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HIGH-ENERGY PARTICLES ASSOCIATED WITH SOLAR FLARES

by

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Abstract

High-energy particles, the so-called solar cosmic rays, are often generated in association with solar flares, and then emitted into interplanetary space. These particles, consisting of electrons, protons, and other heavier nuclei, including the iron-group, are accelerated in the vicinity of the flare. At present, these particles are observed at the earth or nearby by means of both direct and indirect observational techniques. By studying the temporal and spatial variation of these particles near the earth's orbit, their storage and release mechanisms in the solar corona and their propagation mechanism can be understood. The details of the nuclear composition and the rigidity spectrum for each nuclear component of the solar cosmic rays are important for investigating the acceleration mechanism in solar flares. The timing and efficiency of the acceleration process can also be investigated by using this information. In this paper, the various problems mentioned above are described in some detail by using observational results on solar cosmic rays and associated phenomena.

1. Introduction

Since the first observation on 28 February 1942 (Lange and Forbush, 1942a,b), more than one hundred solar cosmic ray events have been reported and investigated (e.g., Bailey, 1964; Obayashi, 1964; McCracken and Rao, 1970; Sakurai, 1974). It is now known that these cosmic rays range from Mev to Bev energies and that their behavior in both the solar envelope and interplanetary space is very complicated in space and time. Since these cosmic rays are generally produced by major solar flares, it is necessary to understand the flare mechanism and the associated phenomena such as X-ray, white-light and radio emissions as well as flare-associated particle acceleration processes. For example, after the association of wide band type IV radio bursts was discovered by Boischot and Denisse (1957) and Hakura and Goh (1959), it became clear that most flares accompanied by these bursts produced solar cosmic rays over a wide energy range (e.g., Fichtel and McDonald, 1967; McCracken and Rao, 1970). This relation suggested that high-energy electrons, say Mev energy, are generated in the so-called proton flares; the first direct observation of the electrons was made in 1961 (Meyer and Vogt, 1962). Since then, many Mev electron events have been observed by satellites (e.g., Cline and McDonald, 1968). Hence it is now known that "solar cosmic rays" consist of high-energy protons, alpha-particles, heavier nuclei and relativisitic electrons. This fact suggests that the acceleration mechanism of these particles is closely related to the physical condition of the flare region and the development of solar flares. For this reason, it becomes important to study the nuclear abundance of solar cosmic rays relative to that for the solar photosphere and corona (e.g., Biswas and Fichtel, 1964, 1965; Sakurai, 1965e; Lambert, 1967a). It also seems useful to refer to the solar cosmic ray isotopic composition, and associated gamma ray and neutron emissions because these can give us important information on the physical state of the accelerating regions.

In this paper, we, therefore, first consider the observational results of solar cosmic rays, taking into account various wave emissions associated with solar flares. In so doing, we also review the characteristics of solar proton flares. In order to understand the acceleration and propagation of solar cosmic rays, it is necessary to understand such features of solar cosmic rays as spectral forms, nuclear composition, isotopic abundance, and others in great detail. These features are considered using recent direct observational data obtained on rockets and satellites; their relations to solar cosmic ray generation mechanisms is also explored. Then a general discussion of solar cosmic ray acceleration mechanisms is given.

2. Solar cosmic ray phenomena (Bev and Mev particles)

Solar flares sometimes produce high-energy particles called solar cosmic rays. The characteristics of such flares are, in general, different from solar flares not associated with those particles. Here we shall consider the characteristics of solar cosmic ray events which are mainly detected at the earth using both direct and indirect methods.

2.1 Characteristics of Bev cosmic ray events

Based on the time profiles and the peak intensities of solar cosmic ray events, they are temporarily classified within two types, called "Unusual increases" and "Small increases." If the peak intensity of a solar cosmic ray increase is over 10% of the background intensity of galactic cosmic rays, following Forbush (1946), we call this event an "Unusual increase." Other cosmic ray events, which do not show an intensity increase higher than 10% are called "Small increases" (Kodama, 1962).

(a) Unusual increases

Forbush (1946) discovered the solar component of cosmic rays. Three solar cosmic ray events which occurred on 28 February and 7 March, 1942 and 25 July, 1946 were analyzed by him using the mu-meson data obtained by several globally distributed observatories. The time profiles of these events are shown in Fig. 2-1 (a) and (b). Forbush noticed that the magnitudes of the peak intensities of these events were dependent on the positions of the cosmic ray observatories. As shown in Fig. 2-1 (a), at the observatory at Huancayo which is located near the geomagnetic equator, the magnitude of the observed peak intensity was very small (7 March, 1942) or negligible (28 March, 1942 and 25 July, 1946). This fact indicated that cosmic rays of energy higher than ~15 Bev were hardly produced from solar flares.

The fourth cosmic ray event associated with solar flares was observed on 19 November, 1949. This event was the first to be detected by a neutron monitor which was thus shown to be very useful in studies of the low-energy component of solar cosmic rays (~ 4 Bev). The fifth event occurred on 23 February, 1956. The time profile of the neutron component observed at Chicago is shown in Fig. 2-2 (Meyer, Parker and Simpson, 1956; Simpson, 1960a). In this case, it was found that the pattern of the time profiles is to some extent different from one observatory to another over the earth. The initial phase of a solar cosmic ray event showed especially high anisotropies. Similar features were also found for the first four events and were explained by using the results calculated for the orbital motion of solar cosmic rays in the earth's magnetic field (e.g., Schlüter, 1951; Firor, 1954).

The solar flare on 23 February 1956, which produced cosmic ray particles, was accompanied by white light emission (Notsuki, Hatanaka and Unno, 1956) and by a type IV radio burst (Boischot and Denisse, 1957). It is known at present that these two emissions are good indicators for solar cosmic ray production in flares.

As is seen in Figs. 2-1 and 2-2, solar cosmic ray events usually start with a sudden increase of cosmic ray intensity just after the onset of solar flares. After it peaks, the intensity gradually decreases following a power low function of time such as $t^{-3/2}$ at first and then like an exponential, exp $(-t/t_o)$, later. A time profile such as this was found in the event of 23 February 1956 and was explained by means of a diffusion process in the inner solar system ($\lesssim$ 1.4 AU) (Meyer, Parker and Simpson, 1956). As we shall see later, the time profiles of solar cosmic ray events at the earth are highly dependent on the position and other characteristics of the associated flares, and on the physical state of the interplanetary space. Three more examples of these time profiles are shown in Fig. 2-3.

(b) Small increases

As shown in Fig. 2-3(c), the increase of solar cosmic ray intensity on 7 July, 1966 was only a few percent above the galactic background intensity, although the super-neutron monitor was in operation (Carmichael, 1968, 1969). This event, though small, was very important in the progress of solar cosmic ray physics since it occurred during the period of the Proton Flare Project (1-13, July, 1966). The increase of solar cosmic ray intensity was very small, even in the neutron monitor data. Thus, such events seem to hardly produce high energy particles which yield mu-meson secondaries. The first "small increase" was observed on 31 August, 1956 (Kodama, 1962).

Until recently, about twenty solar cosmic ray events had been observed. They are summarized in Table 2-1. It is clear from the table that the importance of the associated flares is usually 3 or 3+. The two exception are associated with small increase events. Two events which were associated with solar flares beyond the limb of the solar disk are described (20 November, 1960 and 28 January, 1967). Some flares produce an increase of cosmic ray intensity even if they occur beyond the limb. We have no data to identify an associated flare in the case of the 28 January 1967 event. Except for this one event, every solar cosmic ray increase was produced from a solar flare which was accompanied by wide frequency and type IV radio bursts. This observation indicates that high-energy electrons are accelerated simultaneously with solar cosmic ray protons and heavier nuclei (e.g. Ellison, McKenna and Reid, 1961).

It was once thought that solar flares of importance ≤ 2 could also generate solar cosmic rays of Bev energy, but their intensity increase was thought much smaller than for those summarized in Table 2-1. By analyzing statistically the neutron data of the Climax station, Firor (1954) concluded that an intensity increase occurred around 0900 L.T. in association with these small flares; however, Towle et al. (1959) and Ghielmetti et al. (1960) did not find any increase of cosmic rays associated with the small solar flares, although they had taken into account the local time effect of solar cosmic ray incidence.

(c) Increases Associated with Over-Limb Flares

The first evidence for a cosmic ray increase associated with an over-limb flare was obtained in the 20 November 1960 event. In this case, the associated flare was detected through the observation of the ascending Hα emitting clouds and X-ray and radio emissions (e.g., Zirin, 1964, 1965). As shown in Fig. 2-4, the rate of increase in this event seemed to be slow compared to the other unusual increase events shown in Figs. 2-1 and 2-3 (Carmichael, 1962). This can be interpreted by taking into account the diffusion process of solar cosmic rays across the interplanetary magnetic field.

The large increase of cosmic ray intensity on 28 January, 1967 was observed globally, but no likely responsible flare was observed. The observed characteristics of this event were very similar to those which were associated with solar flares of great importance on the western hemisphere on the sun (Lockwood, 1968). This event, therefore, seems to have been produced from a solar flare which occurred just beyond the west limb of the sun. It is very hard to say something about the position of this flare because we have no observational data on radio or X-ray bursts. Further evidence for over-limb events was proposed by Dodson et al. (1969a,b,c) during the period 16 to 19 July, 1966.

Many distinct Kev electron events associated with over-limb solar flares have been reported by Lin and Anderson (1967) and Lin (1970b). All over-limb flares which produced these cosmic ray and electron events occurred beyond the west limb of the solar disk.

(d) Statistical characteristics of Bev Events

Except for a few cases, solar cosmic ray events of Bev energy are associated with solar flares which occur on the western hemisphere of the sun (see Table 2-1). The distribution of positions of solar flares of importance 3 and 3+ on the solar disk, however, is not asymmetric over the solar disk at all; hence, we should say that the western excess is not casually related to the nature of solar activity, but is due to some controlling factors intervening between the sun and the earth. Historically, this observational fact gave a clue to the existence of the interplanetary magnetic field spiralling eastward from the sun. Furthermore, the travel time of solar cosmic rays from the sun to the earth tends to become shorter as the position of associated flares moves westward on the solar disk (Fig. 2-5) (Hakura, 1961 ; Kodama, 1962). We shall see later that this tendency is also seen in the case of Mev solar cosmic rays (Sakurai and Maeda, 1961). These results suggest that the position of solar flares around the region 60^{o} - 90^{o} west on the sun is most favorable for the propagation of solar cosmic rays from the sun to the earth. By using these observations, Cocconi et al. (1958) first proposed the model of the interplanetary magnetic field which explains the western excess of solar proton flares. Initially, the role of solar cosmic ray study was very important in the determination of the physical state of the interplanetary space. The study of solar cosmic rays was to be useful tool for the understanding of the interplanetary magnetic field configuration (e.g., McCracken, 1962a-c; Obayashi and Hakura, 1960a,b; Parker, 1963a).

2-2 Characteristics of Mev Cosmic-Ray events

Solar cosmic rays of Mev energy were first investigated using the data obtained by riometers and ionosondes. Riometers observed the relative increase of the absorption of galactic background radio emissions in the high-frequency band (~30 MHz), while ionosondes measured the increase of f-min due to the absorption of vertical sounding radio waves. However, the first evidence on the production of Mev cosmic rays from solar flares was given by Bailey (1957) on the basis of an analysis of the characteristics of the HF backscatter wave absorption in the case of Bev cosmic ray event on 23 February, 1956. During the period (1956-1962), these radio techniques began to be exclusively used to study solar cosmic ray phenomena of Mev energy (e.g., Leinbach and Reid, 1959; Hakura and Goh, 1959; Hultqvist, 1959; Avignon, Pick and Danjon, 1959; Reid and Leinbach, 1959; Obayashi and Hakura, 1960a,b; Hakura, 1961 ; Sinno, 1961, 1962; Sakurai and Maeda, 1961; Leinbach, 1962).

During the same period, new observing techniques using rockets, satellites and balloons were developed to directly detect solar cosmic ray particles. The first direct observation of solar cosmic rays was made by Anderson (1958) using an ionization chamber on board a balloon flight. Successively, Anderson and his colleagues extensively studied many Mev solar cosmic ray events by using balloon-born instruments (Anderson, Arnoldy, Hoffman, Peterson and Winckler, 1959; Ney Winckler and Freier, 1959; Freier, Ney and Winckler, 1959; Winckler, Peterson, Hoffman, and Arnoldy, 1959; Anderson and Enemark, 1960; Winckler, 1960; Winckler, May, and Masley, 1961; Winckler, Bhavser, Masley, and May, 1961). The first satellite observation of solar particles was done by Rothwell and McIlwain (1959) on the Explorer 6 satellite. These direct observations showed that solar cosmic rays consist primarily of protons, but did not at all give information on the heavy nuclei. For this reason, the Mev cosmic ray events were often called "Solar proton events."

The heavy nuclei were first successfully observed by Fichtel and Guss (1961) using the emulsion technique on board rockets. At present, it is known that the composition of solar cosmic rays is very similar to that of the photosphere of the sun and is, therefore, different from that of the galactic cosmic rays ((Li,Be,B) and iron-groups) (e.g., Biswas and Fichtel, 1964, 1965).

(a) <u>General Characteristics</u>

Solar cosmic rays of Mev energy cannot reach sea level in low and medium geomagnetic latitudes, but they do invade the polar cap regions. Because they lose most of their energy to ionization of the atmospheric constituents, their invasions are detected by ionosondes and riometers. Although these observations do not involve solar particles directly, they can be used to study the patterns of solar cosmic ray incidence over the polar cap regions, the time profiles of the solar cosmic ray flux in the vicinity of the earth and so forth (e.g., Hakura, Takenoshita and Otsuki, 1958; Obayashi and Hakura, 1960a; Reid, 1961, 1966, 1970; Hultqvist, 1965, 1969).

In general, an increase of the solar cosmic ray flux begins within ten minutes to several hours after the onset of an associated flare. Sometimes, the increase of the flux is delayed some ten hours after the onset of an associated flare and is generally associated with the beginning of an SSC geomagnetic storm. Based on his analysis of these two different patterns for the flux increases, Sinno (1961) found that the time profiles of the solar cosmic ray flux are highly controlled by the physical circumstance in the interplanetary space. He assigned the designations, F and S type events to these fast and slow increases. Fig. 2-6 shows these two types of events as seen in the f-min data (Sinno, 1961). This classification has been further extended in order to indicate the relationship between solar cosmic ray and geophysical phenomena (e.g., Hakura, 1961 ; Sakurai and Maeda, 1961; Leinbach, 1962; Obayashi, 1962, 1964).

By using the riometer data of several observatories in the north polar cap region, Leinbach (1962) studied the development of polar cap absorptions associated with solar flares and then developed the classification scheme shown in Table 2-2. As shown in Fig. 2-7, his classification system is more complex than that given by Sinno (1961). Obayashi (1964) and Sakurai (1965a) gave the definitions shown in the last column in Table 2-2, which are an extension of the classification given by Sinno (1961) and Hakura (1961a). Solar cosmic ray events of the F and F* type are often called the "prompt onset type" because the increase of cosmic ray flux at the earth begins within several hours after the onset of the associated solar flare and the flux reaches its maximum before the SSC geomagnetic storm starts (Fig. 2-7 (a) and (b)). In the cases of the events shown in Fig. 2-7 (c) and (d), the peak flux is reached just after the SSC geomagnetic storms occur. Hence these events are closely connected with the propagation of the energetic plasma clouds responsible for the production of SSC storms and are often called "delayed onset type" events. The different behaviors of the prompt and delayed onset type events is closely related to the energy of the solar cosmic rays as shown in Fig. 2-8. This figure indicates that the particles of energy lower than 30 Mev partly propagated with the plasma cloud which produced the SSC storm on 30 September, 1961(Obayashi, 1964). Such delayed onset type events have been extensively studied using satellite data (e.g., Bartley, Bukata and McCracken, 1966; McCracken, Rao and Bukata, 1967; Rao, McCracken and Bukata, 1967; Anderson, 1969).

With the advancement of satellite observational techniques, the lower limit of the observable cosmic rays energy has been gradually reduced. At present, very weak cosmic ray events produced by protons of energy $\leqslant$ 1 Mev are detectable by satellites and deep space probes. Such events have never been detected by ground based experiments, nor by satellite observations during the period 1958-1961. This progress necessitated a new definition and classification scheme of solar cosmic ray events.

The classification of polar cap absorptions was initiated by Obayashi (1960) based on his analysis of the characteristics of f-min increases associated with solar proton events. This classification was adopted in tabulating the importance of proton events observed from 1957 to 1967 by Obayashi et al. (1967) and Hakura (1968). As shown in Table 2-3, solar proton events are divided into five classes defined as importance 1-, 1,2,3, and 3+. The relation of these importances with f-min increase recorded at Resolute Bay is indicated in this table.

In order to compare the records for f-min increases with the directly observed solar cosmic ray flux and others, we would like to find a formula or transform for relating these two different data. Such studies were done by Bailey (1964), Van Allen, Leinbach and Lin (1964), Fichtel, Guss and Ogilvie (1963), Masley and Goedeke (1968) and Juday and Adams (1969), but it is now clear that we are hardly able to obtain a unified formula transforming between these two data. Recently, as shown in Table 2-4, Smart and Shea (1970) have proposed a new classification system of solar cosmic ray events to synthesize all data recorded by different methods. If we refer to this classification, "Small increase" events of Bev cosmic rays correspond to the index numbers 1 and 2.

(b) Statistical Characteristics

Type IV radio bursts are usually associated with solar proton flares. In the analysis of the time profiles of Mev solar cosmic rays, we have seen that there exist two different type events, prompt and delayed onset types. It has been found that the cause of these two types of events is related to some characteristics of the associated flares. The dynamic spectra of type IV radio bursts, for example, are characteristically different between them as shown in Fig 2-9 (Sakurai and Maeda, 1961; Sakurai, 1963). Castelli et al. (1967, 1968) have shown that the peak flux spectra for prompt onset type events are good indicators for solar cosmic ray production.

By examining the development patterns of type IV radio bursts, Hakura (1961) showed that the intense emission of the microwave component is usually associated with solar flares which produce the prompt onset type cosmic ray events. If the emission of this component is low, associated flares do not produce prompt onset events, but are sometimes accompanied by delayed onset type events (Fig. 2-10). Figs. 2-10 (a) and (b) are respectively associated with the prompt and delayed onset type events. It has also been shown (Sakurai, 1967a) that the rise time of the Hα brightening in flares is shorter for the former than for the latter events. Hakura (1966) showed that the development patterns of SID's, especially SWF's, are different for these two types of cosmic ray events. An example of an SWF event is also schematically shown in the above figure. This result suggests that the emission processes of thermal X-rays from solar flares are strongly connected with those of type IV radio bursts and Hα flare emissions (Hakura, 1966; Sakurai, 1967a).

From Figs. 2-7 (c) and (d), we can infer that the delayed onset type cosmic ray events are closely related to the energetic storm plasma clouds which produce SSC geomagnetic storms (e.g., Sinno, 1961; Obayashi and Hakura, 1960b; Leinbach, 1962). This fact suggests that the main portion of the Mev cosmic ray particles from solar flares is transported by these clouds being trapped in them. Parker (1965), however, has interpreted such cosmic ray events by assuming a local acceleration of energetic particles due to the interaction of the interplanetary plasma with the blast waves propagating into interplanetary space. A similar idea was recently considered by Fisk (1971), who showed that the increase of the particle flux incident to the polar cap regions would only be observed just before or at the time of the sudden commencement of the geomagnetic storms. However, as is clearly seen in Figs 2-7 (c) and (d), the enhancement of particle flux is generally observed continuously after the onset of the sudden commencement and sometimes reaches its maximum during the main phase of the geomagnetic storm (e.g., Obayashi, 1962,; Leinbach, 1962).

As we have discussed above, there exist the western excess of Bev proton flares and the shortness of the travel times of Bev particles from such western events. Similar tendencies are also found in the propagation of Mev solar cosmic rays, but the interpretation, in this case, is complicated by the distinct F, F* and S types of events which can occur. The western excess was first shown by Reid and Leinbach (1959) and Obayashi and Hakura (1960a). This excess is most conspicuous for F type events. In fact, most of these events are associated with solar flares which occur on the western hemisphere of the sun as shown in Fig. 2-11 (Obayashi, 1964). Because of the similarity of their time profiles (Fig. 2-7 (a)), are thought of as small-scale versions of the "unusual increase" events of Bev solar cosmic rays. Fig. 2-11 also shows that the travel time of Mev cosmic ray particles from the sun to the earth tends to become shorter as the position of the associated flares moves westward over the solar disk and that this tendency is seen independent of the type of cosmic ray events (Sakurai, 1960; Obayashi and Hakura, 1960a,b). However, the travel time for F* type events is slightly longer than that for F type events (Sakurai, 1965a).

The differences between the F and F* type events seems to be related to physical circumstances in interplanetary space. The configuration of the magnetic field is of primary importance. This configuration is affected strongly by the plasma clouds ejected from solar flares (Gold, 1959; Steljes et al., 1961; Schatten, 1970). The ejection of these clouds several days before the occurrence of a proton flare at the sun seems to be an important cause of the different F, F*, and S types of cosmic ray events. Sakurai (1965a) found that the ejection of such clouds plays an important role on the origin of these types of events. In fact, F type events are produced when the time interval between the onset of the SSC geomagnetic storm produced by these clouds and the start of the next solar cosmic ray event is shortest (Fig. 2-12). S type events are associated with the passage of the plasma clouds which are produced from the same solar flares which generate the cosmic ray particles.

2-3 Generation of High Energy Electrons

Solar proton flares are usually associated with solar radio bursts of spectral types II, III and IV, but the most important association is with wide band type IV radio bursts. The type IV bursts are currently believed to be emitted from mildly relativistic electrons in the sunspot magnetic fields (Boischot and Denisse, 1957; Boischot, 1959; Takakura, 1959, 1960a,b). The direct observation of relativisitic electrons was first made by Meyer and Vogt (1962) on 18 July, 1961, on a day during which, as is shown in Table 2-1, a proton flare occurred. Cline and McDonald (1968) found that most proton flares are associated with the production of relativisitic electrons. At present, we have many observational data for these electrons (e.g., McDonald, Cline, and Simnett, 1969; Simnett, Cline, Holt, and McDonald, 1969; Detlawe, 1970; Simnett, 1971, 1972).

Just as in the case of proton flares, solar flares which generate relativisitic electrons observed at the earth and its vicinity are almost always located on the western hemisphere of the sun. This fact is also explained by taking into account the configuration of the interplanetary magnetic field. The maximum intensity of thosê electrons at the earth's orbit varies largely from < 0.05 to ~3 in the unit of electrons cm^{-2} sec^{-1} str^{-1} (Cline and McDonald, 1968). An example for the time profiles of 3 Mev solar electons is shown in Fig. 2-13. As shown in this figure, for the 7 July 1966 event, the travel time of these electrons is shorter than that for 16-80 Mev solar protons.

The acceleration of these relativisitic electrons is done during the explosive phase of solar flares (e.g., Svestka, 1970; Sakurai, 1971b). A portion of the accelerated electrons would be the source for the emission of type IV radio bursts. It should also be noted that solar flares which produce relativistic electrons usually occur in sunspot groups which are active in the emission of metric continuum radiation.

Solar electrons of Kev energy are more frequently produced than relativistic electrons. These electrons ($\gtrsim$ 45 Kev) were first discovered by Van Allen and Krimigis (1965) and Anderson and Lin (1966). Solar flares which produced these electrons are almost always accompanied by microwave impulsive and type III radio bursts (Sakurai, 1967c). This fact indicates that the origin of these bursts is closely connected with the production of these Kev electrons. The time profiles of Kev electrons observed by satellite born instruments are shown in Fig. 2-14. (Anderson and Lin, 1966). The increase of the electron flux usually starts several to some ten minutes after the onset of an associated flare and the event of this type is called a "Prompt" electron event.

As in the case of solar proton flares, the positions of solar flares producing Kev electron events are almost always located on the western hemisphere of the sun (Lin, 1970b; Sakurai, 1971a). However, there is no clear dependence on solar longitude of the travel time of these electrons from the sun to the earth.

Sakurai (1971a) showed that most solar flares associated with Kev electron production occurred in the sunspot groups which were active in the emission of metric continuum radiation (often called type I noise). This result suggests that the energetic electrons responsible for the continuum radiation are accelerated to even higher energy in solar flares and then ejected.

We have, so far, considered the prompt onset type electron events; but, as reported by Anderson (1969), there also exists the "delayed" onset type events. In these events, the flux of Kev electrons begins to increase several hours or less before the sudden commencement of geomagnetic storms. This characteristic is very similar to that of S type Mev solar cosmic ray events.

3. Characteristics of Solar Proton Flares

In this section, we shall consider various phenomena associated with solar proton flares and the relation of these flares to the configuration of sunspot groups.

3-1 Optical characteristics and white light emissions

At present, we have data on proton flares since 1954 (e.g., Fritsova-Svestkova 1966; Hakura, 1968). Systematic observation of solar cosmic rays started with the beginning of the IGY (July, 1957-December, 1958). Solar cosmic ray events were observed 128 times during the 14 years from 1954 to 1967 (Hakura, 1968). The distribution of the importance of proton flares for this period is shown in Fig. 3-1. This distribution indicates that the generation of solar cosmic rays are mainly associated with solar flares of importance $\geq$ 2N. Most Bev cosmic ray events were accompanied by solar flares of importance 3 and 4 (or 3+). These facts suggest that the generation of solar cosmic rays is closely connected with the Hα brightening process.

White light emissions are often observed in association with solar flares which produce Bev cosmic ray particles. McCracken (1959) first remarked that these emissions are one of the important indicators of such flares. The white light emissions are mainly emitted during the explosive phase for several to some ten seconds.

