heliosheath plasma. If they are solitons, they propagate perpendicular the magnetic field relative to the plasma, and the sizes would be larger than quoted. On day 81, 2012, V2 observed what appeared to be two partial magnetic holes, which might be two magnetic holes that were interacting with one another, assuming that they have the character of solitons. It would be of interest to determine whether the observed magnetic field profiles could be modeled by the interaction of two solitons.

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**REFERENCES**

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Baumgärtel, K. 1999, JGR, 104, 295

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**Table 4**

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<th>Event</th>
<th>nh</th>
<th>mh</th>
<th>mh</th>
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<td>56.2</td>
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<td>254.961</td>
<td>259.109</td>
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<td>0.96</td>
<td>0.97</td>
<td>0.97</td>
<td>0.94</td>
<td>0.94</td>
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<td>0.148</td>
<td>0.21</td>
<td>0.269</td>
<td>0.272</td>
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<td>0.212</td>
<td>0.153</td>
<td>0.268</td>
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<td>0.084</td>
<td>0.002</td>
<td>0.116</td>
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<td>$(B_2 + B_1)/2 nT$</td>
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<td>0.190</td>
<td>0.211</td>
<td>0.211</td>
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<td>$B_{min}$</td>
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<td>0.57</td>
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<td>98</td>
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<td>7</td>
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