At first, in order to explain those emissions, Stein and Ney (1963) suggested a possibility of synchrotron emission from relativistic electrons, but later this was abundaned (see, Korchak, 1965; Svestka, 1966). At about the same time, Svestka (1963) considered the effect of H^- ions on the generation of those emissions. Recently, Najita and Orrall (1970) have estimated the contribution from the bombardment of accelerated Mev protons into the photosphere. This process seems to be promising, but we still have to say that we have no theory which wholly explains the white light emissions associated with proton flares.

Dodson and Hedeman (1959, 1960) first pointed out that, to some extent, the umbral regions of sunspots are covered by the Hα brightening areas in association with solar proton flares. This phenomenon is now called "umbral coverage" of the flare brightening area, and its cause seems to be closely related to the configuration of sunspot magnetic fields in which proton flares occur. Ellison et al. (1961) and Avignon et al. (1963, 1964) found that the occurrence of proton flares is associated with sunspot groups of some unusual configurations.

According to Malville and Smith (1963), the frequency of occurrence of solar proton flares and associated type IV radio bursts increases with the percentage of umbral coverage. Furthermore, this percentage varies with the longitude position of the associated flares (Sakurai, 1965a, 1970a). Solar proton flares are closely associated with the later formation of loop prominences. According to Bruzek (1964a,b), most proton flares are accompanied by the formation of loop prominence systems after the associated flares cease. Thus the formation of these systems seems to be an after effect of proton flares (e.g., Jefferies and Orrall, 1965a,b).

In association with solar proton flares, dark halos are often formed above the Hα flare regions and are now defined as "flare nimbuses" (Reid, 1963). Generally speaking, such halos appear in correlation with the striation patterns in the Hα observations. Ellison et al. (1960a,b, 1961) found a close association of the flare nimbus phenomenon to type IV_m B emissions.

According to Reid (1963), the characteristics and development of flare nimbuses are summarized as follows:

when the efflux of plasma from the flare region is of sufficient energy to carry the local magnetic fields away from the sun, then the magnetic energy withdrawn from the active regions may result in the bright chromospheric intensity, with a consequent disappearance of any form of nearby striation pattern.

Later on, these magnetic fields are extended from the flare regions to form the stationary sources of type IV_m B emissions. These nimbuses are most conspicuous 20-30 minutes after the flare maximum; their length scale and duration are $2\text{-}4\times10^5$ Km and 1-2 hours, respectively.

3-2 Characteristics of associated radio and X-ray bursts

As shown in Fig. 3-1, (in [3-1]), solar cosmic rays are mainly generated in solar flares of importance $\geq$ 2N. These flares are generally accompanied by radio bursts of spectral type II, III and IV and microwave radio bursts. Type IV radio bursts are especially important indicators of the generation of solar cosmic rays. As shown in Fig. 3-2, these bursts consist of four components from microwave to metric (or decametric) wave frequencies.

In general, as indicated in figure 3-3, the microwave emission (IV_μ) starts during the explosive phase and then, as time goes on, the lower limit of the emission frequency tends to decrease to metric frequencies via decimetric frequencies.

Figure 3-3 shows that for the July 7 1966 event, at first, a microwave radio burst occurred and then the microwave component of a type IV radio burst followed, but, in general, we are unable to separate these two bursts in the frequency range $10^3 - 10^4$ MHz. It seems, however, that the emission of the type IV microwave component begins in the explosive phase of solar flares. As shown in this figure, both microwave and decimetric components (IV_μ and IV_{dm}) reached maximum simultaneously, but the metric component (IV_m) showed a frequency drift in peak flux, which is similar to that which was observed for the type II radio burst. The speed of the source for this burst was estimated to be ~ 1,000 $Kmsec^{-1}$ by considering the frequency drift rate (Sakurai, 1971d).

The peak flux spectrum for this event was obtained by Castelli, Aaron and Michael (1967), and is shown in

Fig. 3-4. This is a typical example of these spectra (Castelli et al., 1967, 1968; Sakurai, 1969c). In general, these spectra show a deep depression of the peak flux at decimetric frequency around 1000 MHz and are therefore called "U-shaped" spectra (Castelli et al., 1967). Hence we are now able to forecast the production of solar cosmic rays by observing the peak flux spectra of type IV radio bursts.

Hard X-ray bursts are usually associated with microwave impulsive radio bursts. As shown in Fig. 3-5, for the solar flare on 7 July 1966, hard X-ray emissions were observed and their time profile was very similar to that of the microwave impulsive burst. This similarity suggests that the causes for the X-ray and microwave bursts are related to each other (Anderson and Winkler, 1962; de Jager and Kundu, 1963; Kane, 1969).

It is known that gamma-rays are sometimes emitted from solar flares (Anderson and Winckler, 1962). The emission of gamma-rays seems to be related to some nuclear reactions produced by high energy protons in the solar atmosphere (e.g., Lingenfelter and Ramaty, 1967). The discussion on the gamma-ray emission will be later given in section 4.5.

Soft X-rays are also emitted from solar proton flares. Since these X-rays are thermally produced in or near the flare regions, the rise time of these emissions are probably related to the rapidness of the thermalization of flare plasma. Hence this time can be correlated to the rise time of the Hα brightness. In fact, these two rise times tend to become shorter as the energy of accelerated protons becomes higher (Sakurai, 1970b) (Fig. 3-6). This result indicates that the speed of flare development is an important factor for the acceleration efficiency of flare particles.

3-3 Sunspot configuration

The regions where solar flares occur are closely related to the configuration of sunspot magnetic fields. Recently, it has been found that the solar proton flare mechanism is associated with the formation of δ-type sunspot groups. We shall here review some characteristics of sunspot groups which produce proton flares.

a) Relation to sunspot type:

By examining the types of sunspot groups which produced Bev proton flares, Ellison et al. (1961) concluded that they were generally classified as $\beta\gamma$ or γ type. As Noyes (1962) showed later, about 70 percent of the Mev proton flares also occurred in sunspot groups which were classified as $\beta\gamma$ or γ type. Thus, sunspot groups of $\beta\gamma$ and γ types are very important for the production of both Mev and Bev proton flares.

Solar flare occurrence frequency is highly dependent on the age of sunspot groups. As shown in Fig. 3-7, sunspot groups are most active for flare production when they are of type F (Waldmeier, 1957). In fact, most proton flares have occurred during the periods while the associated sunspot groups were of types E and F (Anderson, 1961).

The results of the above paragraph indicate that both Mev and Bev proton flares generally occur in sunspot groups of $\beta\gamma$ or γ types, the phases of which are of types E or F.

By taking into account the new classification by Künzel (1960) of sunspot groups, Warwick (1966) examined the magnetic polarity distribution of the groups which produced proton flares and found that both south and north polarity regions coexist in the umbrae of these groups, i.e., they are classified as δ-type. Sakurai (1967b, 1969b, 1970a) also concluded that most proton flares occurred in the sunspot groups of δ-type. These results indicate that the morphology of sunspot groups plays an important role in the origin of proton flares. The formation of neutral regions and sharp magnetic gradients in sunspot magnetic fields are apparently related to the δ-type sunspot groups.

b) Configuration of sunspot groups:

The importance of the sunspot group configuration for the production of solar proton flares was first pointed out by Avignon et al. (1963, 1964), who classified sunspot groups into three different types defined as A, A' and B (Fig. 3-8). The first two configurations, A and A', are important for the occurrence of proton flares. As shown by the shaded areas in figure 3-8, the Hα brightening areas cover the umbral regions of configuration A and A' sunspot groups. In addition, Avignon, Caroubalos, Martres and Pick (1965) quantitatively measured the ratios of the distances between the two main sunspots to the diameters of the umbrae for these spots. According to them, the occurrence frequency of proton flares tends to become higher as this ratio becomes smaller. This result is very similar to that obtained by Severny (1964a, 1965).

The distribution of magnetic fields above sunspot groups was investigated extensively by Severny (1957, 1958) by measuring the field intensity along the line of sight ($B_{||}$). He found that solar flares tend to occur along the neutral line for this field component (i.e., $B_{||} = 0$). Later on, he used this fact to develop the neutral line discharge theory of solar flares (e.g., Severny, 1960, 1964a,b, 1965).

Recently, Sakurai (1967b) found that the magnetic polarity distribution of sunspot groups which produced solar proton flares was unusual in comparison with that for most sunspot groups which did not. Two examples of such unusual polarity distributions are shown in Fig. 3-9 (Sakurai, 1972a). In this figure, the chain line indicates the neutral line for $B_{||}$. Type I describes a sunspot group in which the polarity distribution is reversed from the normal sunspot groups which have the north pole region in the preceding spot. The type II spot is rather normal compared to the type I spot, but the north pole region is located northward of the following south pole region.

The formation of such polarity distributions for sunspot groups can be associated with the proper motion of the sunspot groups themselves (Sakurai, 1967b, 1969b).

c) The change of the gradient of sunspot magnetic fields during solar flares:

Severny (1958, 1959) first pointed out the importance of changes in the gradient of sunspot magnetic fields near the "neutral line" during triggering of solar flares. As useful as the measurements might be, however, it is very difficult to see the change of the field gradient associated with solar flares. In different attempts at measuring this change for the sunpot group which produced proton flare on 16 July, 1959, Howard and Babcock (1960) obtained a result quite different from that by Howard and Severny (1963). Howard and Severny found that the field intensity changed by a factor of 3 during this flare, whereas Howard and Babcock did not see a detectable change in B in and near the flare site. At present, we cannot say conclusively whether or not the magnetic fields of sunspot groups change significantly during solar flares. (see the discussion between Howard (1963, 1969) and Sivarman (1969)).

On the other hand, Severny (1964a,1965) did find that the gradient of sunspot magnetic fields is closely related to some characteristics of solar proton flares. According to Sakurai (1972a), the importance of polar cap absorption, which is numerically related to the integral flux of Mev solar protons, tends to become greater as the gradient of associated sunspot magnetic fields become sharper (see Fig. 3-10). In this analysis, the data on sunspot magnetic gradients was taken from Severny (1965).

d) Rotating motion of sunspot groups:

Sunspot groups which have produced solar proton flares have often rotated counterclockwise (clockwise) in the northern (southern) hemisphere for several days before proton flares occurred (Sakurai, 1967b, 1969b). Sawyer and Smith (1970) pursued such rotating motion for the sunspot group MacMath No. 9760 in November, 1968 by measuring the day-to-day variation of the magnetic axis of this group and found that this axis gradually rotated counterclockwise (Fig. 3-11). By examining this rotating motion in the sunspot groups which produced proton flares, McIntosh (1969, 1970) also found that these sunspot groups were accompanied by counterclockwise and clockwise rotating motion in the northern and southern hemisphere, respectively. Generally speaking, such rotating motion is usually observed for sunspot groups which produce proton flares. This fact indicates that this motion is related to the triggering of solar proton flares (e.g., Sakurai, 1970a).

As remarked by Hale (1908, 1927) and later by Richardson (1941), this rotating motion seems to be related to the formation of the fibrille structure in the chromosphere as observed by the Hα line. For the latter seems to be controlled by sunspot magnetic lines of force. Hale (1927) and Richardson (1941) have pointed out the importance of the Coriolis force for the formation of the fibrille structure in the convective motion in the photosphere.

3-4 Development of solar proton flares

A typical solar flare of importance 4 (or 3+) develops through four distinct phases defined as 1) Pre-explosive phase (or Precurser), 2) Explosive phase, 3) Main phase and 4) Late phase. The pre-explosive phase is related to the increase of various solar activities before the onset of a solar flare or explosive phase. Before the onset, microwave and soft X-ray emissions are generally intensified from sunspot groups in which these flares will occur. Sunspot structures also vary with the appearance of satellite sunspots (Rust, 1969). The Hα plage bright spots and the activation of dark filaments are often observed in the optical range. These features are generally thought of as the phenomena preceding the occurrence of solar flares. However, all sunspot groups with these features do not produce solar flares.

When the explosive phase does occur, the Hα plage bright spots are enhanced very quickly and the sudden expansion of the Hα bright regions also begins. Simultaneously, the emission of both microwave impulsive and hard X-ray bursts starts and associated type III radio bursts are often observed.

It is now thought that the acceleration of solar cosmic ray protons and heavier nuclei and high-energy electrons occurs during the explosive phase. Some part of these electrons, in the energy range 100 Kev-10 Mev, becomes the source of type IV radio bursts of microwave frequencies (IV_{μ}). In the flare regions, furthermore, hydromagnetic disturbances like the Moreton waves are generated during the explosive phase and then propagate outward. The development of these phenomena associated with solar flares are summarized in Table 3-1.

Ellison et al. (1961) have studied the distribution of the Hα brightening areas over the sunspot groups which produced solar proton flares. According to their analysis, these areas are usually distributed as in the two cases shown in Fig. 3-12 a) and b); namely, the two Hα brightening areas form between two main sunspots of different polarity. Krivsky (1963a,b) has pointed out that, just after the onset of the explosive phase, these brightening areas pass through a Y-shape, for which he gave the definition of the Y-phase. This phase may be characteristic of proton flares.

3-5 Energetics of solar flares

Various phenomena such as optical, EUV, X-ray, radio and particle emissions are associated with a solar flare of great importance (say 3+ or 4). By analyzing the energetics of the solar proton flare of 23 February 1956, Parker (1957) first estimated the total amount of the energy released from this flare to be $\sim 10^{32}$ ergs. Since the typical duration such great flares is ~30 minutes, the rate of flare energy release is estimated as $10^{28} - 10^{29}$ ergs. sec^{-1}. This energy is expended in many different ways in various particle and electromagnetic emissions. Later, Ellison (1963) also estimated the total amount of flare energy and concluded that most of the energy goes into plasma cloud and visible light emissions. The amount of energy expended for these emissions is of the order of 10^{32} ergs. The energy released as radio, X-ray and particle emissions is smaller at least by a factor 10 than the above two emissions. According to Ellison (1963), the total energy of solar cosmic rays is $\sim 10^{30}$ ergs. Thus we can say that a typical solar proton flare releases a minor amount of flare energy in the form of solar cosmic rays.

Recently, Bruzek (1967) examined the partition of flare energy to various emissions in more detail. His results are shown in Table 3-2. Bruzek also concludes that a typical flare releases a total of 10^{32} ergs. Hence we may conclude that the amount of energy released from a typical flare is $\sim 10^{32}$ ergs (de Jager, 1969). Since the duration of such a flare is estimated as ~30 minutes, the rate of energy release is calculated to be $\sim 10^{29}$ ergs sec^{-1}; this amount corresponds to about 10^{-4} of the continuum emission from the quiet sun (see Zirin, 1966).

The volume of a flare region is estimated to be $\sim 10^{29}$ cm^3 by using the observed characteristic size and height. Thus 3×10^3 $ergs/cm^3$ must be released in the flare region. Our next problem is to find out how so much energy can be released in the flare region and what the energy source is.

The energy of solar flares seems to be supplied from the region where flares occur or from the vicinity of the flare region. As the source of flare energy, several possibilities are considered; they are (1) thermal, (2) gravitational, (3) magnetic, (4) rotational energies and (5) the energy stored by high-energy particles.

Thermal and gravitational energy:

Excess material is trapped by sunspot magnetic fields in the chromosphere and the corona over the active regions, and this provides thermal and gravitational energy. The thermal energy can be estimated if the number density and temperature of this material are known. Since the electron number density and the optical temperature of the flare regions are respectively $\sim 10^{12} - 10^{13}$ cm^{-3} and $\sim 10^{3} - 10^{4}$ ^{o}K, this energy is estimated as $\sim 10^{-1} - 10^{1}$ ergs cm^{-3}. If we assume the height of the flare regions to be $10^{4} - 5\text{x}10^{4}$ Km above the photosphere (Warwick, 1955), the thermal energy stored in the vertical column of the area 1 cm^{2} becomes of the order of $10^{8} - 5\text{x}10^{9}$ ergs cm^{-2}. As the characteristic size of the flare regions is typically $\sim 10^{10}$ cm, the total thermal energy reaches $10^{28} - 5\text{x}10^{29}$ ergs. Since the amount of energy $5\text{x}10^{29}$ ergs is obtained with $T=10^{4}$ ^{o}K and $n_e=10^{13}$ cm^{-3}, this value seems to be an upper limit for the thermal energy stored in the flare regions (see Sweet, 1969).

By using the plasma density referred to above, the gravitational energy is estimated to be $\sim 10^{25} - 10^{26}$ ergs.

This result means that the contribution of the gravitational force to the energy of a solar flare is negligible. The importance of a gravitational energy was once suggested by Sturrock and Coppi (1966), but it is now clear that this energy is unable to supply the whole energy of solar flares.

Magnetic energy:

Solar flares generally occur in sunspot magnetic regions. The intensity of sunspot magnetic fields is from 10^2 to several X 10^3 guass in most cases. If we assume that the field intensity is 500 gauss over the average flare region of 10^{29} cm^3, the total energy of the sunspot magnetic fields is calculated to be $\sim 10^{33}$ ergs. Thus the field energy is sufficient to explain the flare phenomenon even if only a part of it is released (e.g., Ellison, 1963, 1964; Parker, 1957; de Jager, 1969).

If the intensity of the magnetic field is 500 gauss, the field energy density is $\sim 10^4$ ergs cm^{-3}. This energy density is high enough to explain the flare energy expended per unit volume in the flare regions. Therefore, it seems that we could explain the flare phenomenou by taking into account the conversion of field energy to flare energy if a mechanism for the conversion could be found.

At present, it is believed that such a high amount of sunspot magnetic energy could be stored in the form of a force-free field configuration (Parker, 1957; Gold and Hoyle, 1960). Moreton and Severny (1968) have observed that the distribution of vertical electric currents in sunspot groups can be explained on the basis of a force-free magnetic configuration.

Ambient high-energy protons:

The importance, to the generation of solar flares, of Mev protons trapped in sunspot magnetic fields has been suggested by Elliot (1964,1969). According to him,these protons are the source of flare energy which is transfered to various other forms of flare energy during the flare. If the protons are trapped well before the flare onset, they may be a source for background emissions of gamma rays and neutrons. However, we have never observed such emission from the quiet sun.

Interception and storage of magnetoacoustical flux:

Acoustic and hydromagnetic wave energies are steadily transported up to the chromosphere and the corona. This phenomenon is also observed in and above sunspot regions. Some amount of this energy seems to be trapped by sunspot magnetic fields (Parker, 1964; Pnewman, 1967). This energy was once thought significant but now the total amount of trapped energy does not seem to be enough to supply much of the flare energy (Sweet, 1969).

From the above discussion, we see that the sunspot magnetic field is the only probable source of flare energy. Thus, at present, most theories of solar flares are based on this energy as the most important for triggering and development of flares. In these theories, however, we need always consider the configuration of the sunspot magnetic fields and its stability.

4. Nature of Solar Cosmic Rays

In the last section, we have reviewed various topics of solar cosmic ray phenomena, but we have not considered the rigidity (or energy) spectra, the flux and the nuclear composition of solar cosmic rays. These subjects are very important for understanding the acceleration mechanism of solar cosmic rays and related topics.

4.1 Flux, Rigidity and Energy Spectra

As summarized in Table 2-1, the value of the peak flux of Bev solar cosmic rays is highly variable from event to event. This variation is also seen in Mev solar cosmic ray events. In fact, the range of flux of > 10 Mev solar protons which are measured by satellite borne instruments is very wide; from 10^{-2} to 10^{4} cm^{-2} sec^{-1} or more. As shown in Table 2-4, the increase of the > 10 Mev proton flux to greater than 10^{4} cm^{-2} sec^{-1} is comparable to that of the Bev proton flux which is measured by ground based neutron monitors.

In order to study the nature of solar cosmic rays, it is important to obtain detailed information on the distribution of these cosmic rays in energy or rigidity. As the flux of solar cosmic rays varies with time, these spectra also change. This change provides information on the acceleration and propagation mechanism of solar cosmic rays.

In early studies of the differential energy spectra, it was assumed that,

$$\frac{dJ}{dE_K} = C(t)\ E_K^{-\beta(t)} \tag{4-1}$$

where dJ/dE_k is the differential particle intensity per unit energy and E_k is the kinetic energy of particle. The range of β (t) was found to be from 3 to 7, though different from event to event (e.g., Simpson, 1960a,b). At present, it is well known that the energy spectra of relativisitic solar cosmic rays are usually given by the power laws as given by (4-1). However, such power laws are only fitted within some limited energy ranges and even then, with $\beta(t)$ a function of time. Several examples of those spectra are shown in Fig. 4-1. For comparison, the energy spectrum of galactic cosmic rays is indicated also.

It has been shown (Freier and Webber, 1963; Webber, 1964) that the form

$$\frac{dJ}{dR} = \left(\frac{dJ}{dR}\right)_o (t) \cdot \exp\left\{-R/R_o\ (t)\right\} \tag{4-2}$$

is more successful in the expression of particle spectra. Here, R is the particle rigidity, given by

$$R = \frac{Pc}{Ze} , \tag{4-3}$$

where P and Z are the momentum and the atomic numbers of a particle, respectively. The quantity R_o is normally in the rigidity range 40 - 400 MV/C and it generally decreases with time during an event (Freier and Webber, 1963; Webber, 1964). The spectral form given in (4-2) is, therefore, applied during the decaying period of the solar cosmic ray flux in the energy range above ~ 20 Mev. The rigidity spectra of solar cosmic rays are shown in Fig. 4-2 for several events. The expression given by (4-2) is applied to both protons and helium nuclei with similar and sometimes identical values of R_o, but it seems to be only useful for solar cosmic rays in non- and mildly-relativistic energy ranges (Hayakawa et al., 1964; Sakurai, 1965b, 1971c).

4.2 Nuclear abundance

Solar cosmic rays consist of protons, helium and heavier nuclei. Since the first observation by Fichtel and Guss (1961), of nuclei heavier than protons in solar cosmic rays, many data on the nuclear composition of solar cosmic rays have been accumulated by the Minnesota and Goddard groups (e.g., Biswas and Freier, 1961; Ney and Stein, 1962a, b; Biswas, 1962; Freier, 1963; Biswas, Freier and Stein, 1961. 1962; Biswas, Fichtel and Gauss, 1962; Biswas, Fichtel, Guss and Waddington, 1963; Waddington and Freier, 1964; Biswas and

Fichtel, 1964, 1965; Freier and Webber, 1963; Biswas, Fichtel, and Guss, 1966; Durgaprasad, Fichtel, Guss, and Reames, 1968; Bertsch, Fichtel, and Reames, 1972).

As has been shown in the last sub-section the rigidity distribution of non- and mildly-relativisitic solar cosmic rays is expressed by (4-2). This expression can be used for every nuclear specimin (e.g., Freier and Webber, 1963). When the magnitude of R_o is equal for both protons and helium nuclei, the relative abundance ratio of protons to helium nuclei at constant rigidity is given by

$$\frac{\text{number of protons } (p)}{\text{number of helium nuclei } (\alpha)} = \frac{\left(\frac{dJ}{dR}\right)_{o,P}}{\left(\frac{dJ}{dR}\right)_{o,\alpha}} \qquad (4\text{-}4)$$

which will be hereafter notated by P/α. By using the same method for the heavier than helium nuclei at the same rigidity, we also obtain the relative abundance ratio of helium to medium (M) and heavy (H) nuclei. The notations "medium" and "heavy" nuclei technically are used to define the CNO group and the iron group, respectively. The relative abundance of the iron group was recently observed by Bertsch et al. (1969, 1972).

The relative abundance ratio P/α is highly variable (e.g., Freier and Webber, 1963; Sakurai, 1965e, 1971c). The magnitude of P/α is related to the types of solar cosmic ray events such as F,F* and S. For this reason, this ratio P/α cannot be uniquely determined by observing solar protons and helium nuclei. On the other hand, the realtive abundance ratios α/M and α/H are almost constant for most solar cosmic ray events, and, therefore, are very similar to those of the photosphere of the sun (e.g., Biswas and Fichtel, 1964, 1965). Even now, we do not know the helium abundance at the photosphere because no spectroscopic method can be applied to deduce the existence of helium there (e.g., Aller, 1961, 1965; Goldberg et al., 1960; Unsold, 1969). Hence, any information on the helium content in solar cosmic rays is very useful for estimating the helium abundance in the photosphere of the sun. As mentioned above, the ratio P/α, however, is not so useful since as shown in Fig. 4-3, this ratio varies from $\geqslant$ 50 to ~ 1 depending on the time after the onset of the associated flares (Sakurai, 1971c). In this figure, solid and open circles indicate, respectively, those ratios before and after the sudden commencement of geomagnetic storms. This figure suggests that, statistically speaking, the ratio P/α tends to monotonically decrease with time before the beginning of geomagnetic storms (Sakurai, 1965a). As shown by Durgaprasad et al. (1968) and Fichtel (1971), however, this ratio increased with time in the case of the 2 September, 1966 event. It seems, therefore, premature to make a definite statement on the effects of modulation on the ratio P/α.

We have mentioned that the magnitude of P/α is dependent on the type of solar cosmic ray events. As shown by Hakura (1965, 1967a) and later by Sakurai (1971c), F type events usually consist of protons with negligible amounts of other species. Early in F* type events the flux also mainly consists of protons, while later the content of helium nuclei tends to gradually increase with time (see Sakurai, 1971c). Furthermore, the ratio P/α is always ~ 1 in the case of S type events. As mentioned above, it does not seem useful to use this ratio in determining the relative abundance of helium at the sun since this ratio varies so much from one event to another.

As a result of progress in the observing techniques for UV and XUV lines from the sun above the earth's atmosphere, data on the relative element abundances in the solar corona have been recently accumulated (e.g., Pottasch, 1963, 1966, 1967; Jordan, 1966; Warner, 1968; Lambert and Warner, 1968a,b). The analysis of these data shows that the element abundances in the solar corona are very similar to those of the photosphere except for the iron group. The presently available data on these relative element abundances, are summarized in Table 4-1 along with the data for solar cosmic rays and stony meteorites (Pottasch, 1966). In this table, hydrogen is normalized to 1,000,000. Notice that the relative nuclear abundances are very similar among the five observations shown. Recently, Bertsch et al. (1969) obtained the relative abundance ratio of iron nuclei to protons to be $10.1/10^6$. This result is in good agreement with the results obtained from UV, and several forbidden lines of iron from the solar corona.

It seems evident from the above table that the heavy nuclei group abundance in the photosphere is an order of magnitude lower than those of the solar cosmic rays and the solar corona. As remarked by Pottasch (1963,1964) and Jordan (1966), the abundance of the iron group in the solar corona is definitely higher than that of the photosphere. If so, this result is very important to considerations of the distribution of the nuclear species in the solar atmosphere (e.g., Lambert, 1967a,b; Cameron, 1967; Nakada, 1969). Since the relative nuclear abundance of solar cosmic rays is similar to that of the solar corona, it is likely that solar cosmic ray particles are accelerated in the regions high up in the solar atmosphere, i.e., the solar chromosphere (e.g., McCracken, 1969).

As shown in Table 4-1, the abundance of the iron group in the solar corona is a factor 10 higher than that of the photosphere. However, recent experiments have shown that the transition probabilities currently used for the interpretation of the optically visible lines from the iron group, Fe, Ni, and Co, are not correct, but should be reduced by factor of ~10 (Garz and Kock, 1969; Whaling, King and Martinez-Garcia, 1969; Bridges and Wiese, 1970). If these new experimental values are correct, the element abundances in the photosphere are then almost the same as those in the solar corona. Although the transition probabilities are still uncertain, the difference in the relative abundances between the photosphere and the corona does not seem to be as serious as mentioned above.

At present, many new observational results have been accumulated on the overabundance, compared with the photospheric abundance, of the heavy nuclei of solar cosmic rays in the energy range less than ~ 5 Mev/nucleon (e.g., Cowsik and Price, 1971; Price, 1973; Mogro-Capero and Simpson, 1972a,b). These results have not been reconciled with those obtained earlier by Fichtel and co-workers (Bertsch et al., 1972, 1973; Biswas and Fichtel, 1965): that is, in the case of low-energy solar cosmic rays, the difference between the abundances of these cosmic rays and of the solar photosphere becomes greater with increasing charge number Z. In particular, the overabundance is clearly observed for the particles of high charge numbers, such as the iron group (Armstrong and Krimigis, 1971; Armstrong et al., 1972; Price at al., 1971; Mogro-Capero and Simpson, 1972a,b; Price, 1973). It should be noted that this overabundance for these heavy nuclei is only seen in the observational data on low energy solar cosmic rays (say, less than ~ 5 Mev/nucleon.) The nuclear abundance of heavy nuclei in solar cosmic rays of energy higher than 10 Mev/nucleon is very similar to that of the solar photosphere and the corona (Bertsch et al., 1972, 1973; Teegarden et al., 1973).

The difference in the nuclear abundance mentioned above may be explained by taking into account some specific condition in the accelerating regions near flare sites (e.g., Mogro-Capero and Simpson, 1972a,b; Cartwright and Mogro-Capero, 1972). This energy dependent difference may be related to the ionization states of the heavy nuclei in the flare regions and their variation in association with the onset of solar flares. Further investigations are necessary in order to reach some definite conclusion on this subject (e.g., Price, 1973).

As has been discussed above, the relative abundance of helium cannot be determined by means of spectroscopic methods (e.g., Aller, 1961, 1965). If we neglect the slight difference between the relative nuclear abundances of solar cosmic rays and the photosphere, we may use the data for solar cosmic rays to estimate the relative abundance of helium at the photosphere. In doing this, Lambert (1967a, b) assumed that the relative abundances in solar cosmic rays is almost equal to those for the photosphere. Although the ratio P/α is variable from event to event, the ratios such as α/M and α/H are usually constant and independent of any characteristic of solar cosmic ray events. Furthermore, the ratios P/M and P/H are as well known as α/M and α/H in the case of the photosphere. Hence, by using the ratios α/M and α/H in solar cosmic rays and the ratios P/M and P/H in the photosphere, Lambert (1967a,b) estimated the ratio α/P for the photosphere as follows:

$$\frac{\alpha}{P} = 0.063 \pm 0.015$$

Thus, the ratio P/α is 15.9 ± 0.67 (~ 16), but this is much greater than that which is currently used (for example, Goldberg et al., 1960; Aller, 1961). Gaustad (1964) has also proposed a method to estimate the helium abundance in the solar atmosphere by referring to solar cosmic ray data. He finds that the ratio α/P is 0.09. This value is consistent with that which is estimated by Goldberg et al. (1960).

Recently, Durgaprasad et al. (1968) found for this ratio, 0.062 ± 0.008, based on their analysis of the nuclear relative abundance observed on the 16 September, 1966 event. The above value is in good agreement with that estimated by Lambert (1967a,b). However, these results of Lambert (1967a,b) and Durgaprasad et al. (1968) are not in agreement with those which are currently accepted in the cosmic abundances (e.g., Suess and Urey, 1956; Cameron, 1959; Hayakawa, 1970).

The abundance ratio of protons to helium plays an important role in the study of the internal structure and evolution of the sun (e.g., Aller, 1963; Clayton, 1968). In fact, the mass-luminosity relation is entirely determined by the chemical composition of the sun (e.g., Schwartzschild, 1958; Clayton, 1968). The sun is representative of a large class of stars which belong to the main sequence. Thus, it is important to estimate the energy production rate necessary for the sun to remain in the main sequence. In doing this, it is necessary to first assume the ratio of protons to helium (P/α) in the solar interior (e.g., Sears, 1964; Demarque and Percy, 1964; Weyman and Sears, 1965; Morton, 1967). The ratios, necessary to explain the observed mass-luminosity relation of the sun, are adopted as summarized in Table 4-2. From this table, the values of P/α are clearly greater than those which are estimated by Lambert (1967b). Values similar to those given in the table ($\alpha/P \sim 0.086$) were also used by Schwarzschild (1958) in the study of the internal structure of the sun. Hence the result obtained by Lambert (1967a,b) suggests that the chemical composition in the solar interior is different from that in the solar atmosphere. It is interesting to note that the ratio P/α in cosmic space is estimated as 6.25 which is also much smaller than that obtained by Lambert (1967a).

It is known that the observed flux of solar neutrinos at the earth is also useful for estimating the ratio P/α. By using the neutrino data obtained by Davis, Harmer and Hoffman (1968), Bahcall, Bahcall and Shaviv (1968) estimated the upper limit of the ratio α/P. Later, Iben (1968) critically reviewed the result of Bahcall et al. (1968) and then estimated this upper limit to be 0.049 - 0.064. The ratio α/P estimated by Lambert (1967a,b) from solar cosmic ray data is, therefore, almost equal to the largest value of these upper limits. If we adopt the ratio α/P of 0.064, we are then led to the conclusion that the chemical abundance in the solar interior is the same as those in the solar cosmic rays and in the solar atmosphere. But, such a low helium abundance is in conflict with the helium abundance as deduced from the theory of the solar interior which explains the mass-luminosity relation of the sun.

At present, the ratio P/α for high temperature stars is estimated to be ~6.2, despite the fact that their evolution rate is really much higher than that of the sun (e.g., Hayakawa, 1970). We can see that the exact determination of the chemical composition of the sun is very important for our understanding of the evolution of main sequence stars and their relation to the genesis of helium in the universe (e.g., Hoyle and Taylor, 1964; Taylor, 1967; Peebles, 1971).

The abundances of the light nuclei (L), Li, Be, and B are also important for the understanding of the behavior of solar cosmic rays in the solar atmosphere. Practically speaking, these nuclei have never been detected in the photosphere of the sun, but the upper limit of the ratio of the light nuclei to protons, defined as L/P is estimated as $< 10^{-2}$ in the unit of 10^6 protons (e.g., Goldberg et al., 1960; Aller, 1963). Based on the observations of solar cosmic rays, the relative abundance of $_4Be$ to $_5B$ in solar cosmic rays has recently been estimated to be < 2 in the same unit mentioned above (Biswas and Fichtel, 1965; Fichtel and McDonald, 1967). This value is $\sim 10^3$ smaller than the relative abundance of the light nuclei in galactic cosmic rays (e.g., Aller, 1961; Aizu, Ito and Koshiba, 1964). This fact gives us some information concerning the fragmentation processes of heavy nuclei in solar flares.

As discussed above, the determination of the relative nuclear abundances in the interior and the atmosphere of the sun is very important for the understanding of the acceleration of solar cosmic rays and of the evolution of the sun. Furthermore, the study of solar cosmic rays is useful in our attempts to understand the genesis of helium nuclei during the evolution of the universe (e.g., Peebles, 1971).

4.3 Neutrons and electrons in solar cosmic rays

After acceleration in solar flares, high-energy protons and heavier nuclei seem to interact with ambient atoms and ions in the solar atmosphere. Due to this interaction, neutrons and high-energy electrons (and positrons) apparently are produced in the flare regions.

Neutrons:

The production of neutrons in solar flares was first discussed by Biermann, Haxel and Schlüter (1951). Since Mev or Bev protons are sometimes produced in solar flares of great importance, neutrons can be produced as a result of nuclear interaction with atoms and ions in the vicinity of the flare.

As summarized in Table 4-3, there are several possible nuclear interactions that could produce neutrons in solar flares. The expression H^1 $(P, n\,\pi^+)$ H^1 in this table, for example, means the nuclear reaction

$$P + P \longrightarrow P + n + \pi^+.$$

Namely, the above reaction produces a neutron as a result of the proton-proton interaction. In order for this reaction to occur, the kinetic energy of the interacting protons must be higher than 292.3 Mev.

In order to calculate the rate of neutron production due to the above process, we must know the production cross section as well as the numbers and energy spectra of the ambient and accelerated high-energy protons. Hence, the result obtained inevitably includes some uncertainty for the total number of neutrons produced. There exists a variety of ambiguities, including, in particular, the ejection rate of the accelerated protons into the dense atmosphere and the direction of the neutron ejection. In spite of these difficulties, the total numbers of neutrons produced in solar flares and the quiet sun have been recently estimated and the possibility of detection of these neutrons at earth has been considered (e.g., Chupp, 1964, 1971; Lingenfelter, Flamm, Confield, and Kellman, 1965a,b; Ito, Okazoe and Yoshimori, 1968; Lingenfelter, 1969; Forrest and Chupp, 1969).

Bame and Asbridge (1966) tried to satellite detection of the neutron flux from the quiet sun, but did not obtain positive evidence for neutron emission. On the other hand, Daniel et al. (1970) and Apparao et al. (1966) showed that neutrons were possibly emitted from solar flares. These two results suggest that neutrons are sometimes emitted from the sun in association with solar flares; however, there are no positive observational data from either the quiet or the disturbed sun (e.g., Lingenfelter and Ramaty, 1967; Chupp, 1971).

Since these neutrons decay to protons and electrons in about 14 minutes, most neutrons decay during their flight between the sun and the earth (e.g., Lingenfelter et al., 1965a,b). The protons thus produced must then propagate under the guidance of the interplanetary magnetic field (Roelof, 1966a).

<u>Electrons</u>:

We shall here consider the possible positron production processes because we have already discussed the electron production in solar flares (see 2-5).

High-energy positrons are produced by the process of positive pion decay as follows:

$$\pi^+ \longrightarrow \mu^+ + \nu$$

and

$$\mu^+ \longrightarrow e^+ + \nu + \nu.$$

Since these positive pions are produced by such processes as PP and Pα interactions, it seems important to study these processes in solar flares. The time for the above decay ($\pi^+ \rightarrow \mu^+ \rightarrow e^+$) is about 10^{-6} seconds. Hence such positrons are immediately produced as soon as positive pions are generated. As suggested by Lingenfelter and Ramaty (1967), these positrons may play an important part in the emission of microwave impulsive radio bursts which usually begin with the start of the explosive phase of solar flares.

If many Mev protons are trapped in sunspot magnetic fields before the onset of solar flares as proposed by Elliot (1964) and Reid (1966), the production rate of the positive pions could be quite large if raised by a factor of 10 or more. However, this possibility does not seem plausible from the viewpoint of energetics of solar flares.

These positrons, through their annihilation, also seem to be important as a source of gamma ray emissions (e.g., Dolan and Fazio, 1965; Lingenfelter and Ramaty, 1967; Chupp, 1971). Because this process is closely related to nuclear processes in the solar atmosphere, the subject of gamma ray emissions will be considered in detail in (4-5).

4.4 Isotope abundance (deuterons and He^3)

The nuclear reaction processes related to the production of deuterons and helium isotope He^3 are summarized in Table 4-4. The most important of the reactions are those between solar cosmic ray protons and the ambient protons, heliums (He^4) and medium nuclei (CNO group) because these latter particles are relatively abundant in the solar atmosphere (e.g., Lingenfelter and Ramaty, 1967). We have already discussed the processes for neutron production in (4-3). The other important nuclear interaction is related to the production of tritons in solar flares. Part of these nuclear products would be later ejected from the sun into interplanetary space and then be detected near the earth.

The cross sections for those nuclear reactions shown in Table 4-4 have been calculated by Lingenfelter and Ramaty (1967). The measurement of Goebel at al. (1964) indicates that the yield ratio He^3/H^3 is of the order of two when produced from protons of energy from one to several 100 Mev. In reality, helium and the medium nuclei (CNO, Ne) in the solar atmosphere seem to play important roles in the production of deuterons, tritons and He^3.

The production of deuterons and He^3 is proportional to the path length of solar cosmic rays in the photosphere. This length can be estimated by using the cross sections for nuclear reactions, the ambient mass density and the observed results on these isotopes. The measurement of deuterons in solar cosmic rays was made for the two events on 12 November

1960 and 18 July, 1961 (Biswas, Freier and Stein, 1962; Freier and Waddington, 1964; Waddington and Freier, 1964). If this path length is about equal to the diameter of the Hα brightening areas, the mean density of ambient plasma in the region traversed by solar cosmic rays is estimated as 3×10^{10} to 3×10^{11} electrons/cm^3. These values are smaller than the plasma density in the Hα flare regions by about two orders of magnitude, but are consistent with the values estimated by Jefferies and Orrall (1961a,b) and Sakurai (1971b).

The observed ratios of deuterons to protons are $\lesssim 2 \times 10^{-3}$ in the energy range > 50 Mev/nucleon on the 12 November, 1960 event (Biswas et al., 1962) and $\sim 5 \times 10^{-3}$ for protons of 15-75 Mev and for deuterons of 20-100 Mev on both the 16 March, 1964 and 5 February, 1965 events (McDonald et al., 1965).

The production of He^3 has been calculated by Lingenfelter and Ramaty (1967). However, relatively little is known for the relative abundance of He^3 (e.g., Biswas and Fichtel, 1965). The measurement of He^3 for solar cosmic rays was done only by Schaeffer and Zähringer (1962) and Biswas et al.

(1962), but we do not know as yet the relative abundance of He^3 in solar cosmic rays in detail (e.g., Comstock et al., 1972; Dietrich, 1973; Hsieh et al., 1962; Anglin et al., 1973).

4.5 Positrons and gamma-ray emissions

Positrons are produced as a result of the decay of the π^+ mesons through $\pi^+ \rightarrow \mu^+ \rightarrow e^+$. Since the energies of the π^+ mesons produced from PP and Pα reactions is generally higher than ~ 285 and ~180 Mev, respectively, the energy of positrons thus produced is ultrarelativisitic. The important pion producing processes are summarized in Table 4-5 (Lingenfelter and Ramaty, 1967). The interactions of solar cosmic ray protons with the ambient protons and helium (He^4) are most effective for the production of pions (π^+, π^- and π^o). In this table, a and b are artificial multiple numbers because π^+, π^- and π^o mesons are multiply produced by the reactions described in the table.

The pions seem to be produced in and near the flare regions. Since the life time for the decay $\pi^+ \rightarrow \mu^+ \rightarrow e^+$ is ~ 10^{-6} seconds, positrons are made simultaneously. These positrons, in the sunspot magnetic fields in and near the flare regions, would then be a source of microwave emissions in the explosive phase through the synchrotron emission process. This idea was once proposed by Lingenfelter and Ramaty (1967) to explain microwave impulsive radio bursts.

Some processes like the PP and Pα reactions can also produce neutral π mesons (π^{o}) (See Table 4-3). They instantaneously decay with a half-life less than 10^{-15} seconds into two gamma rays with an energy spectrum peaked at 67.6 Mev. Moreover, the de-excitation of excited nuclei, the capture of neutrons by hydrogen and the annihilation of positrons produced by the decay of charged pions can produce gamma rays, but the most important process seems to be the decay of π^{o} mesons (Lingenfelter and Ramaty, 1967). Actually, the relative importance among these processes is dependent on the rigidity spectrum of the solar cosmic rays. Since the energy range of solar cosmic rays is usually lower than ~1,000 Mev except for the Bev particle events as described in Table 2-1, the de-excitation of the excited nuclei C^{12}, N^{14},O^{16}, and Ne^{20} would be important in addition to both the neutron capture and the annihilation processes as mentioned above. The energies of gamma rays from the capture and annihilation processes are 2.23 and 0.51 Mev, respectively.

The energy range of gamma rays produced from the above de-excitation processes has been calculated by Lingenfelter and Ramaty (1967).

Recently, these characteristic gamma ray emissions have been observed by Chupp et al. (1973). They were associated with the solar proton flares on 4 and 7 August, 1972. These observations shows that high energy neutrons are produced through various nuclear interactions as summarized in Table 4-3. Theoretical interpretations of these observations have been given by Chupp et al. (1973b) and Ramaty and Lingenfelter (1973).

5. Acceleration mechanism of solar cosmic rays

Solar cosmic rays are produced from solar flares which have the characteristics discussed in section 3. Although we do not yet fully understand the mechanism of these flares, we can investigate the mechanism of solar cosmic ray acceleration to explain the observed nature of solar cosmic rays. This mechanism seems to be related to the space-time variation of sunspot magnetic fields. In this section, we shall first consider the general theory of particle acceleration and then examine its relation to the observations.

5.1 Reviews of acceleration theories

Acceleration of charged particles, in general, occurs through their interaction with electric fields of various origins. The theory of the acceleration was first developed in order to explain the origin of galactic cosmic rays. In 1933, Swann (1933) studied the acceleration of charged particles on the basis of the interaction of charged particles with time-varying magnetic fields. This mechanism is often called "betatrón" acceleration. In this mechanism, the electric field induced according to Faraday's law energized the charged particles (e.g., Alfvén and Fälthammer, 1963; Northrop, 1963a).

In the early 1940's, the behavior of magnetic field lines in ionized media was investigated by Alfvén (1942). He showed that in the astrophysical setting where the electrical conductivity is very high, the magnetic field lines move with the ambient ionized medium, i.e., it is said that the field lines are "frozen" to the plasma. In the coordinate system moving with the plasma, because of the high conductivity, we do not observe an electric field (e.g., Cowling, 1953), but in any other coordinate system we obtain an electric field given by

$$\underline{E} = - \frac{1}{c} \underline{u} \times \underline{B}, \tag{5-1}$$

where $\underline{E}$,$\underline{B}$ and $\underline{u}$ are respectively the electric and magnetic fields and the velocity of the medium as a whole. This electric field can sometimes accelerate charged particles.

The equation of motion of a charged particle in electromagnetic fields is given as

$$\frac{d\underline{P}}{dt} = Ze \ (\underline{E} + \frac{1}{c} \underline{v} \times \underline{B}), \tag{5-2}$$

where $\underline{P}$, Ze, $\underline{v}$ and t are respectively the momentum, the electric charges, and the velocity of the particle and the time. By substituting (5-1) into (5-2), we obtain

$$\frac{dW}{dt} = \frac{Ze}{c} \ \underline{u} \cdot (\underline{v} \times \underline{B}), \tag{5-3}$$

where W is the total energy of the particle. This equation indicates that the charged particle can be accelerated by the electric field induced in the medium (Parker, 1958). The acceleration of charged particles by this electric field was first considered by Fermi (1949, 1954). As shown by Fermi (1949), there exists two distinct types of "Fermi acceleration." They are schematically described in Fig. 5-1: the Fermi I acceleration is associated with the reflection of particles from moving magnetic scattering centers (Fig 5-1 (a)); whereas the Fermi II mechanism is due to the motion of particles along magnetic flux tubes between moving mirrors (Fig. 5-1(b)).

From the statistical point of view, the efficiency of the two Fermi acceleration processes is the same (Fermi, 1949; Sakurai, 1965a,b). The statistical rate of energy gain is given by,

$$\frac{d}{dt} \left(\frac{W}{W_o}\right) = 2 \gamma \frac{u^2}{c^2} \cdot \frac{1}{T}, \tag{5-4}$$

where W_o, T and γ are respectively the rest energy of the particle, the effective mean scattering time and the Lorentz factor. Since this acceleration is proportional to $(U/c)^2$, it is not usually very effective. An important feature of this acceleration rate is that it is proportional to the particle energy since $W = W_o \gamma$ (Parker, 1957, 1958; Ginzburg and Syrovatskii, 1964).

In contrast to the stochastic Fermi mechanism, the betatron mechanism, given by

$$\frac{\partial W}{\partial t} = \frac{M_r}{\gamma} \frac{\partial B}{\partial t}, \tag{5-5}$$

is a reversable process since the energy change is dependent on the sign of $\partial B/\partial t$. Here the magnetic moment, M_r, is given by

$$M_r = \frac{P^2 \sin^2 \alpha}{2m \quad B} \tag{5-6}$$

where m and α are the rest mass and the pitch angle of the particle (e.g., Northrop, 1963 a; Hayakawa et al., 1964). The manner in which these two acceleration mechanisms change the energy of the charged particles is very different. Furthermore, it has been shown that, in a plasma in which the magnetic field lines are very effectively frozen in, the betatron mechanism does not function.

The role of electric fields other than those which are induced by space and time variations of magnetic fields was first taken up by Schlüter (1952). A static electric field would, of course, accelerate particles; but it is very difficult to maintain such fields in a high conductivity plasma. A quasi-static electric field would typically, have to last for a few minutes or more to be very effective for the acceleration of charged particles. Sometimes, the effectiveness of oscillating electric fields associated with plasma waves is considered (e.g., Bohm and Pines, 1949; Spitzer, 1962).

In solar flares, electric fields can be produced by space and time variation of sunspot magnetic fields in the flare regions. These fields may be effective for the acceleration of solar cosmic rays. Therefore, the acceleration mechanisms of solar cosmic rays are being investigated by taking observed characteristics of solar flares and flare regions into account (e.g., Parker, 1957; Syrovatskii, 1961; Severny and Shabanskii, 1961; de Jager, 1962; Schatzman, 1963, 1967 a,b; Sakurai, 1965 b, 1971 b). In order to understand particle acceleration in solar flares we need to consider the relationships between the theories of particle acceleration and the observed properties of solar cosmic rays. The observational results considered in the last section, on the energy and rigidity spectra and the relative nuclear abundances of solar cosmic rays have been shown to be useful in this respect (e.g., Waddington and Freier, 1964; Hayakawa et al., 1964; Biswas and Fichtel, 1964; Sakurai, 1965 b,c; Wentzel, 1965).

5.2 Principal mechanism of particle acceleration

The general theory of particle acceleration in varying magnetic fields has been investigated by many authors since the first important work of Fermi (1949) (see Northrop, 1963 a,b; Hayakawa et al., 1964; Hayakawa and Obayashi, 1965; Sakurai, 1974). In developing the theory, the guiding center approximation is adopted. Because the general theory has been well established, we will here refer to the results which bear directly on the principal mechansim of solar cosmic ray particle acceleration.

Let us assume that the variation of the ambient magnetic field is expressed as

$$\frac{dB}{dt} = \frac{\partial B}{\partial t} + (u \cdot \nabla)B. \tag{5-7}$$

When the field variation is expressed as above, the equation for energy gain is given by

$$\frac{dW}{dt} = \frac{C^2 P^2 \sin^2\alpha}{2\ W\ B} \left(\frac{\partial B}{\partial t} + (u \cdot \nabla)B\right). \tag{5-8}$$

As indicated in the last section, the first term on the right side of eq. (5-8) corresponds to betatron acceleration. The second term is called Fermi acceleration, since the energy gain is due to the motion of the field lines. Hence, in the guiding center approximation, the acceleration consists of both the betatron and Fermi mechanism (Northrop, 1963 a,b; Sakurai, 1965 b,c). By substituting eq. (5-7) into (5-8), both acceleration mechanisms could be considered together through the expression

$$\frac{dW}{dt} = \frac{C^2P^2}{2W} \sin^2\alpha \; (\frac{\dot{B}}{B}),$$

where "." indicates the time derivative of B. However, since these two mechanisms are quite different from each other we will consider them separately in the following.

The average energy gains for the two acceleration mechanisms averaged over time long enough for several mirrorings to take place, are given by

$$<(W)_B> = \frac{C^2P^2}{W} < \frac{\sin^2\alpha}{2B} \frac{\partial B}{\partial t}> \tag{5-9}$$

and

$$<(W)_F> = \frac{W}{\Delta t} \; 2 \; (\frac{u}{c})^2, \tag{5-10}$$

(Hayakawa et al., 1964; Sakurai, 1965 b,c). Thus, the total average rate of energy change is,

$$<\frac{dW}{dt}> = \bar{f} \; W + b \; v \; p, \tag{5-11}$$

where

$$\bar{f} = 2 \; (\frac{u}{c})^2 \; \frac{1}{\Delta t}$$

$$b = <\frac{\sin^2\alpha}{2B} \frac{\partial B}{\partial t}>.$$

The above expression is applicable only after enough time has elapsed to allow the particle to mirror many times.

Equation (5-11) can be rewritten in several useful forms. Using the relation $\partial W/\partial t = v\,(\partial P/\partial t)$, equation (5-11) can be rewritten as

$$\left\langle\frac{dP}{dt}\right\rangle = \bar{f}_P \sqrt{P^2C^2 + m^2c^4} + bP \qquad (5\text{-}12)$$

where $\bar{f}_P = f/v$. By taking the definition of rigidity into account, we can rewrite (4-51) as an equation for the rigidity gain:

$$\left\langle\frac{dR}{dt}\right\rangle = a\sqrt{R^2 + \frac{m^2c^4}{(Ze)^2}} + bR, \qquad (5\text{-}13)$$

where $R = Pc/Ze$ and $a = \bar{f}\,(c/v)$. This equation will be referred to in the following discussion. As is evident from this equation, the Fermi acceleration is proportional to $[R^2 + (\frac{mc^2}{Ze})^2]^{1/2}$ while the betatron acceleration is only dependent on the rigidity R. If the rigidity is low, the Fermi acceleration becomes independent of R. When $R \gg mc^2/Ze$, the Fermi acceleration also becomes proportional to R, but the rate of acceleration is not the same as that for the betatron acceleration since, in general, $a \neq b$. The rate of rigidity gain for the two acceleration mechanisms is shown as a function of R in Fig. 5-2.

Equations (5-11) and (5-13) can be used to estimate the energy or rigidity spectra of accelerated particles, respectively, by solving the continuity equation in the energy-time or the rigidity-time space (Roederer, 1964 a,b: Hayakawa et al., 1964; Wentzel, 1965; Sakurai, 1966a). The continuity equation in rigidity-time space is given by

$$\frac{\partial N\,(R,t)}{\partial t} = - \frac{\partial}{\partial R}\,(N(R,t)\,\langle \frac{dR}{dt} \rangle) - \frac{N(R,t)}{t_e} + q(R,t), \qquad (5\text{-}14)$$

where $N(R,t)$, t_c and $q\,(R,t)$ are respectively, the number of particles in the rigidity interval $(R,R+dR)$, the mean confinement time in the acceleration region and the injection rate of particles. For the injection rate, there should be an upper cut off: $q\,(R,t) = 0$ for $R > R_i$, where the maximum injection rigidity R_i probably lies somewhere in the upper tail of the thermal spectrum. In the above equation, the term $\langle dR/dt \rangle$ is given by (5-13). If we change from R to W, (5-14) becomes the continuity equation in energy-time space.

We shall separately consider the betatron and Fermi mechanisms in solving the continuity equation (5-14). First, the betatron mechanism will be considered. In this case, the rigidity gain is given by

$$\left\langle \frac{dR}{dt} \right\rangle = bR, \tag{5-15}$$

or,

$$R = R_o \exp (bt),$$

where R_o is the initial rigidity. The rigidity increases exponentially with time, for as long as the guiding center approximation continues to hold.

Let us assume now that the confinement time τ_c is short enough so that a steady state is reached shortly after the beginning of acceleration. Then, $\partial N/\partial t = 0$, and, by integrating (5-14), we obtain

$$N(R) = \tau_a R^{-(1 + \frac{\tau_a}{\tau_c})} \int_o^R q(R')R'^{\frac{\tau_a}{\tau_c}} dR' \tag{5-16}$$

where

$$\tau_a = \frac{1}{B}$$

since we are only interested in high rigidity particles $(R > R_i)$, it follows from this equation that

$$N(R) = kR^{-\gamma} \tag{5-17}$$

with

$$\gamma = 1 + \frac{\tau_a}{\tau_c} \text{ and } K = \tau_a \int_o^{R_i} q(R')R'^{\frac{\tau_a}{\tau_c}} \, dR'$$

The above result indicates that the betatron acceleration gives a power law spectrum in particle rigidity with spectral index, γ, determined by the ratio of the acceleration to the confinement times. We shall next consider the Fermi acceleration alone. In this case

$$<\frac{dR}{dt}> = a \sqrt{R^2 + (\frac{mc^2}{Ze})^2} \tag{5-18}$$

with considerably different behavior in the non-relativisitic and relativistic energy ranges.

In the non-relativistic range, the rigidity gain after time, t is given by

$$R = R_i + a \left(\frac{mc^2}{Ze}\right) t,$$

where R_i is the initial rigidity. In the steady state, the rigidity spectrum is

$$N(R) = \frac{1}{a} \frac{Ze}{mc^2} \exp\left(- \frac{Ze}{amc^2} \frac{R}{\tau_c}\right) \int_o^R q(R') \exp\left(\frac{Ze}{amc^2} \frac{R'}{\tau_c}\right) dR' \tag{5-19}$$

For $R > R_i$, this result is

$$N(R) = k \exp\left[-\frac{R}{R_o}\right] \tag{5-20}$$

with

$$R_o = a\,\frac{mc^2}{Ze}\,\tau_c + R_i \quad \text{and} \quad k = \frac{Ze}{amc^2}\,q(R')\int_o^{R_i} \exp\left(-\frac{Ze}{amc^2}\,\frac{R'}{\tau_c}\right) dR'$$

In the non-relativistic range, the Fermi acceleration gives an exponential rigidity spectrum.

Since the Fermi acceleration is almost proportional to R in the relativistic range, a power law spectrum is obtained just as in the case of betatron acceleration. The quantity τ_a for the betatron acceleration must be replaced by $\tau_F = 1/a$ In this case, the spectral index is given by

$$\gamma = 1 + \frac{\tau_F}{\tau_c} \tag{5-21}$$

As a result, the rigidity spectrum changes from exponential to power law as the particle rigidity increases from non-relativistic to relativistic (e.g., Hayakawa et al., 1964; Sakurai, 1965b,c, 1966 b).

5.3 Interpretation of the observation

The rigidity spectrum of solar cosmic rays is exponential for energy $\lesssim$ 500 Mev. This fact can be explained by taking into account the results discussed above; namely, in order to explain this exponential spectrum, we can assume that solar cosmic rays are mainly accelerated by the Fermi mechanism (e.g., Hayakawa et al., 1964; Wentzel, 1965; Sakurai 1965 b).

Except for the proton component, furthermore the relative nuclear abundances of solar cosmic rays are very similar to those which are observed in the photosphere of the sun. This result can also be explained by the Fermi acceleration as follows: the acceleration rate for this process is proportional to mc^2/Ze in the non-relativistic range. The particle mass is $m = Am_p$, where A is the mass number of the nucleus under study. Thus the acceleration rate is

$$\frac{dR}{dt} = a \frac{A}{Z} \left(\frac{m_p c^2}{e}\right) \tag{5-22}$$

It is clear that the acceleration rate is proportional to the ratio of the mass to the atomic number (A/Z). The ratio A/Z is equal, or nearly equal, to 2 for helium and other heavier nuclei, while this ratio is 1 for protons. Thus, the acceleration rate is almost the same for all nuclei except the proton. This difference may explain why the proton abundance in solar cosmic rays is so variable and is further so different from that of the photosphere. In summary, we can say that, if solar cosmic rays are mainly accelerated by the Fermi mechanism in solar flares, the observed rigidity spectrum and nuclear composition are consistently explained (e.g., Sakurai, 1971 c).

The observed relativistic solar cosmic rays are well described by power law rigidity spectra. This fact can also be explained by taking only the Fermi mechanism into consideration. Thus, the acceleration of solar cosmic rays is most likely due to the Fermi mechanism (e.g., Hayakawa et al., 1964; Wentzel, 1965; Sakurai, 1965 a,e).

As shown in Fig. 4-3, the ratio P/α varies with time after the onset of an associated flare. This observational fact makes it difficult to consider the acceleration mechanism of the proton component independent of the propagation mecahnism of these two components in the interplanetary space. However, the exponential spectrum of solar cosmic ray protons in the non-relativistic range suggests that these protons are accelerated by the Fermi mechanism, too.

5.4 Energy loss processes during acceleration

As a result of ionization, collision, bremsstrahlung, Compton scattering and radiation interactions with ambient atoms and ions in the acceleration region, particles also lose energy during the acceleration process (Hayakawa and Kitao, 1956; Parker, 1957; Ginzburg and Syrovatskii, 1964; Holt and Cline, 1968). However, these energy losses can be shown to be negligible during the acceleration of the proton and other heavier nuclear components (e.g., Hayakawa et al., 1956).

For electrons, however, the energy loss processes must be considered. The energy loss rates for several processes are shown in Fig. 5-4 as a function of the electron kinetic energy (Sakurai, 1967 c, 1971 b). This figure shows that, in the relativisitic energy range, the gyro-synchrotron loss process becomes dominant, while, in the non-relativistic range the ionization loss is most effective. In general, the bremsstrahlung and the Compton scattering losses are not significant. Thus, in considerations of the electron acceleration in solar flares, it is enough to consider only the ionization and gyro-synchrotron losses.

As is evident from Fig. 5-3, the injection energy of the accelerating electrons is highly dependent on the ambient electron density in the acceleration region. If the electron acceleration occurs in flare regions with plasma density $10^{12}-10^{13}$ cm^{-3} and with sunspot magnetic fields $10^{2}-10^{3}$ gauss, the range of injection energy is estimated to be several x 10 to 10^{4} Kev or more. Since it is unlikely that electrons of 100 Kev or more are ambient in the flare regions, the injection energy of electrons is probably around 10 Kev. However, if we assume this injection energy, the acceleration rate, from Fig. 5-3, must be higher than 1.5×10^{8} ev sec^{-1} (Sakurai, 1971 b). If this is the case, Mev electrons would be produced almost simultaneously with the start of a solar flare and then optically visible continuum emission would be produced by the synchrotron mechanism. However, we have only rarely observed such continuum emission (see 3-1). This result suggests that the electron acceleration regions are not the same as the flare regions, but are identified with regions of electron density $< 10^{12}$ cm^{-3} and the magnetic intensity $\leqslant 10^{2}$ gauss.

Sakurai (1971 b) has estimated that the plausible plasma density and magnetic field intensity are, respectively,

$$n_p \cong n_e \simeq 10^8 - 10^{10} \text{ cm}^{-3}$$

and

$$B \simeq 10 - 10^2 \text{ gauss}$$

These values are much smaller than those which are given in the Hα flare region (e.g., Svestka, 1966; de Jager, 1969). This result means that the electron acceleration regions are located higher than the Hα flare regions and the injection energy of electrons is reduced to 1-10 Kev. These electrons seem to be present in the acceleration regions before the onset of solar flares in the tail of the thermal Maxwell distribution (e.g., Takakura, 1961, 1962; Sakurai, 1967 c, 1971 b).

5.5 Chronology and related efficiencies of the acceleration process

For the chronology of the acceleration of high energy particles in solar flares, three alternative ideas have been proposed (e.g., Sweet, 1969). They are as follows:

1. The fast nuclei are present in the solar atmosphere before the flares (Elliot, 1964, 1969; Reid, 1966).

In this case, we do not need to consider the acceleration process during the flare. We only need consider the process for the release of these nuclei from the flare regions.

2. All the high energy particles are accelerated simultaneously during the explosive phase of flares (Sakurai, 1965 d, 1971 d; Svestka, 1970). and,

3. The electrons of 1 Kev - 1 Mev energy are accelerated during the explosive phase, but the electrons of energy $\gtrsim$ 1 Mev and the high energy nuclei of 1 Mev - 30 Bev are independently produced during the main phase by the Fermi mechanism (de Jager, 1969).

Among them, ideas 2 and 3 assume acceleration processes in solar flares, but it should be remarked that their essential features are very different. In order to determine which idea is more plausible, we need to examine some characteristics associated with solar proton flares.

The emission of intense Hα line radiation, hard X-rays and microwave and type IV μ radio bursts, in general, starts with the beginning of the explosive phase of solar flares. This observation suggests, at least, that the acceleration of solar cosmic rays starts with or in this phase, although it does not show whether the acceleration finishes during this phase or continues into the main phase. Recently, Sakurai (1971 d) showed that the acceleration of solar cosmic rays is almost completed during the explosive phase of solar flares. As shown in Fig. 3-3, the emission of both type IV μ radio and X-ray bursts started before the beginning of the main phase on 7 July, 1966. Furthermore, the times at which type IV emissions for different frequencies reached peak flux are almost simultaneous and independent of the outward motion of the type II radio source. This result suggests that the development of the main phase is not related to the behavior of the high energy electrons responsible for type IV radio bursts (IV μ, IV_{dm} and IV_m). Thus, we suggest that solar cosmic rays and high energy electrons are mainly accelerated during the explosive phase of solar flares (Svestka, 1970 b; Sakurai, 1971 d). However, we cannot give up the possibility of the secondary acceleration because metric type IV continuum emissions, defined as IV_mB, are often observed from several hours to several days after the end of solar flares.

The efficiency of solar cosmic ray acceleration is related to the length of the rise time of the Hα brightening. In fact, this efficiency becomes higher as the rise time becomes shorter (Sakurai, 1970 b). The length of this rise time seems to be closely related to the rapidness of the development of the explosive phase of flares. Furthermore, this efficiency is decisively dependent on the gradient of sunspot magnetic fields near the neutral regions (Severny, 1964 b, 1965; Sakurai, 1972 a). The relation between the importance of solar proton events and the gradient as mentioned above has been already shown in Fig. 3-9. These results also suggest that the initial stage of the flare development is very important for the production of high energy particles.

7. Concluding Remarks

In this paper, we have reviewed the present state of solar cosmic ray research. As we have shown, there exist many problems to be investigated extensively in the near future: for example, we do not yet understand the solar flare mechanism and its relation to particle acceleration. Even though the understanding of this mechanism is very important for solar cosmic ray physics, we have no promising clues at present. In concluding our paper, we summarize the problems to be studied:

(1) Solar flare mechanism and its relation to particle acceleration,

(2) Acceleration mechanism of high-energy particles and its relation to magnetic field annihilation,

(3) Acceleration phase (one or two?),

(4) Behavior of high-energy particles in the solar envelope,

(5) Physical state of the accelerating region (in relation to the isotopic production and the charge state of accelerating particles),

(6) Relation among shock waves, magnetic bottles and particle acceleration,

(7) The nuclear abundance in the sun,

(8) Propagation mechanism of high-energy particles as deduced from theoretical treatment.

These problems must be investigated in order to understand systematically the generation and propagation of high-energy particles in solar flares. Most important would be the solar flare mechanism, because this mechanism is the source of the high-energy particles at the sun.

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Contents of Tables

Table 2-1 Solar cosmic ray events of Bev energy

	Date	Solar Flare Onset	Solar Flare Position	Imp.	Cosmic Ray Increase Onset	Cosmic Ray Increase Max. Increase(%)
1942	Feb. 28	1200	07° N 04° E	3+	(1230)	(15)
	March 7	(0442)	(07° N 90° W)	–	(0512)	(40)
1946	July 25	1620	29° N 15° E	3+	1700	(22)
1949	Nov. 19	1030	02° S 70° W	3+	1045	(42)
1956	Feb. 23	0332	23° N 80° W	3+	0343	2000
	Aug. 31	1228	15° N 15° E	3	–	2
1957	Sept. 2	1313	35° S 36° W	3+	–	2
	Sept. 21	1330	10° N 08° W	3	–	2
1959	July 16	2118	11° N 30° W	3+	(0018)	10
1960	May 4	1015	12° N 90° W	3	1029	280
	Sept. 3	0037	20° N 87° E	3	0337	4
	Nov. 12	1325	26° N 04° W	3+	1345	120
	Nov. 15	0217	26° N 33° W	3+	0238	80
	Nov. 20	2022	28° N 109° W	3	2054	5
1961	July 18	0938	08° S 60° W	3+	1000	15
	July 20	1552	06° S 90° W	3	1610	4
1966	July 7	0023	34° N 48° W	2B	0113	2
1967	Jan. 28	–	–		–	32
1968	Nov. 18	1017	21° N 87° W	1B	1042	–

Table 2-2 Classification of Mev solar cosmic ray events as observed as polar cap absorptions (Leinbach)

Sudden-onset[a]	Pre-Sc Max.	Fig. 2-7(a)	F[b]
	Sc Max.	Fig. 2-7(b)	F*
Graual-onset	Sc Max. (extensively)	Fig. 2-7(c)	S
	Complex	Fig. 2-7(d)	Complex

(a) Sinno's classification

(b) Obayashi (1964) and Sakurai (1965c)

Table 2-3 Classification of solar proton events as observed as polar cap absorption (Obayashi et al., 1967)

Importance	1-	1	2	3	3+
f-min increase at Resolute Bay	Very small f-min < 3 MHz	f-min > 3 MHz t < 6 hrs*	f-min > 5 MHz or Blackout t > 6 hrs	Blackout t > 24 hrs	Blackout t > 48 hrs

*Starting time adopted is that of an increase of f-min.

t is the duration of an f-min increase (3 MHz)

Table 2-4 Classification of cosmic ray events

Index Number	Sea Level Neutron Monitor Increase	Daily Polar Riometer Absorption	E > 10 Mev Satellite Measured Proton Intensity particles/cm^2 sec
-3			From 10^{-3} to 10^{-2}
-2			10^{-2} 10^{-1}
-1			10^{-1} 1
0	No measurable increase	No measurable increase	1 10
1	Less than 3%	Less than 1.5 db	10 10^{2}
2	From 3 to 10%	From 1.5 db to 4.6 db	10^{2} 10^{3}
3	From 10 to 100%	From 4.6 db to 15 db	10^{3} 10^{4}
4	Greater than 100%	Greater than 15 db	Greater than 10^{4}

Table 3-1 The development of solar proton flares

1) <u>Pre-explosive phase</u>

Hα-plage bright spots around magnetically neutral areas

Activation of dark filaments

Microwave S-component flux increase

Flare start

2) <u>Explosive phase</u>

Sudden increase of the Hα brightness (flare up)

X-ray bursts (non-thermal, 0.1-10Å)

Generation of energetic electrons (10-100 Kev)

→ Microwave bursts

→ Type III bursts

Solar blast wave (Moreton wave)

$100 \sim 1{,}000$ Km sec^{-1}

Acceleration of solar cosmic rays (⩾ 10 Mev)
(and relativistic electrons)

3) <u>Main phase</u>

Radio bursts (Type II, IVμ, IVdm, IVm A)

Ejection of solar cosmic rays and relativistic electrons

Plasma cloud - blast wave

4) <u>Last phase</u>

Stationary type IV radio bursts
(IVm B at metric frequencies)

Flare nimbus

Loop prominence system

Noise storm enhancement (mainly metric frequencies)

Table 3-2 Flare energy and mass distributions released from a typical proton flare

Emission	Particle number	Mass(g)	Energy (ergs)
$H\alpha$			10^{31}
Total line emission			5×10^{31}
Continuum emission			8×10^{31}
Total optical emission			10^{32}
Optical flare region	10^{41}	2×10^{17}	
Soft X-rays (1 - 20 Å)			2×10^{30}
Energetic X-rays ($\geqslant$ 50 Kev electrons)	10^{39}		5×10^{31}
Type IV burst (3 Mev electrons)	10^{33}		5×10^{27}
Type III burst ($\leqslant$ 100 Kev electrons)	10^{35}		10^{28}
Visible ejection ($v \sim 3 \times 10^7$ cm sec^{-1})	10^{40}	2×10^{16}	10^{31}
Energetic protons ($E \geqslant 10$ Mev)	10^{35}		2×10^{31}
Cosmic rays (1 - 30 Bev)			3×10^{31}
Interplanetary blast wave	10^{37}	2×10^{15}	2×10^{32}
Moreton wave			10^{30}

Table 4-1 Relative abundances of the chemical elements (Pottasch, 1966)

Element	Coronal Forbidden Lines	Coronal Ultraviolet Analysis	Solar Cosmic Rays	Photosphere	Stony Meterites
H	1,000,000	1,000,000		1,000,000	
He		200,000	107,000		
C		600	590	520	
N		60	190	95	
O		450	1,000	910	
Ne		50	130		
Mg		90	43	25	63
Aℓ		5		1.6	5
Si		100	33	32	63
P		0.8		0.22	0.5
S	10	14	57*	20	7
Ar	20				
K	0.7			0.55	0.3
Ca	6	3		1.4	4.4
Cr	1			0.23	0.8
Mm	0.6			0.078	0.4
Fe	8	40	20*	3.7	5.3
Co	0.3			0.043	0.2
Ni	5			0.48	3

* P,S,Cl,Ar,K,Ca,Sc

** Ti,V,Cr,Mn,Fe,Co,Ni

Table 4-2 The ratio of protons to heliums in the model of the solar atmosphere

Authors	α/P	P/α
Sears (1964)		
Demarque and Percy (1964)	0.095	10.52
Weyman and Sears (1965)	0.0865	11.56
Morton (1967)	0.077-0.087	11.5-13.0

Table 4-3 Neutron production reactions (Lingenfelter and Ramaty, 1967)

Reaction	Threshold Energy (Mev/nucleon)
$H^1(P, n\pi^+) H^1$	292.3
$He^4(P, nP) He^3$	25.9
$He^4(P, 2Pn) H^2$	32.8
$He^4(P, 2P2n) H^1$	35.6
$C^{12}(P, n\cdots$	19.8
$N^{14}(P, n\cdots$	6.3
$O^{16}(P, Pn\cdots$	16.5
$Ne^{20}(P, Pn\cdots$	17.7

Table 4-4 Nuclear reactions producing deuterons and heliums, He^3

Reaction process	Threshold energy (Mev/nucleon)
Deuteron production reactions	
$H^1(P,\pi^+)\ H^2$	284.9
$He^4(P,He^3)\ H^2$	23.0
$He^4(P,2Pn)\ H^2$	32.8
$He^4(P,Pd)\ H^2$	30.0
C (P,d···	17.9
N (P,d···	8.9
O (P,d···	14.2
Ne (P,d···	15.4
Helium-3 production reactions	
$He^4(P,d)\ He^3$	23.0
$He^4(P,Pn)\ He^3$	25.9
C (P, He^3···	21.3
N (P, He^3···	5.1
O (P, He^3···	16.2
Ne (P, He^3···	16.3

d: deuteron ($=H^2$)

Table 4-5 Pion production reactions (Lingenfelter and Ramaty, 1967)

$P + P \longrightarrow$	$H^2 + \pi^+$
	$P + P + a(\pi^+ + \pi^-) + b\pi^o$
	$P + n + \pi^+ + a(\pi^+ + \pi^-) + b\pi^o$
	$2n + 2\pi^+ + a(\pi^+ + \pi^-) + b\pi^o$
$P + He^4 \longrightarrow$	$P + He^4 + a(\pi^+ + \pi^-) + b\pi^o$
	$P + He^3 + n + a(\pi^+ + \pi^-) + b\pi^o$
	$2P + H^2 + \pi^- + a(\pi^+ + \pi^-) + b\pi^o$
	$4P + n + \pi^- + a(\pi^+ + \pi^-) + b\pi^o$
	$3P + 2n + a(\pi^+ + \pi^-) + b\pi^o$
	$2P + 3n + a(\pi^+ + \pi^-) + b\pi^o$
	$P + 4n + 2\pi^+ + a(\pi^+ + \pi^-) + b\pi^o$

a,b: integers

Caption of Figures

Fig. 2-1. Solar cosmic ray events investigated by Forbush (1946). (a) 28 February and 7 March, 1942 and (b) 25 July, 1946.

Fig. 2-2. Neutron monitor record of the solar cosmic ray event of 23 February 1956 at Chicago (Simpson, 1960 a).

Fig. 2-3. Examples of solar cosmic ray events (a) 4 May 1960, (b) 12-13 November 1960 and (c) 7 July 1966.

Fig. 2-4. The 20 November 1960 event produced by the solar flare on the invisible hemisphere of the sun (Carmichael, 1962).

Fig. 2-5. The travel times of solar cosmic rays between the sun and the earth as a function of the parent flare positions on the solar disk. "S" indicates "small increase" event.

Fig. 2-6. Two types of Mev solar cosmic ray events classified as F and S types (Sinno, 1961).

Fig. 2-7. Classification of Mev solar cosmic ray events as observed by riometers (Leinback, 1962).

Fig. 5-3. Various energy loss processes occurring during electron acceleration: solid line, gyro-synchrotron loss; broken line, ionization loss; and chain line, bremsstrahlung loss.

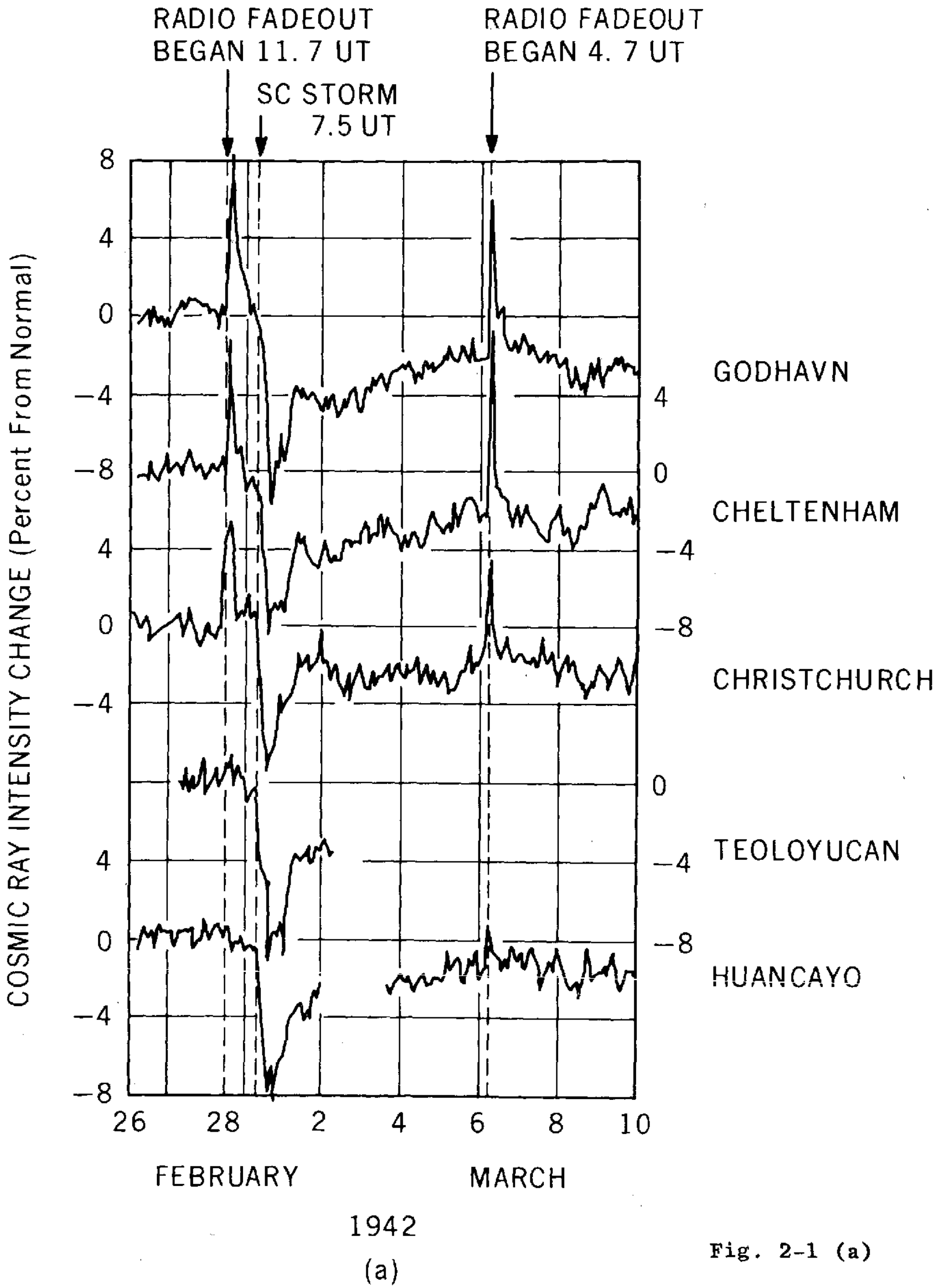

RADIO FADEOUT
BEGAN 11.7 UT
SC STORM
7.5 UT
RADIO FADEOUT
BEGAN 4.7 UT
COSMIC RAY INTENSITY CHANGE (Percent From Normal)
GODHAVN
CHELTENHAM
CHRISTCHURCH
TEOLOYUCAN
HUANCAYO
8
4
0
−4
−8
26
28
2
4
6
8
10
FEBRUARY
MARCH
1942
(a)

Fig. 2-1 (a)

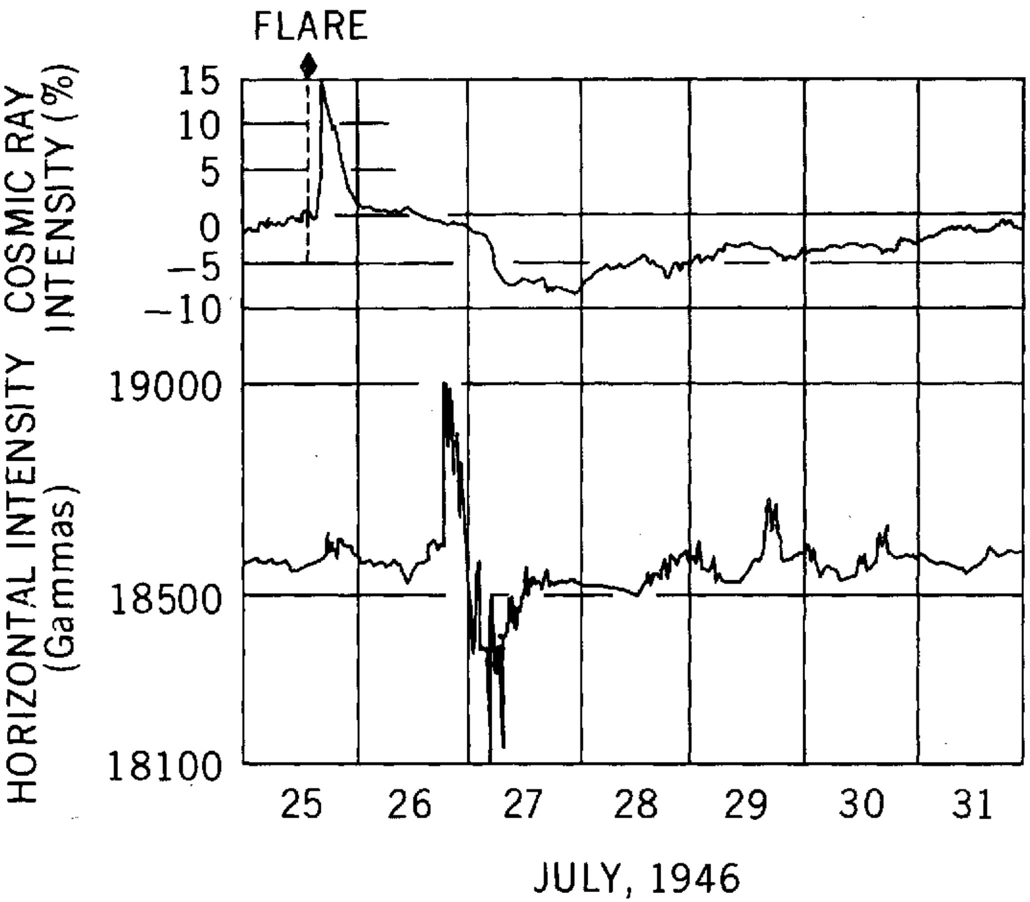

FLARE
COSMIC RAY INTENSITY (%)
15
10
5
0
−5
−10
HORIZONTAL INTENSITY (Gammas)
19000
18500
18100
25
26
27
28
29
30
31
JULY, 1946

Fig. 2-1 (b)

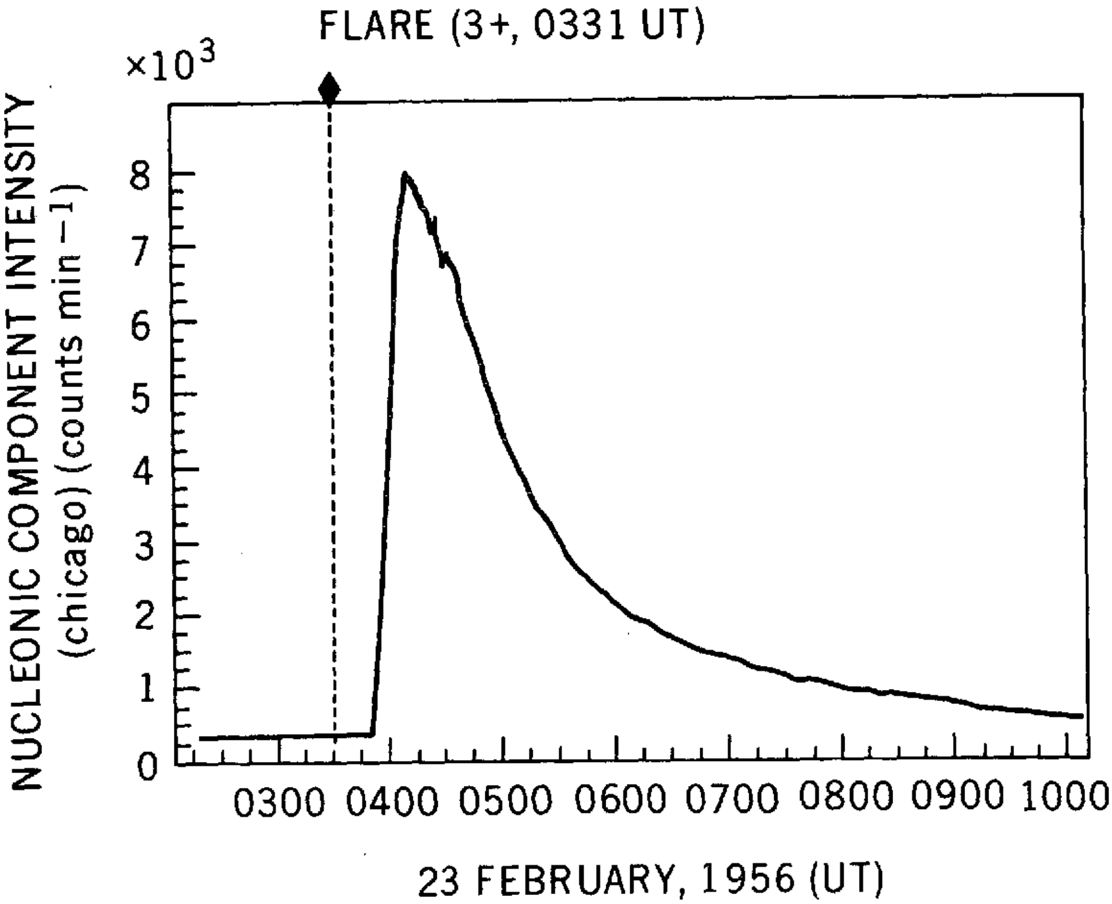

FLARE (3+, 0331 UT)
×10^3
NUCLEONIC COMPONENT INTENSITY
(chicago) (counts min −1)
0
1
2
3
4
5
6
7
8
0300 0400 0500 0600 0700 0800 0900 1000
23 FEBRUARY, 1956 (UT)

Fig. 2-2

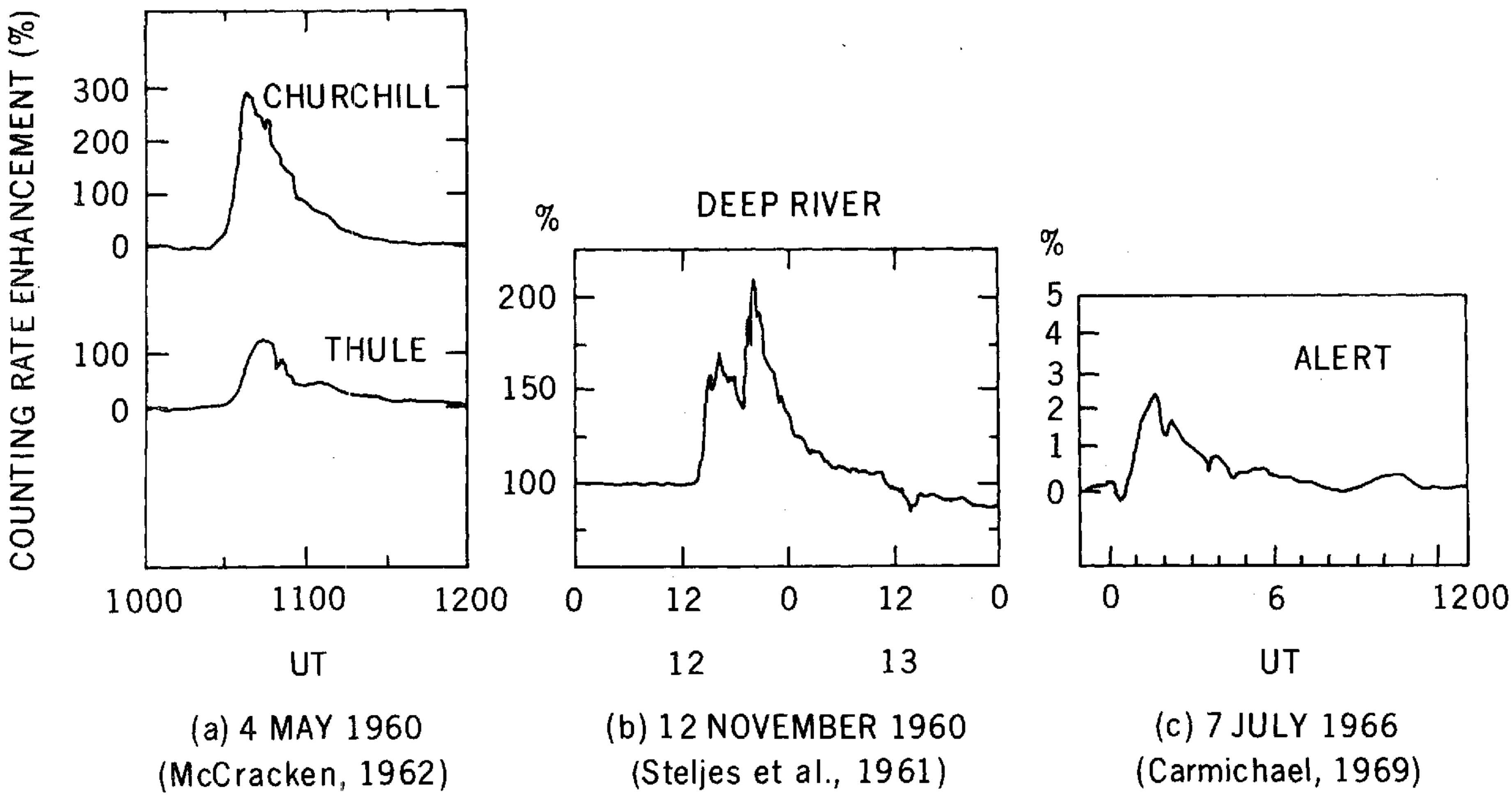

COUNTING RATE ENHANCEMENT (%)
300
200
100
0
100
0
CHURCHILL
THULE
1000
1100
1200
UT
(a) 4 MAY 1960
(McCracken, 1962)
%
DEEP RIVER
200
150
100
0
12
0
12
0
12
13
(b) 12 NOVEMBER 1960
(Steljes et al., 1961)
%
5
4
3
2
1
0
ALERT
0
6
1200
UT
(c) 7 JULY 1966
(Carmichael, 1969)

Fig. 2-3

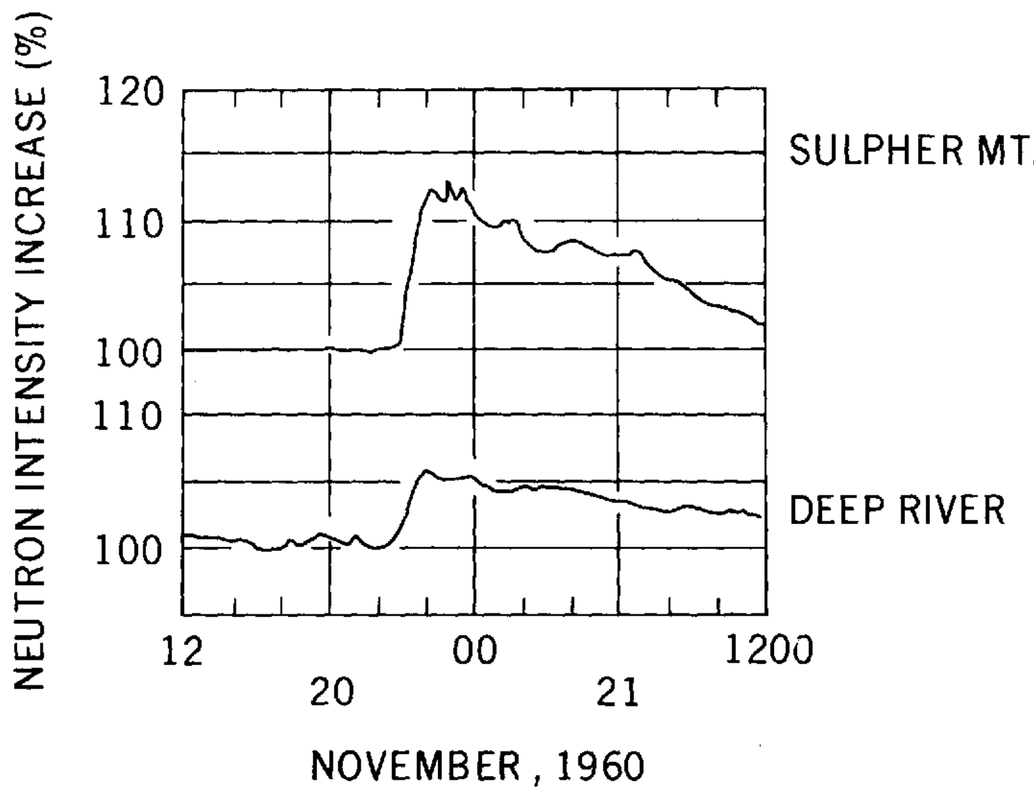
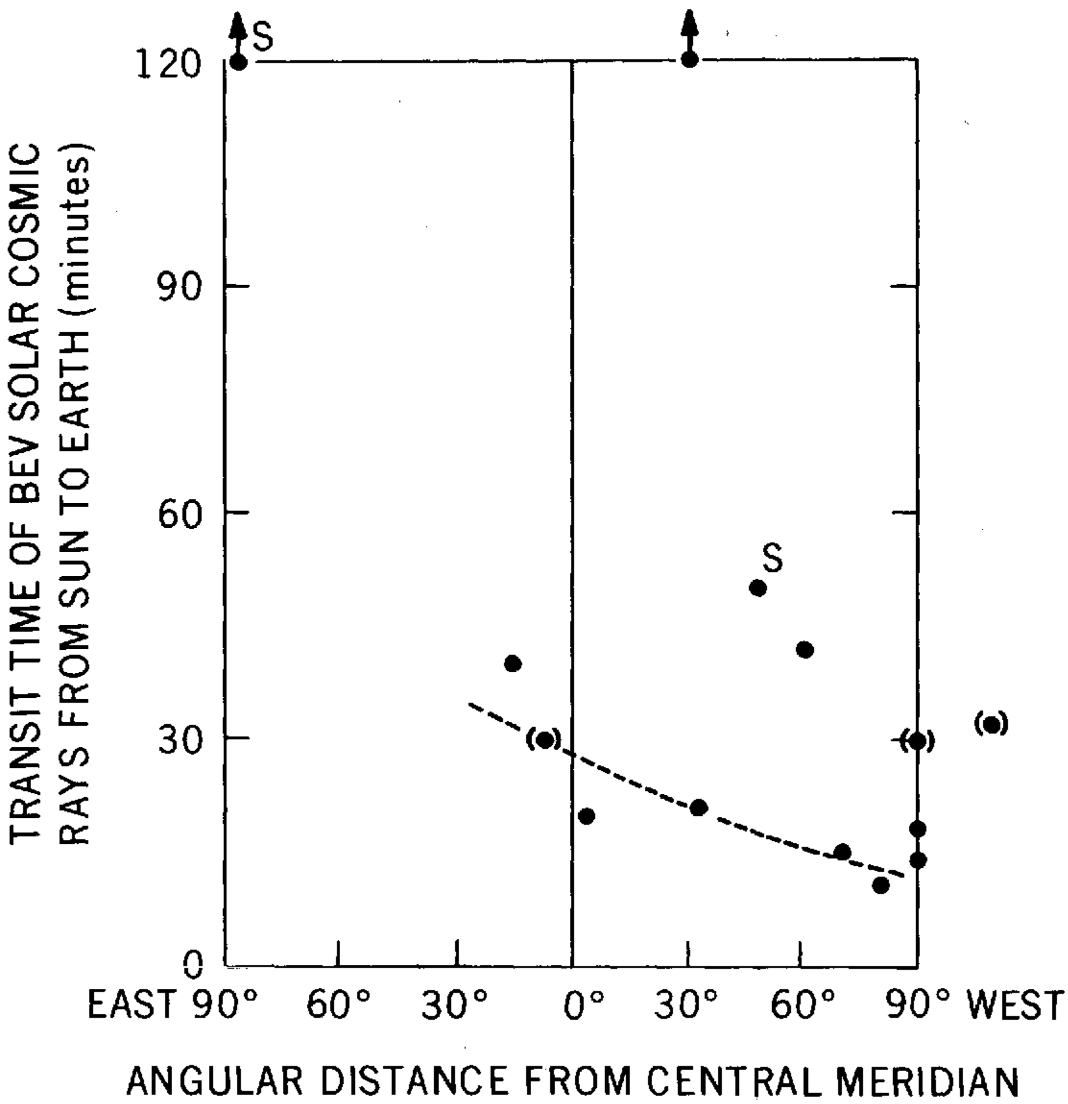

NEUTRON INTENSITY INCREASE (%)
120
110
100
110
100
SULPHER MT.
DEEP RIVER
12
00
1200
20
21
NOVEMBER, 1960

Fig. 2-4



TRANSIT TIME OF BEV SOLAR COSMIC RAYS FROM SUN TO EARTH (minutes)
120
90
60
30
0
S
S
(•)
(•)
(•)
EAST 90°
60°
30°
0°
30°
60°
90° WEST
ANGULAR DISTANCE FROM CENTRAL MERIDIAN

Fig. 2-5

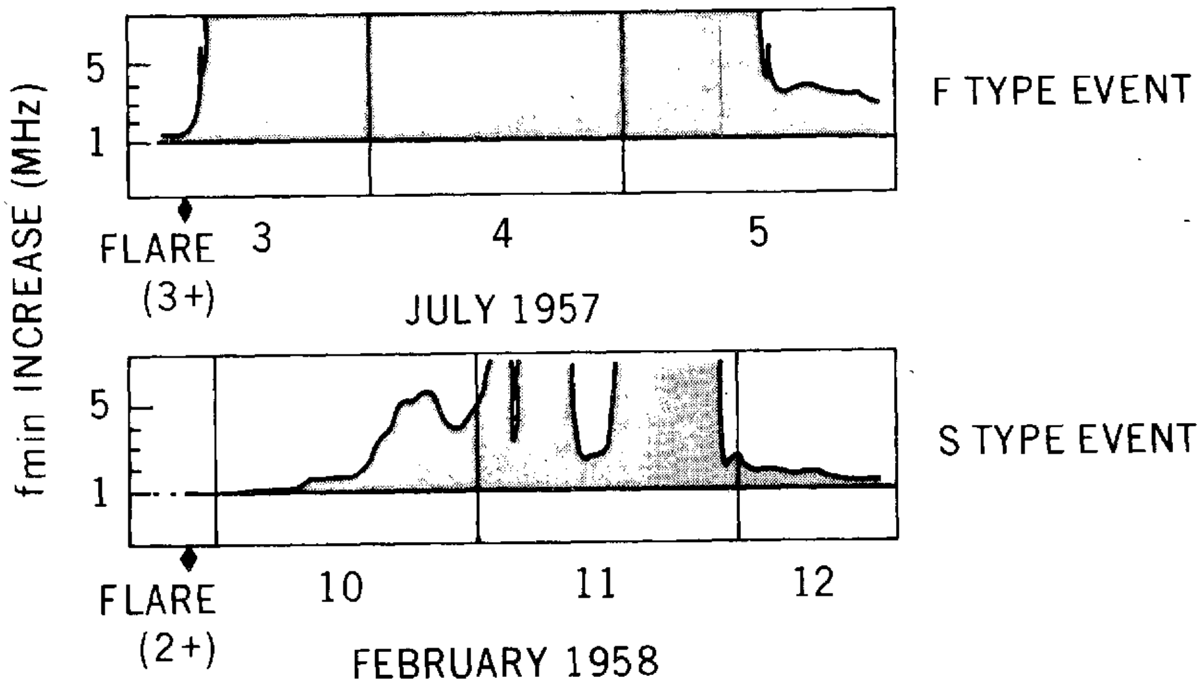

fmin INCREASE (MHz)
5
1
F TYPE EVENT
FLARE
(3+)
3
4
5
JULY 1957
5
1
S TYPE EVENT
FLARE
(2+)
10
11
12
FEBRUARY 1958

Fig. 2-6

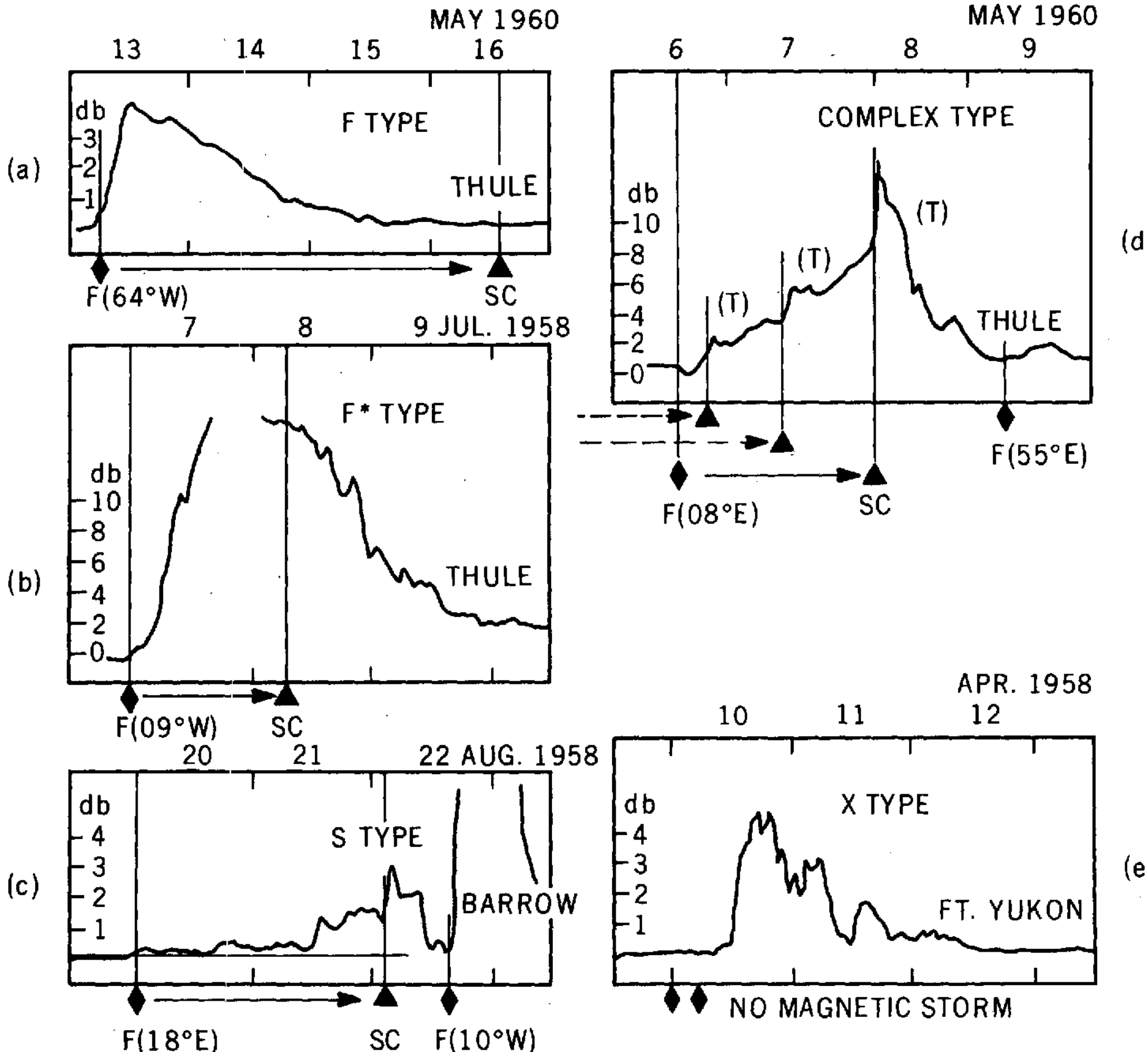

MAY 1960
13 14 15 16
db
3
2
1
F TYPE
THULE
F(64°W)
SC
(a)
7 8 9 JUL. 1958
db
10
8
6
4
2
0
F* TYPE
THULE
F(09°W)
SC
(b)
20 21 22 AUG. 1958
db
4
3
2
1
S TYPE
BARROW
F(18°E)
SC
F(10°W)
(c)
MAY 1960
6 7 8 9
COMPLEX TYPE
db
10
8
6
4
2
0
(T)
(T)
(T)
THULE
F(55°E)
F(08°E)
SC
(d)
APR. 1958
10 11 12
db
4
3
2
1
X TYPE
FT. YUKON
NO MAGNETIC STORM
(e)

Fig. 2-7

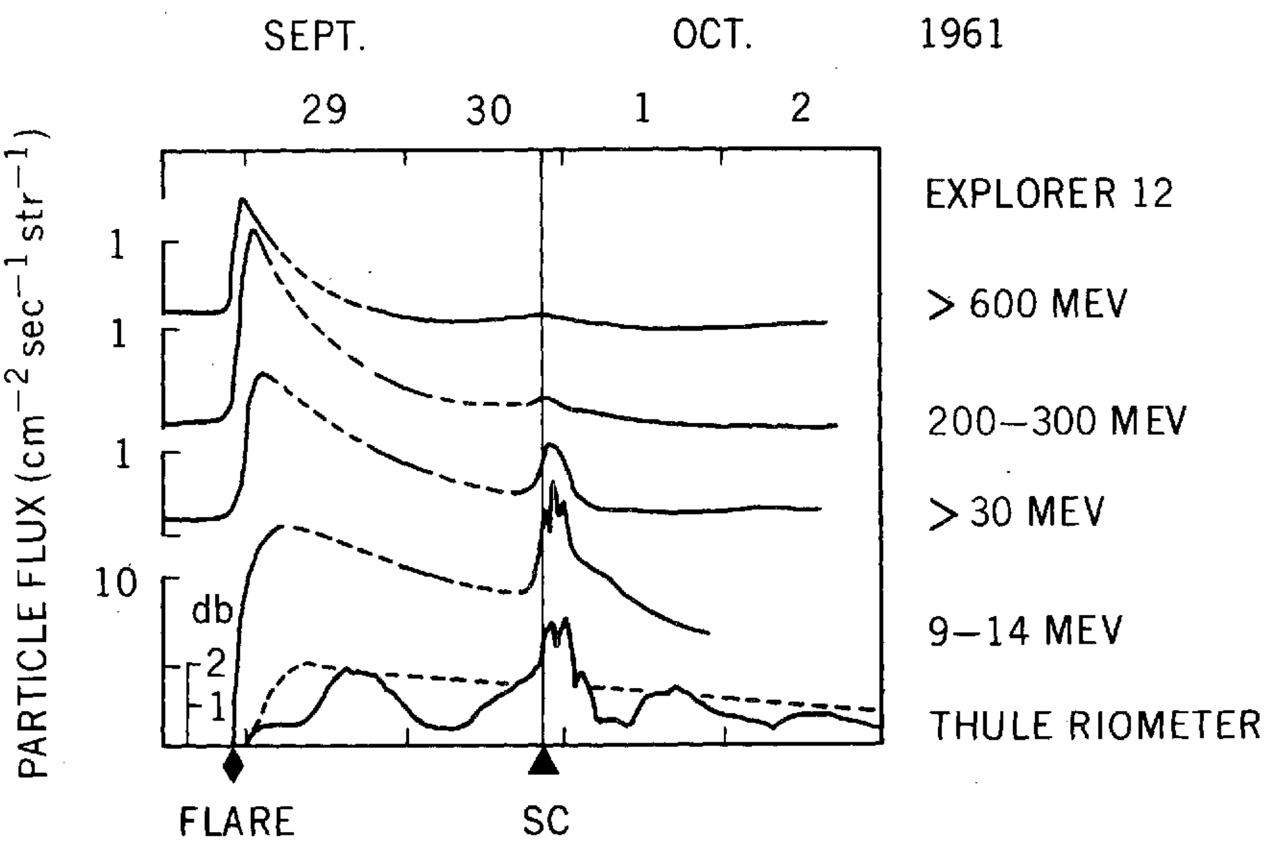
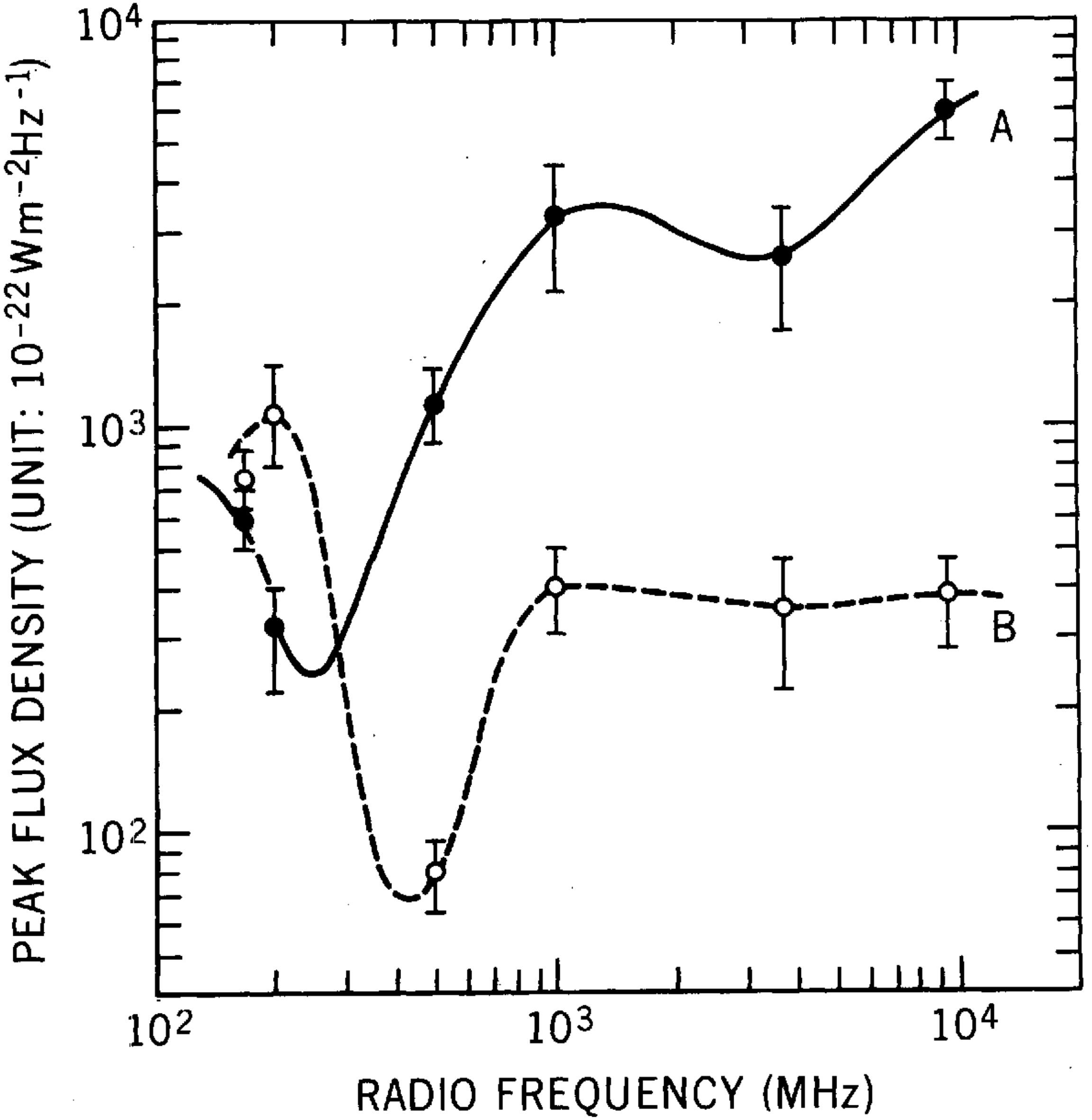

SEPT.
OCT.
1961
29
30
1
2
EXPLORER 12
> 600 MEV
200–300 MEV
> 30 MEV
9–14 MEV
THULE RIOMETER
PARTICLE FLUX (cm^{-2} sec^{-1} str^{-1})
1
1
1
10
db
2
1
FLARE
SC

Fig. 2-8



10^4
10^3
10^2
PEAK FLUX DENSITY (UNIT: 10^{-22} $Wm^{-2}Hz^{-1}$)
A
B
10^2
10^3
10^4
RADIO FREQUENCY (MHz)

Fig. 2-9

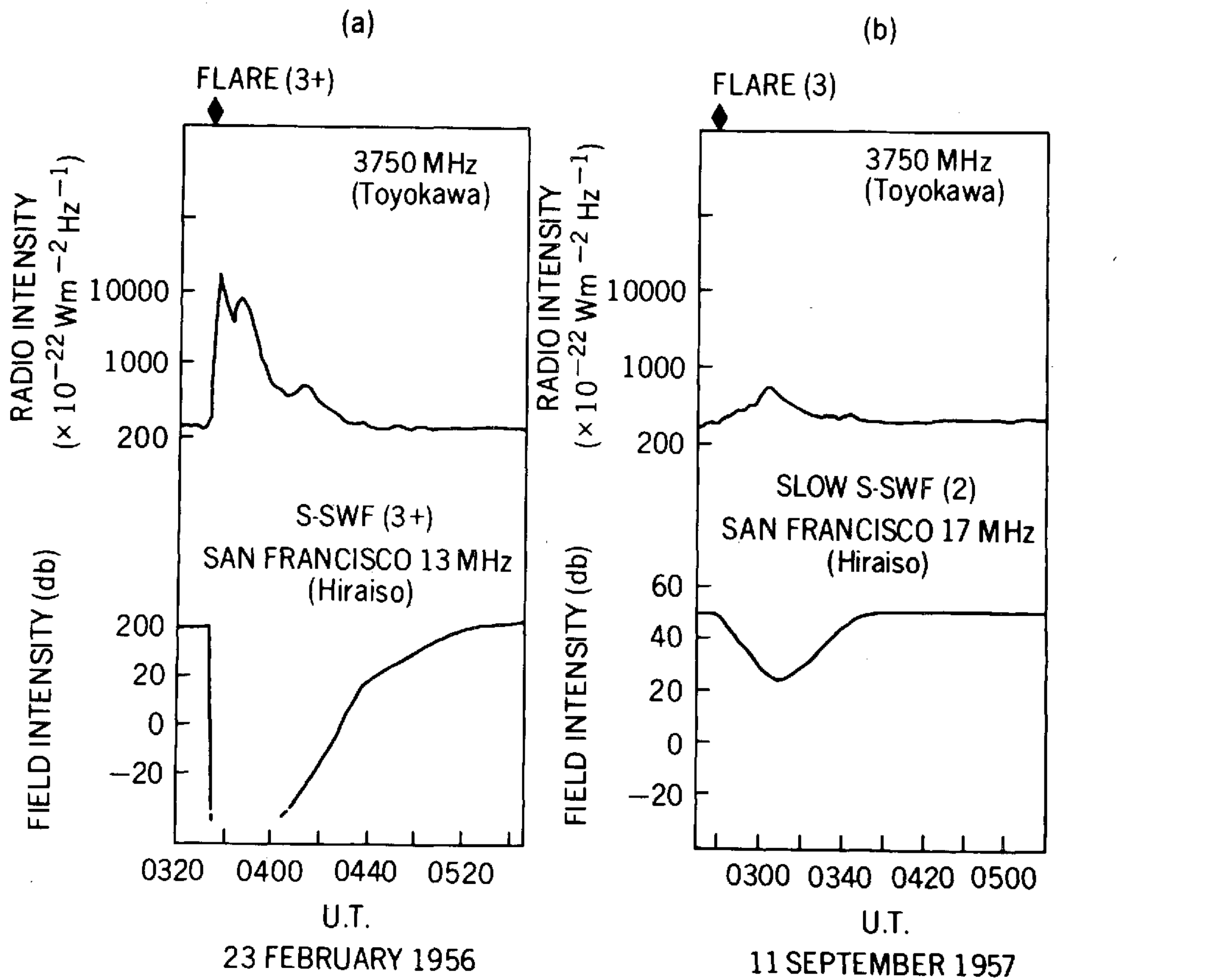

(a)
FLARE (3+)
3750 MHz
(Toyokawa)
RADIO INTENSITY
($\times 10^{-22}$ Wm^{-2} Hz^{-1})
10000
1000
200
S-SWF (3+)
SAN FRANCISCO 13 MHz
(Hiraiso)
FIELD INTENSITY (db)
200
20
0
−20
0320
0400
0440
0520
U.T.
23 FEBRUARY 1956
(b)
FLARE (3)
3750 MHz
(Toyokawa)
RADIO INTENSITY
($\times 10^{-22}$ Wm^{-2} Hz^{-1})
10000
1000
200
SLOW S-SWF (2)
SAN FRANCISCO 17 MHz
(Hiraiso)
FIELD INTENSITY (db)
60
40
20
0
−20
0300
0340
0420
0500
U.T.
11 SEPTEMBER 1957

Fig. 2-10

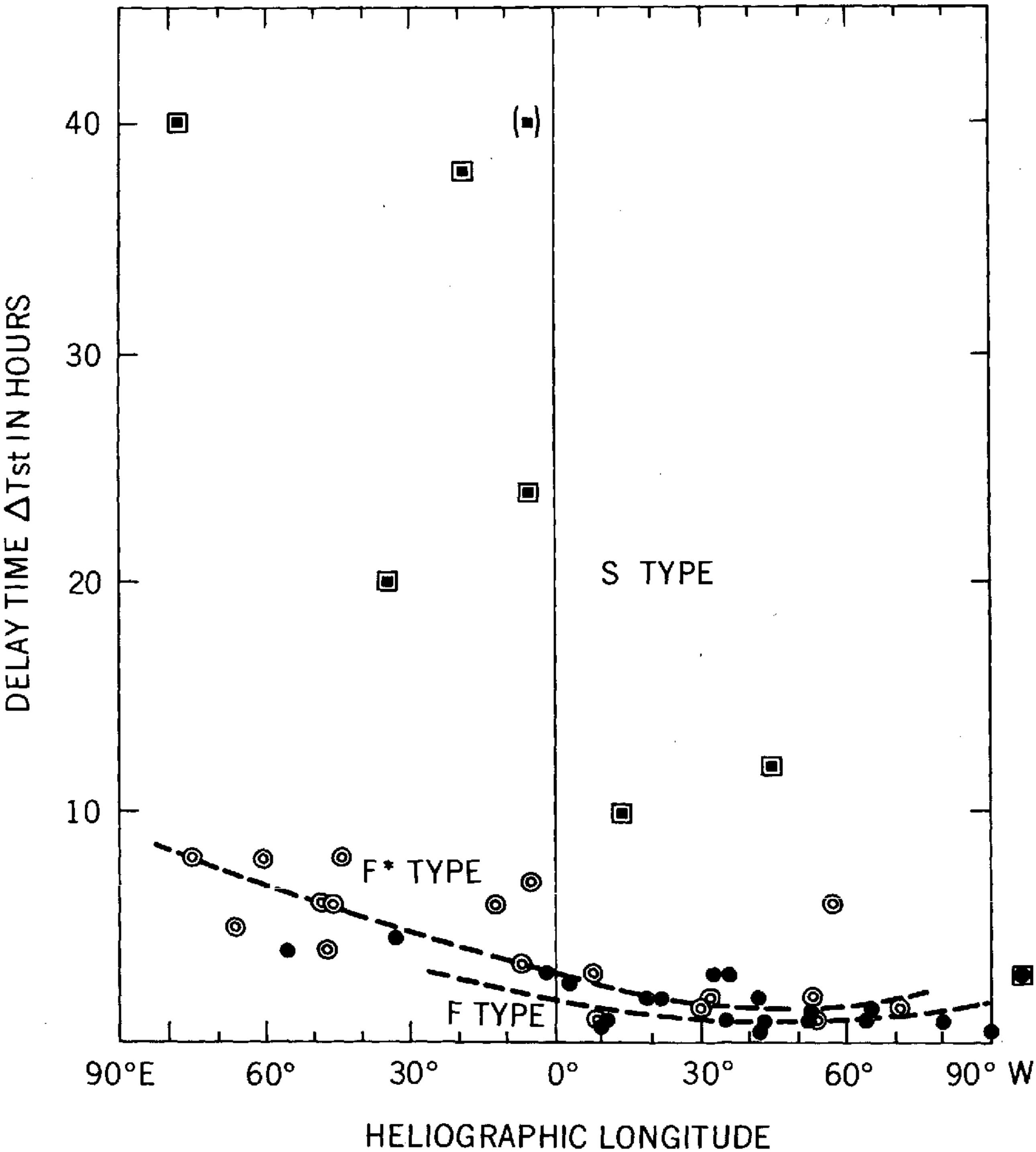

DELAY TIME ΔTst IN HOURS
40
30
20
10
S TYPE
F* TYPE
F TYPE
90°E
60°
30°
0°
30°
60°
90° W
HELIOGRAPHIC LONGITUDE

Fig. 2-11

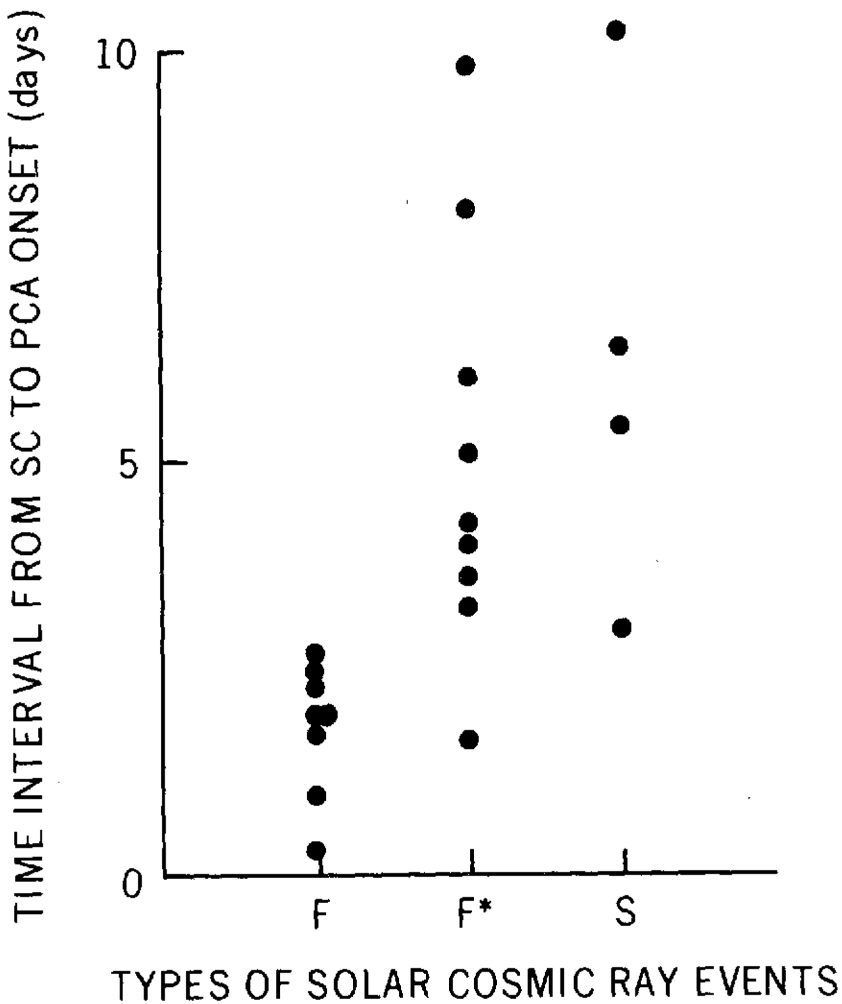
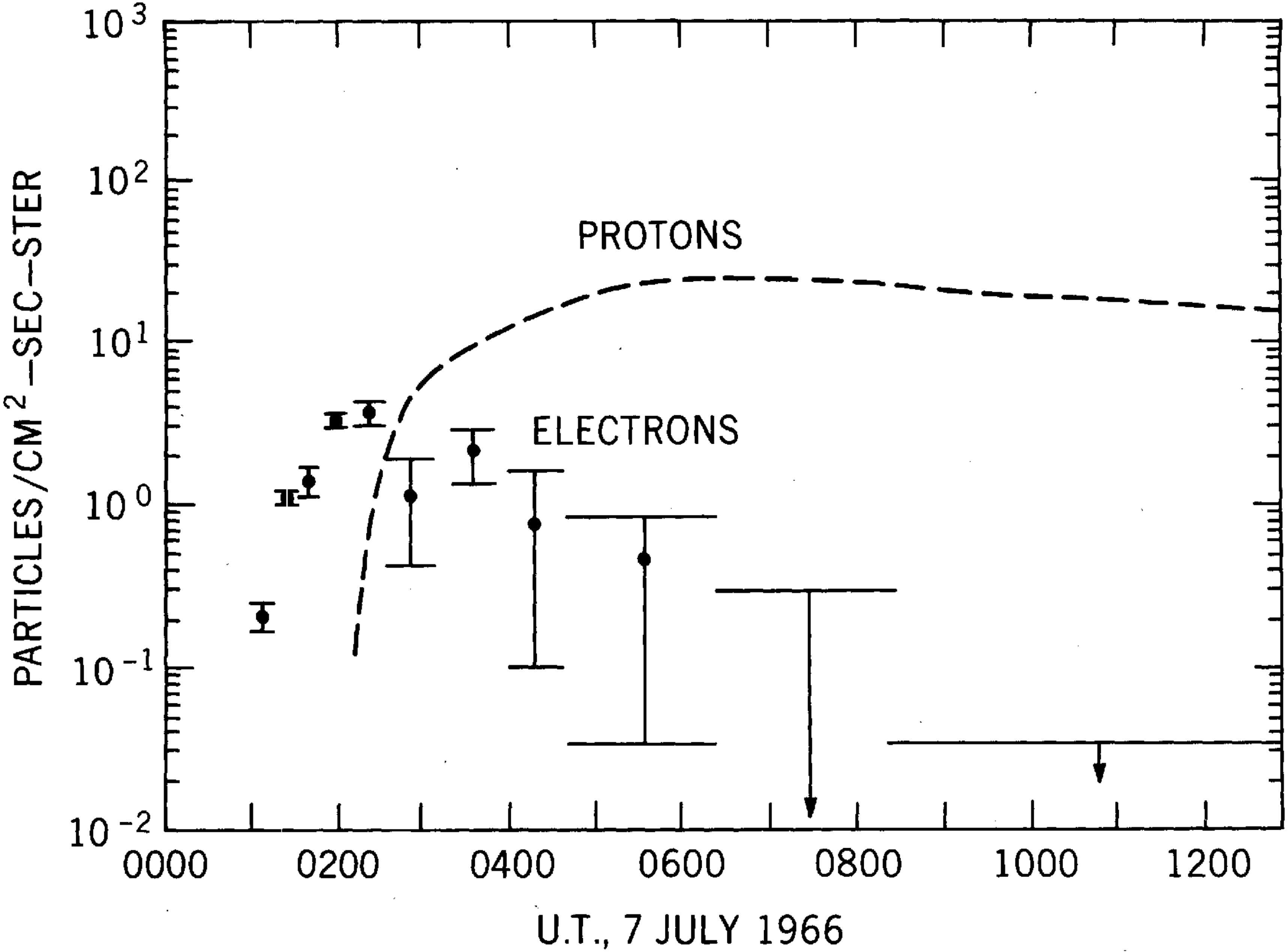

TIME INTERVAL FROM SC TO PCA ONSET (days)
10
5
0
F
F*
S
TYPES OF SOLAR COSMIC RAY EVENTS

Fig. 2-12



PARTICLES/CM² –SEC–STER
10³
10²
10¹
10⁰
10⁻¹
10⁻²
PROTONS
ELECTRONS
0000
0200
0400
0600
0800
1000
1200
U.T., 7 JULY 1966

Fig. 2-13

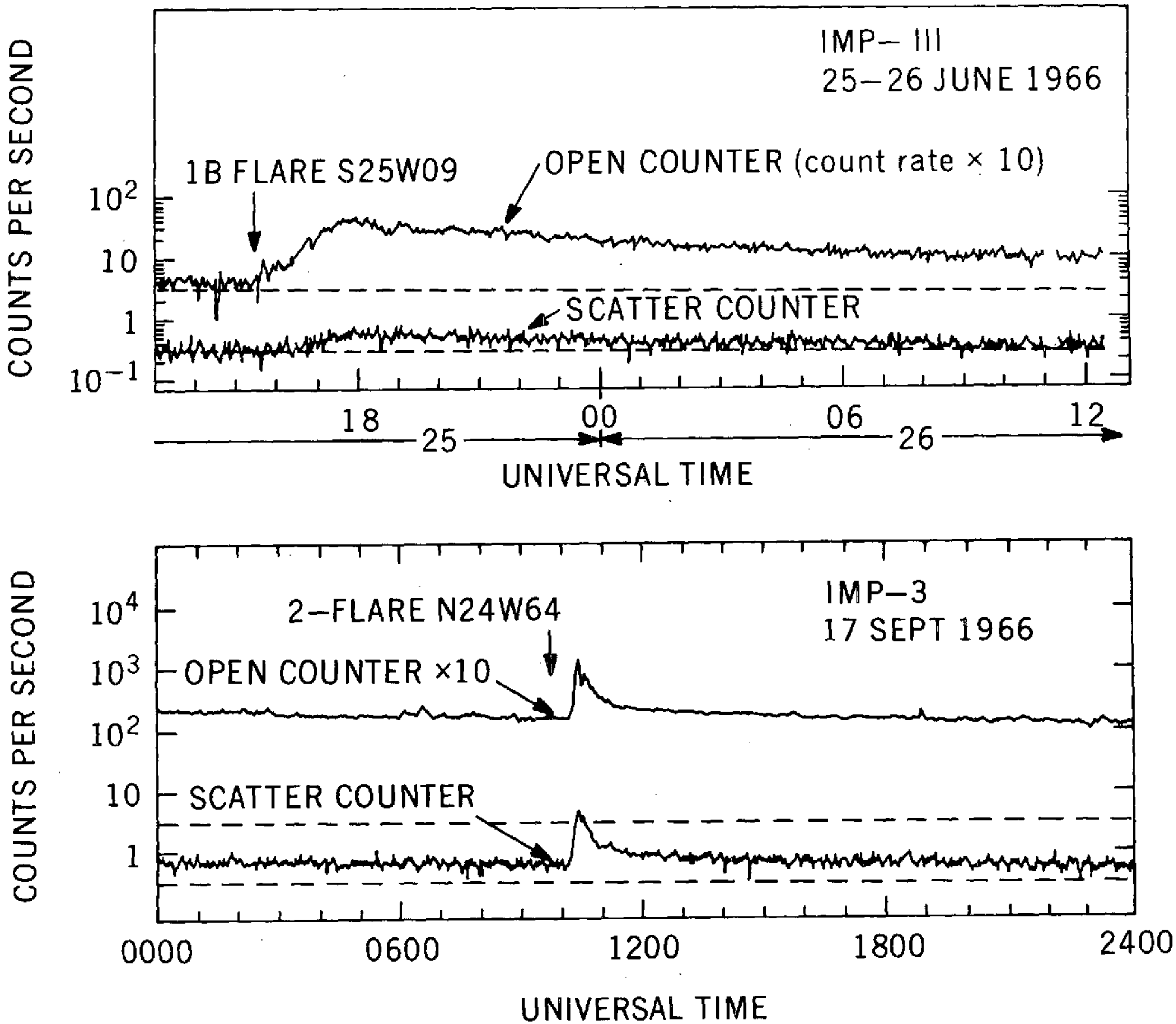

IMP– III
25–26 JUNE 1966
1B FLARE S25W09
OPEN COUNTER (count rate × 10)
SCATTER COUNTER
COUNTS PER SECOND
10^2
10
1
10^{-1}
18
25
00
06
26
12
UNIVERSAL TIME
IMP–3
17 SEPT 1966
2–FLARE N24W64
OPEN COUNTER ×10
SCATTER COUNTER
COUNTS PER SECOND
10^4
10^3
10^2
10
1
0000
0600
1200
1800
2400
UNIVERSAL TIME

Fig. 2-14

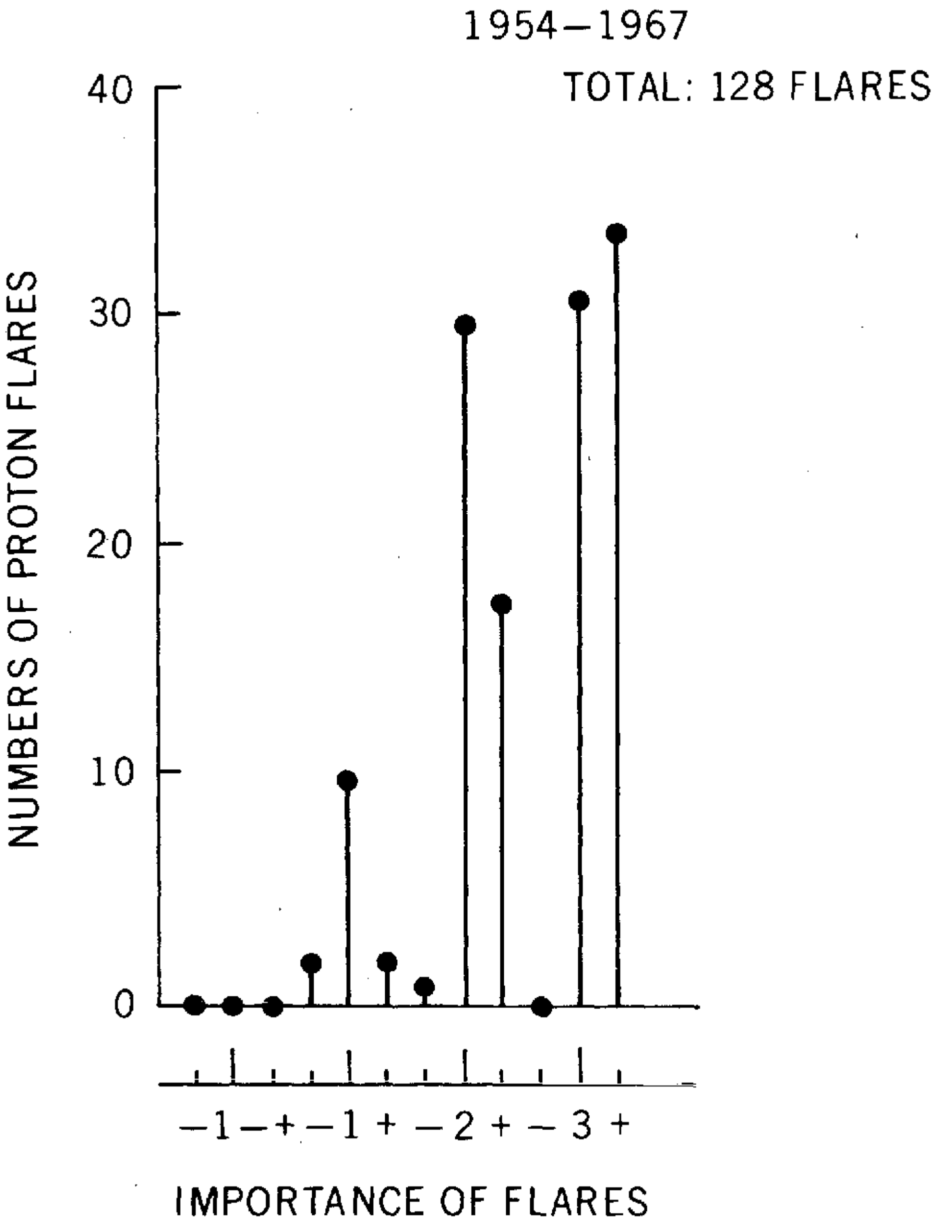

1954—1967
TOTAL: 128 FLARES
40
30
20
10
0
NUMBERS OF PROTON FLARES
—1—+—1 + — 2 + — 3 +
IMPORTANCE OF FLARES

Fig. 3-1

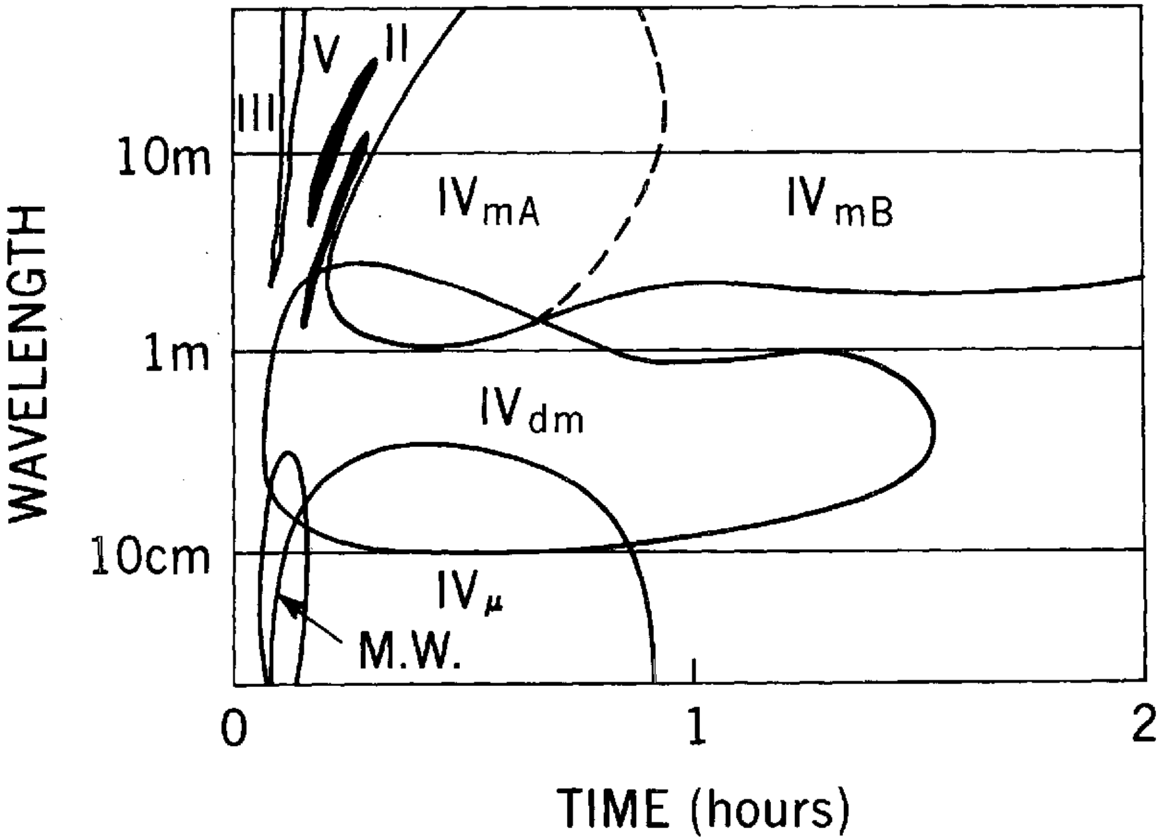
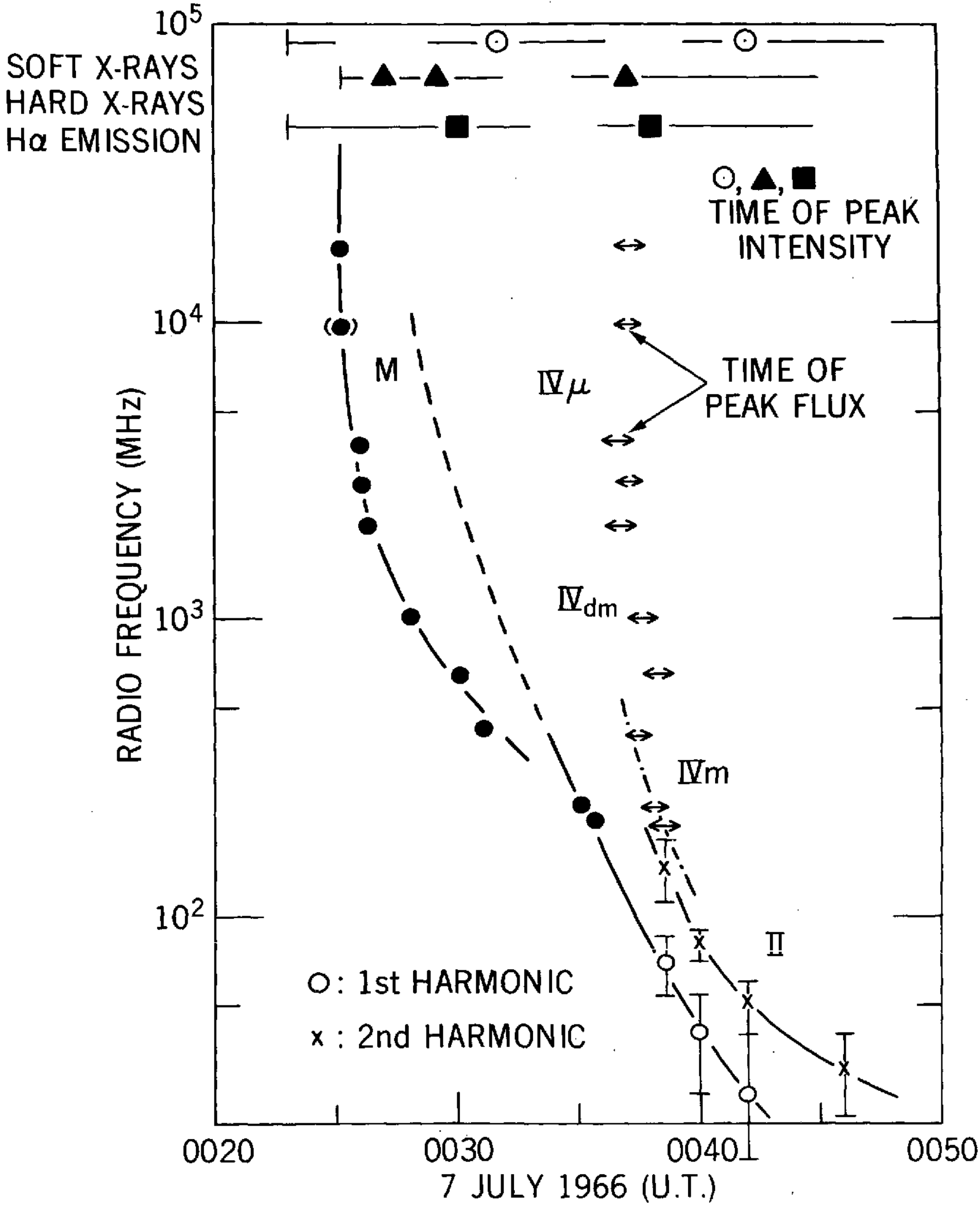
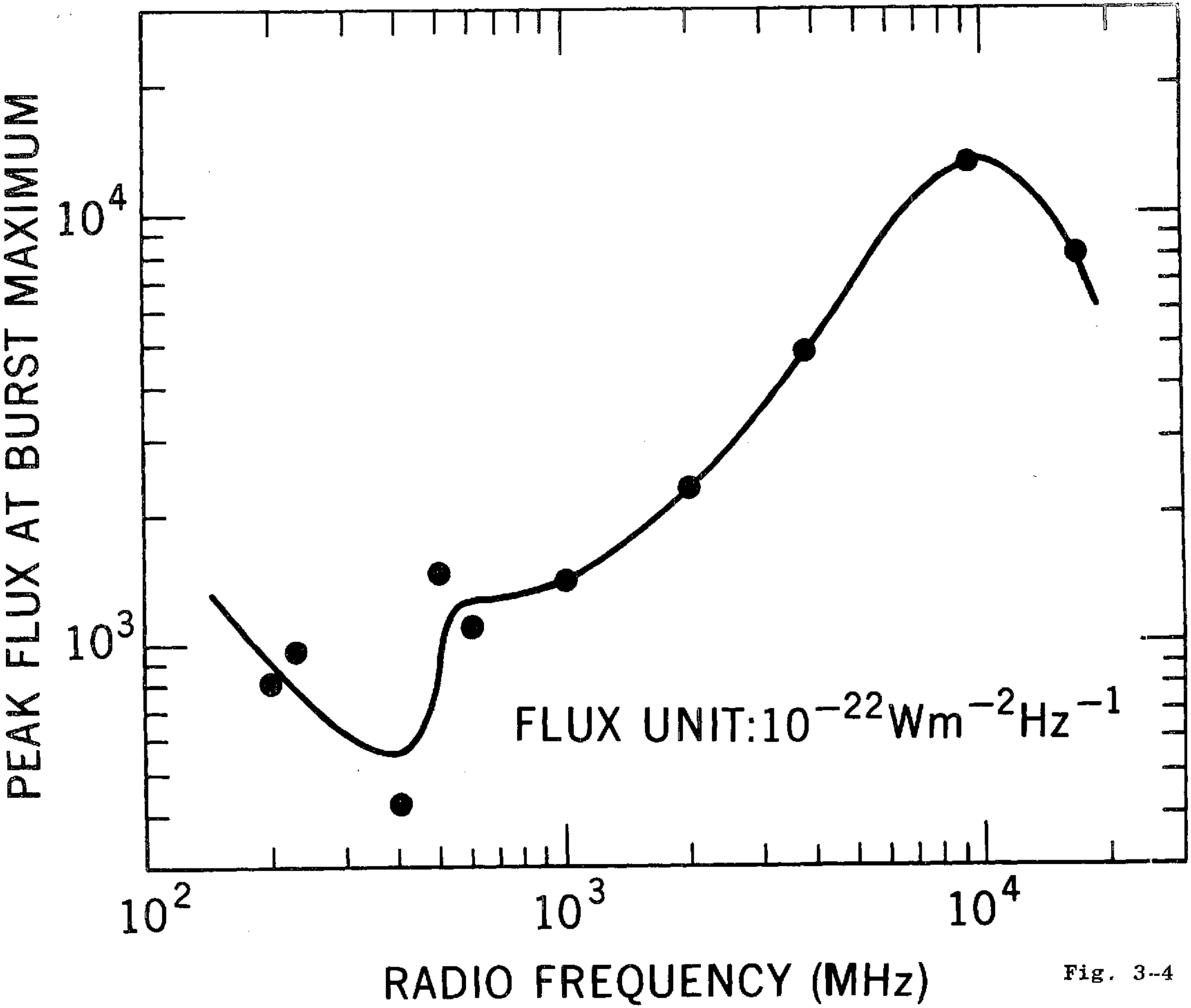

III
V
II
10m
1m
10cm
IV$_{mA}$
IV$_{mB}$
IV$_{dm}$
IV$_{\mu}$
M.W.
WAVELENGTH
0
1
2
TIME (hours)

Fig. 3-2



10^5
SOFT X-RAYS
HARD X-RAYS
Hα EMISSION
TIME OF PEAK INTENSITY
10^4
M
IVμ
TIME OF PEAK FLUX
RADIO FREQUENCY (MHz)
IVdm
10^3
IVm
10^2
II
○: 1st HARMONIC
x : 2nd HARMONIC
0020
0030
0040
0050
7 JULY 1966 (U.T.)

Fig. 3-3



PEAK FLUX AT BURST MAXIMUM
10^4
10^3
FLUX UNIT: $10^{-22}\,Wm^{-2}Hz^{-1}$
10^2
10^3
10^4
RADIO FREQUENCY (MHz)

Fig. 3-4

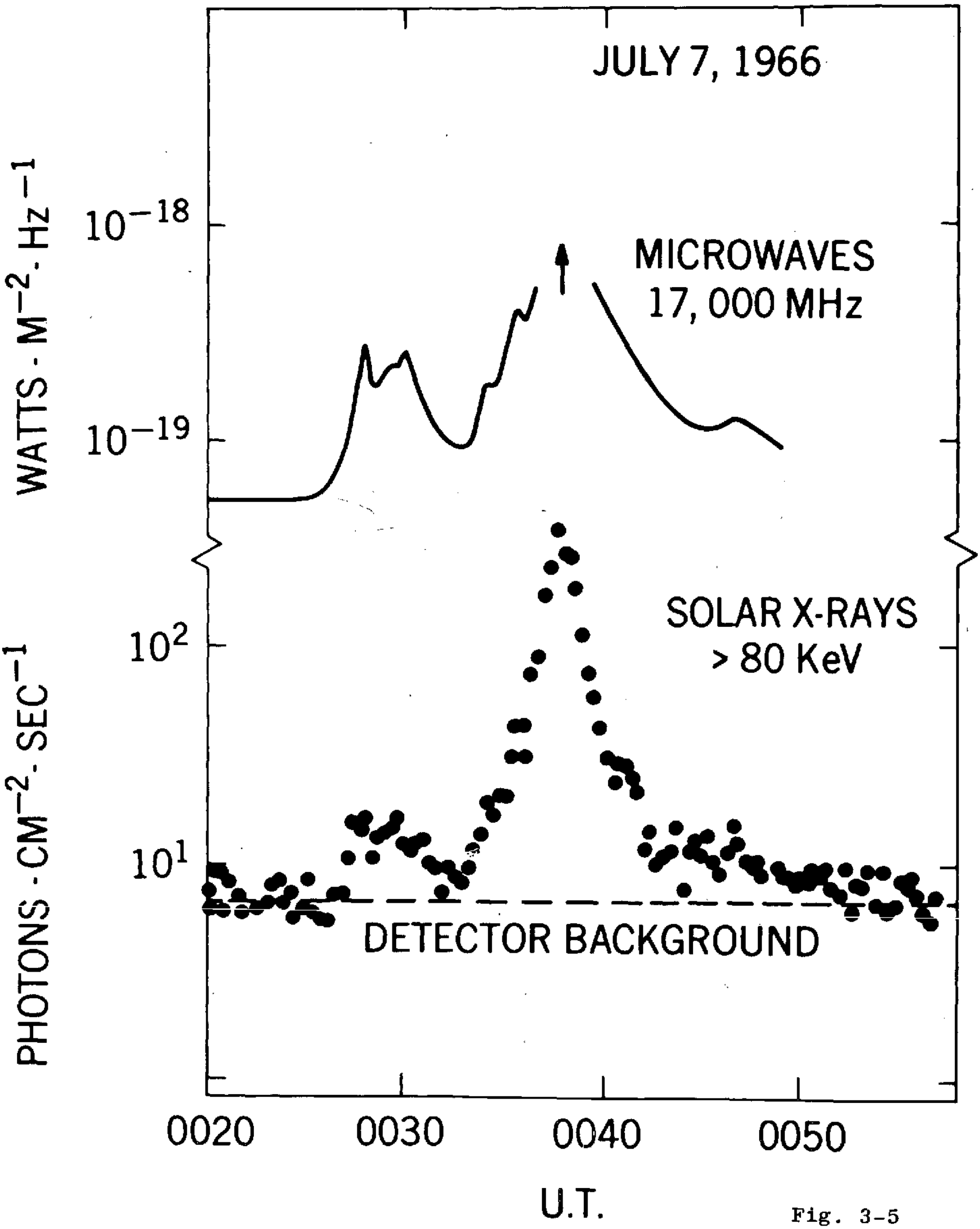

JULY 7, 1966
MICROWAVES
17,000 MHz
SOLAR X-RAYS
> 80 KeV
DETECTOR BACKGROUND
WATTS - M^{-2} - Hz^{-1}
10^{-18}
10^{-19}
PHOTONS - CM^{-2} - SEC^{-1}
10^2
10^1
0020
0030
0040
0050
U.T.

Fig. 3-5

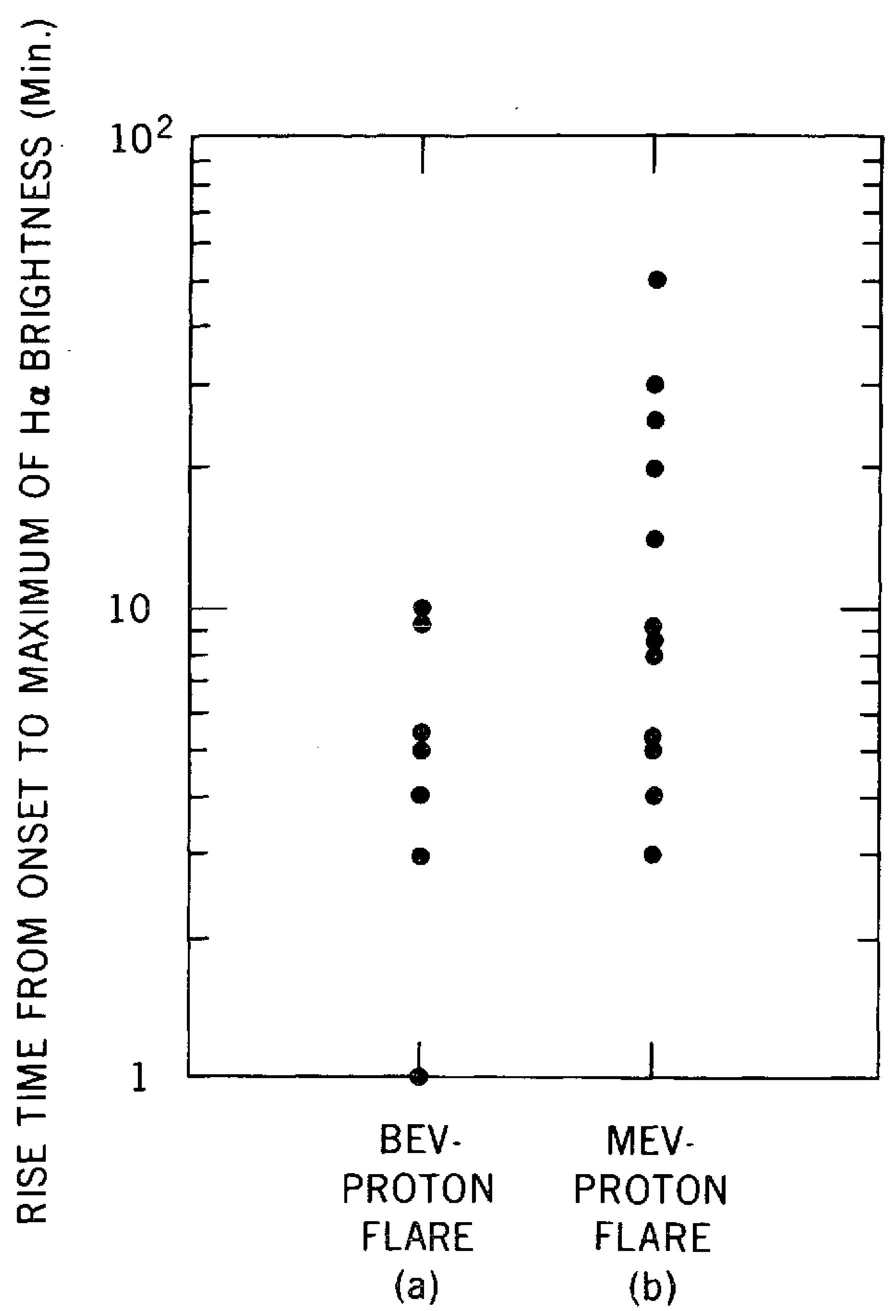

RISE TIME FROM ONSET TO MAXIMUM OF Hα BRIGHTNESS (Min.)
10²
10
1
BEV-
PROTON
FLARE
(a)
MEV-
PROTON
FLARE
(b)

Fig. 3-6

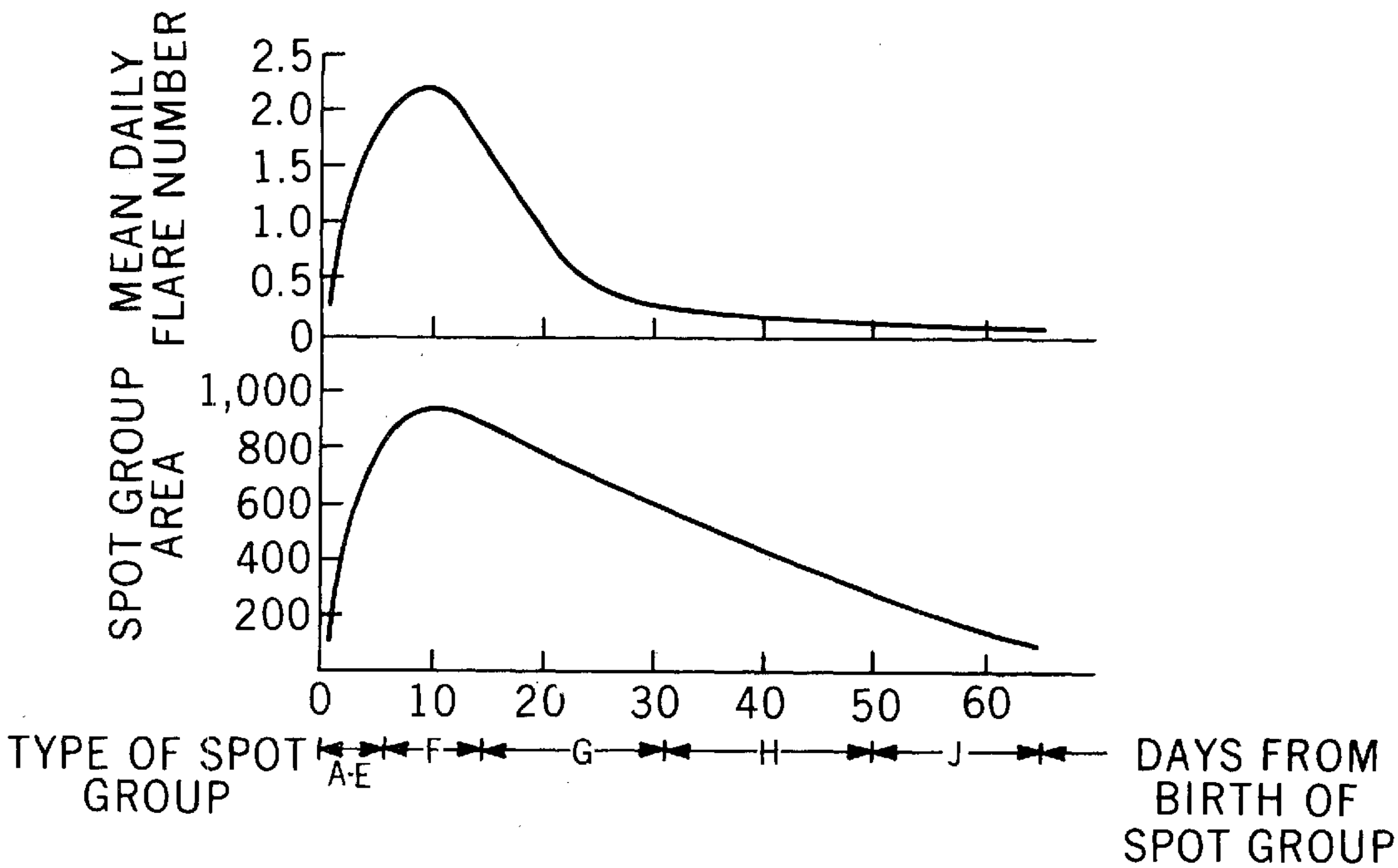

MEAN DAILY FLARE NUMBER
2.5
2.0
1.5
1.0
0.5
0
SPOT GROUP AREA
1,000
800
600
400
200
0
10
20
30
40
50
60
TYPE OF SPOT GROUP
A-E
F
G
H
J
DAYS FROM BIRTH OF SPOT GROUP

Fig. 3-7

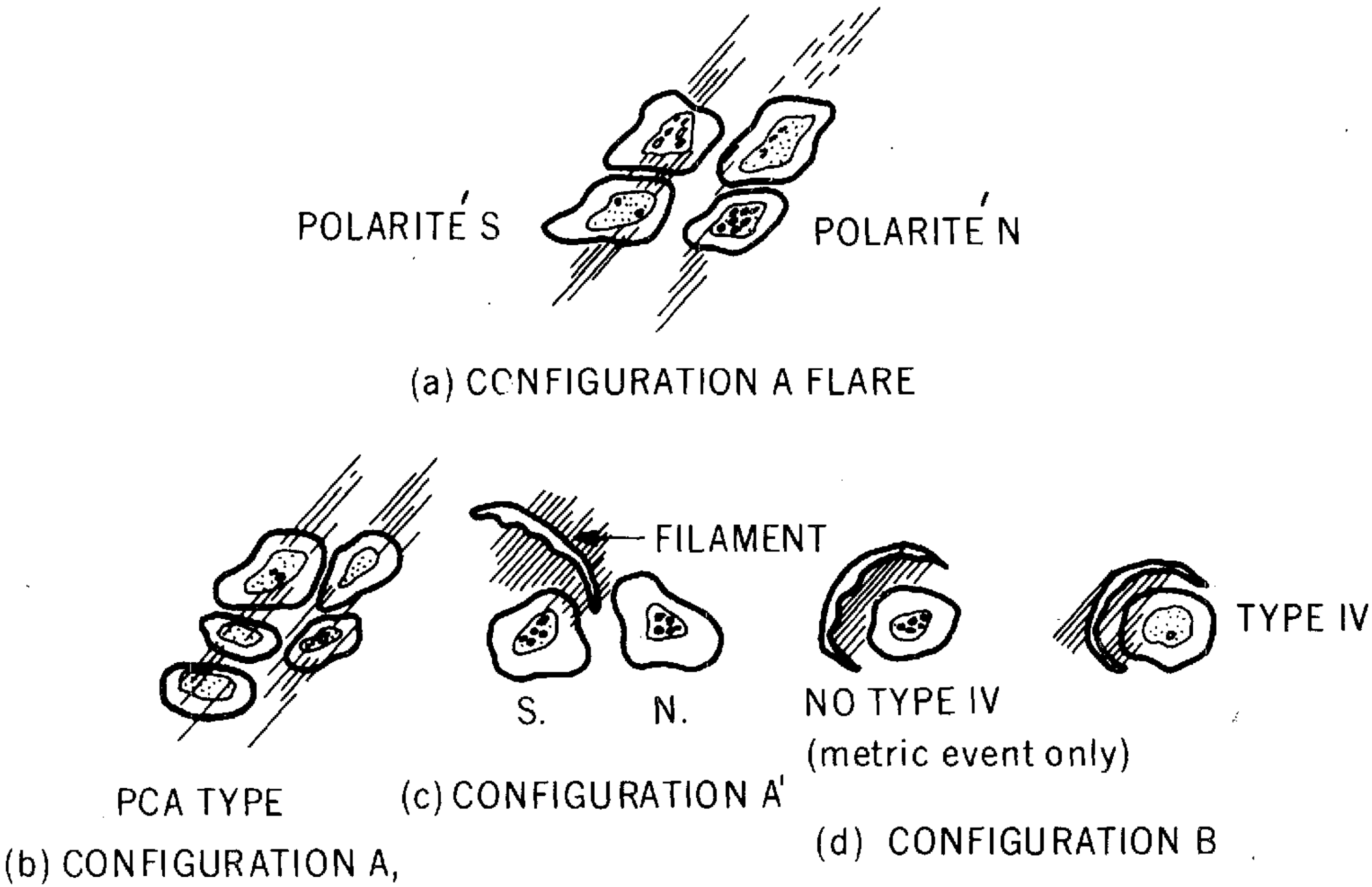

POLARITÉ S
POLARITÉ N
(a) CONFIGURATION A FLARE
FILAMENT
TYPE IV
S.
N.
NO TYPE IV
(metric event only)
PCA TYPE
(c) CONFIGURATION A'
(d) CONFIGURATION B
(b) CONFIGURATION A,

Fig. 3-8

MAGNETIC FIELD DISTRIBUTION OF SUNSPOT GROUPS WHICH PRODUCED PROTON FLARES IN THE NORTHERN HEMISPHERE

TYPE I

(a) JULY 7, 1958

SUNSPOT GROUP
NUMBER 13356

TYPE II

(b) SEPTEMBER 20, 1963

SUNSPOT GROUP
NUMBER 15768

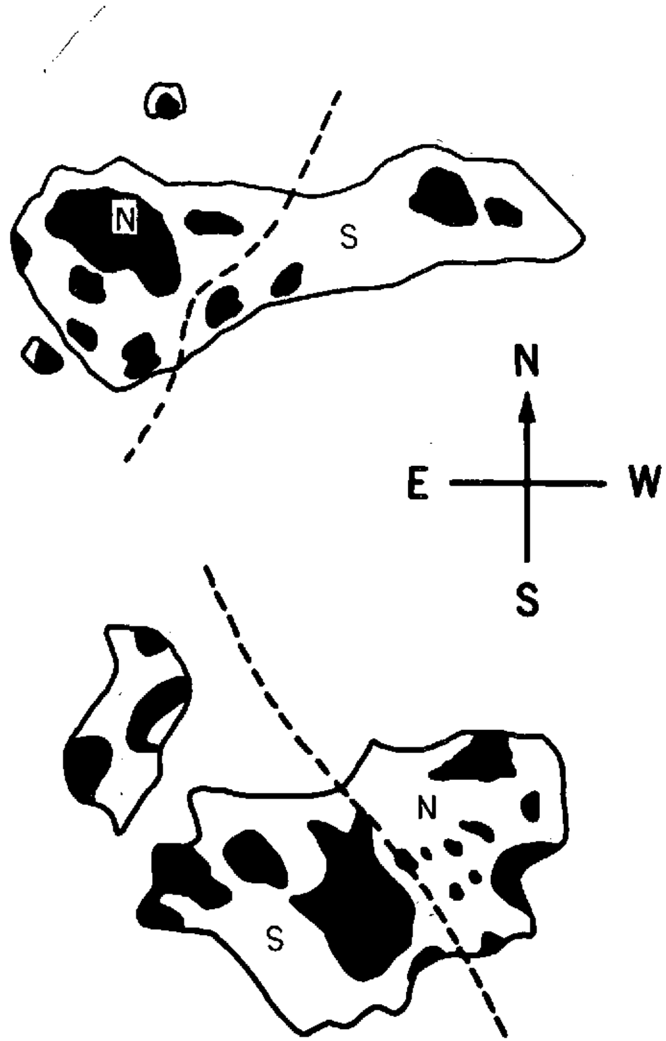


Fig. 3-9

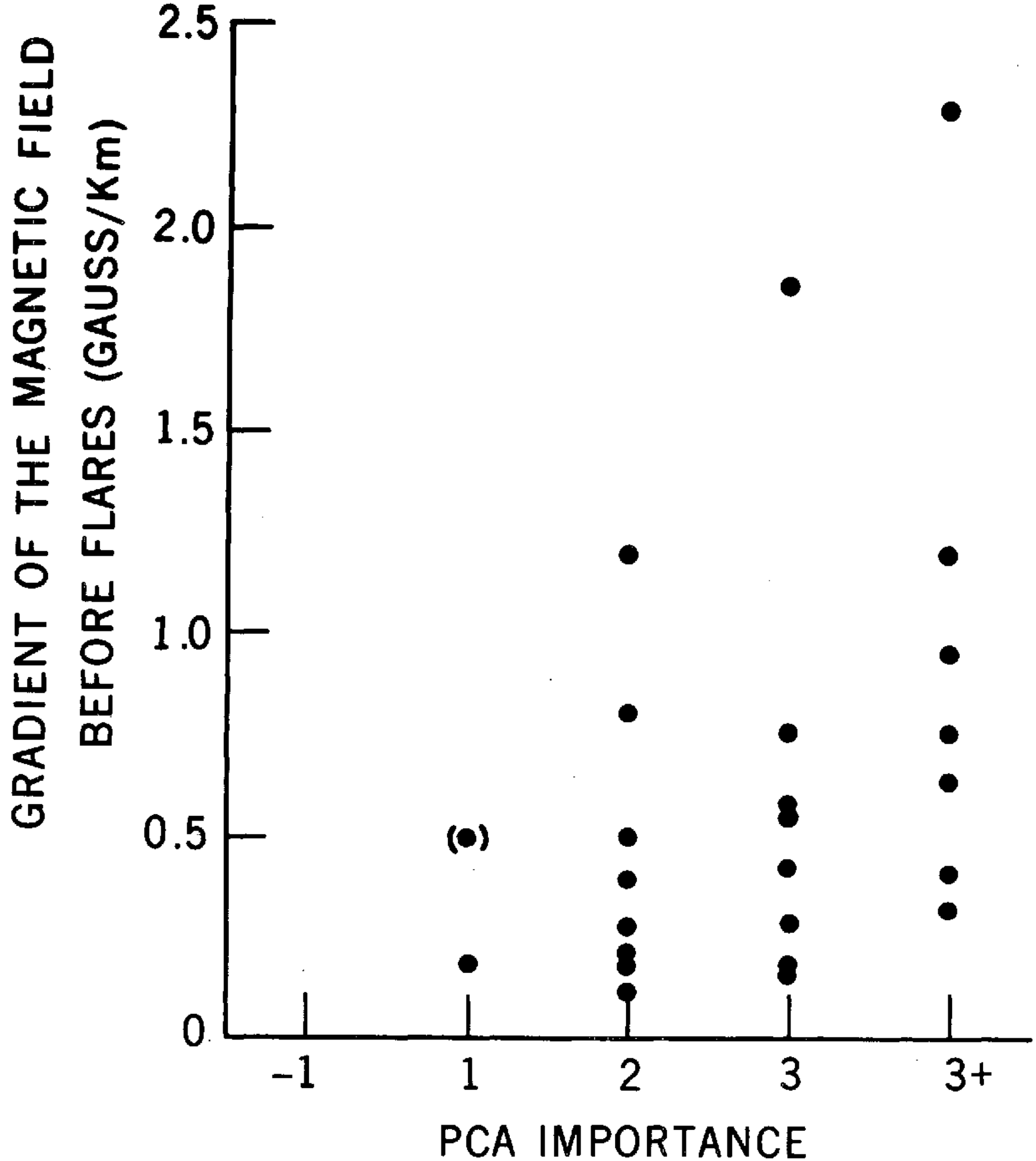
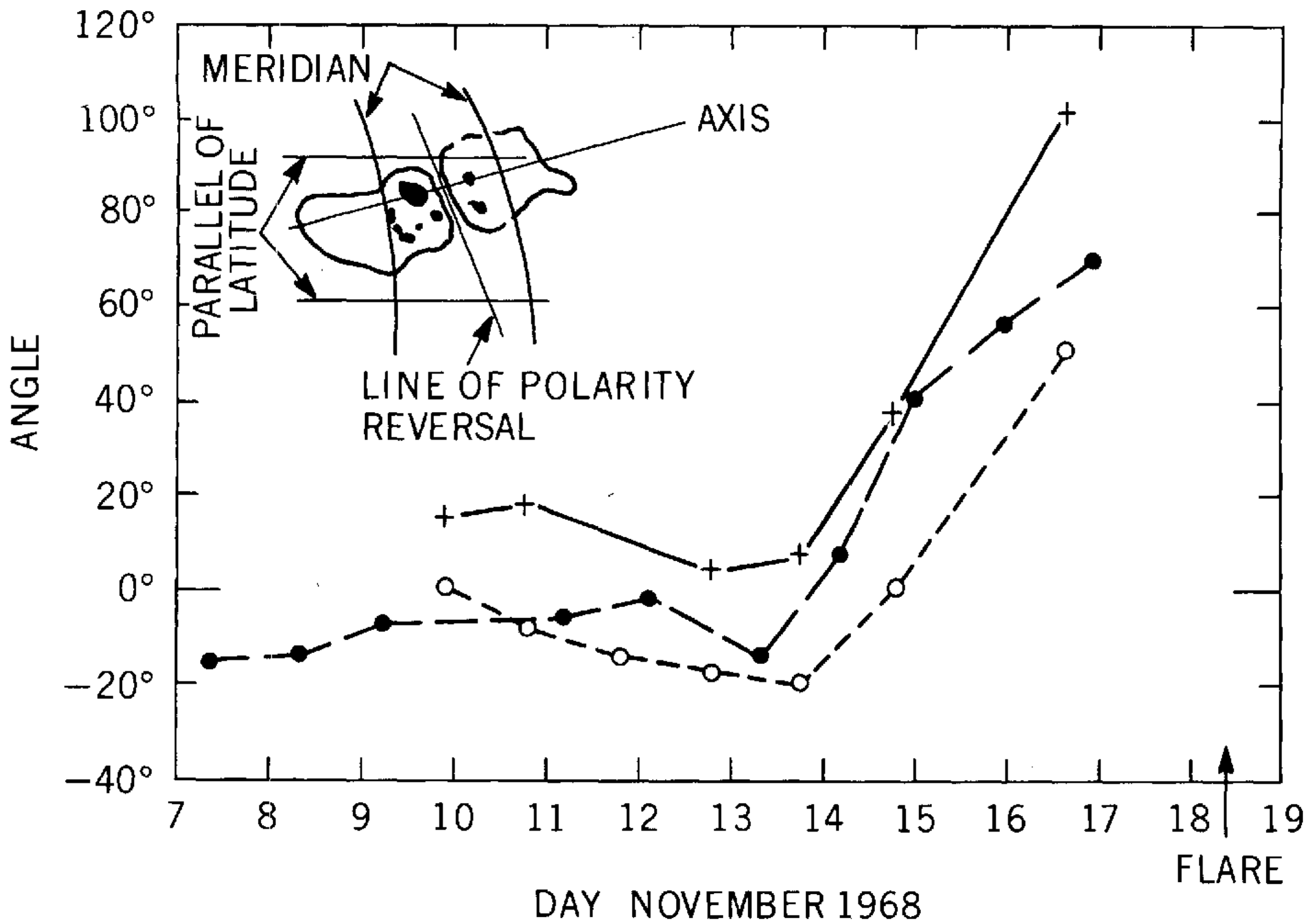

GRADIENT OF THE MAGNETIC FIELD
BEFORE FLARES (GAUSS/Km)
2.5
2.0
1.5
1.0
0.5
0
-1
1
2
3
3+
PCA IMPORTANCE

Fig. 3-10



+——+ ANGLE FROM MERIDIAN TO LIME OF POLARITY REVERSAL, MEASURED ON LOCKHEED FILTER MAGNETOGRAMS.

o- - - -o ANGLE FROM PARALLEL OF LATITUDE TO MAGNETIC AXIS, MEASURED ON LOCKHEED FILTER MAGNETOGRAMS.

●— —● ANGLE FROM PARALLEL OF LATITUDE TO MAGNETIC AXIS, MEASURED ON DRAWING FROM SOLNECHNYE DANNYE

Fig. 3-11

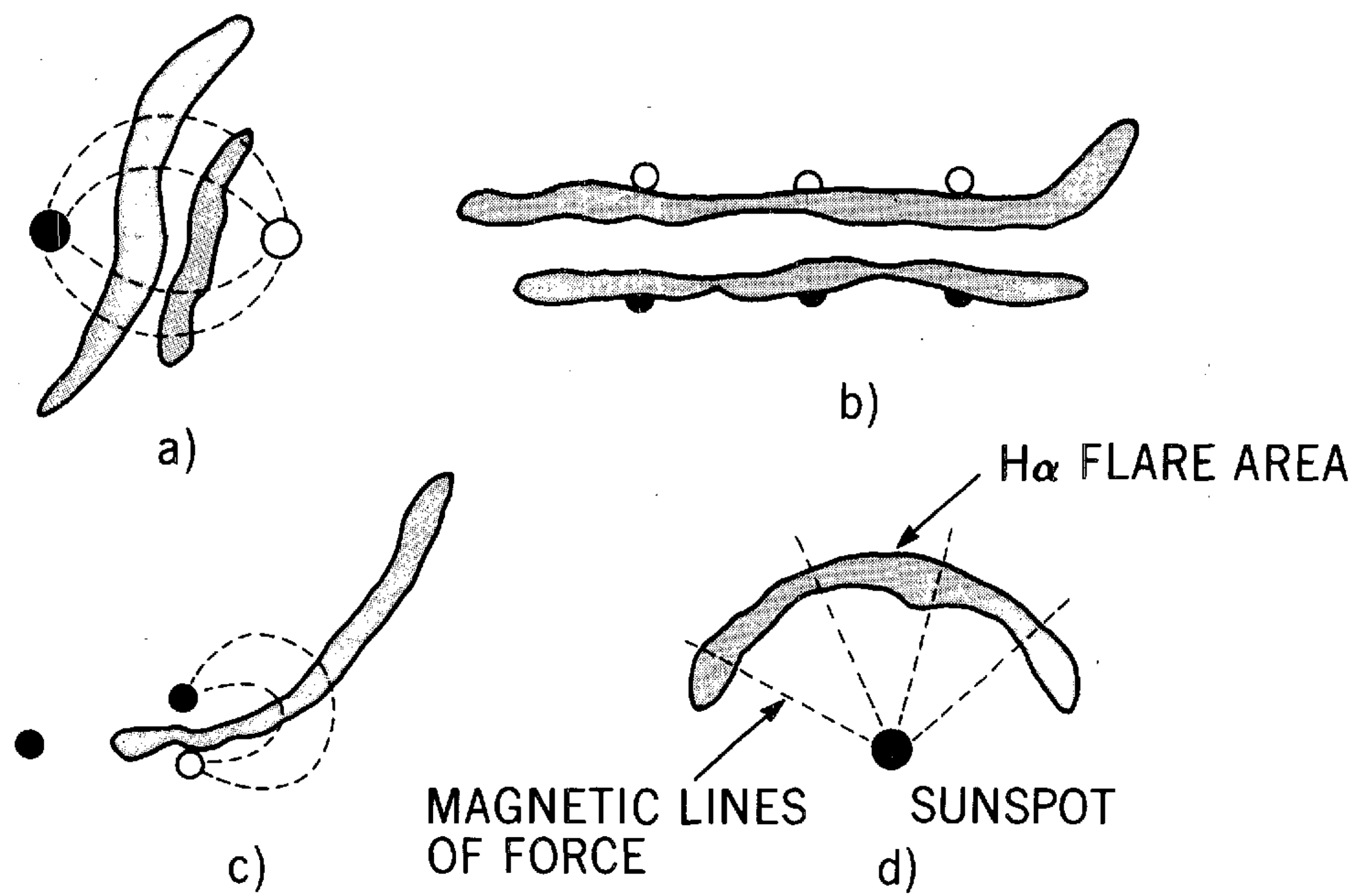


Fig. 3-12

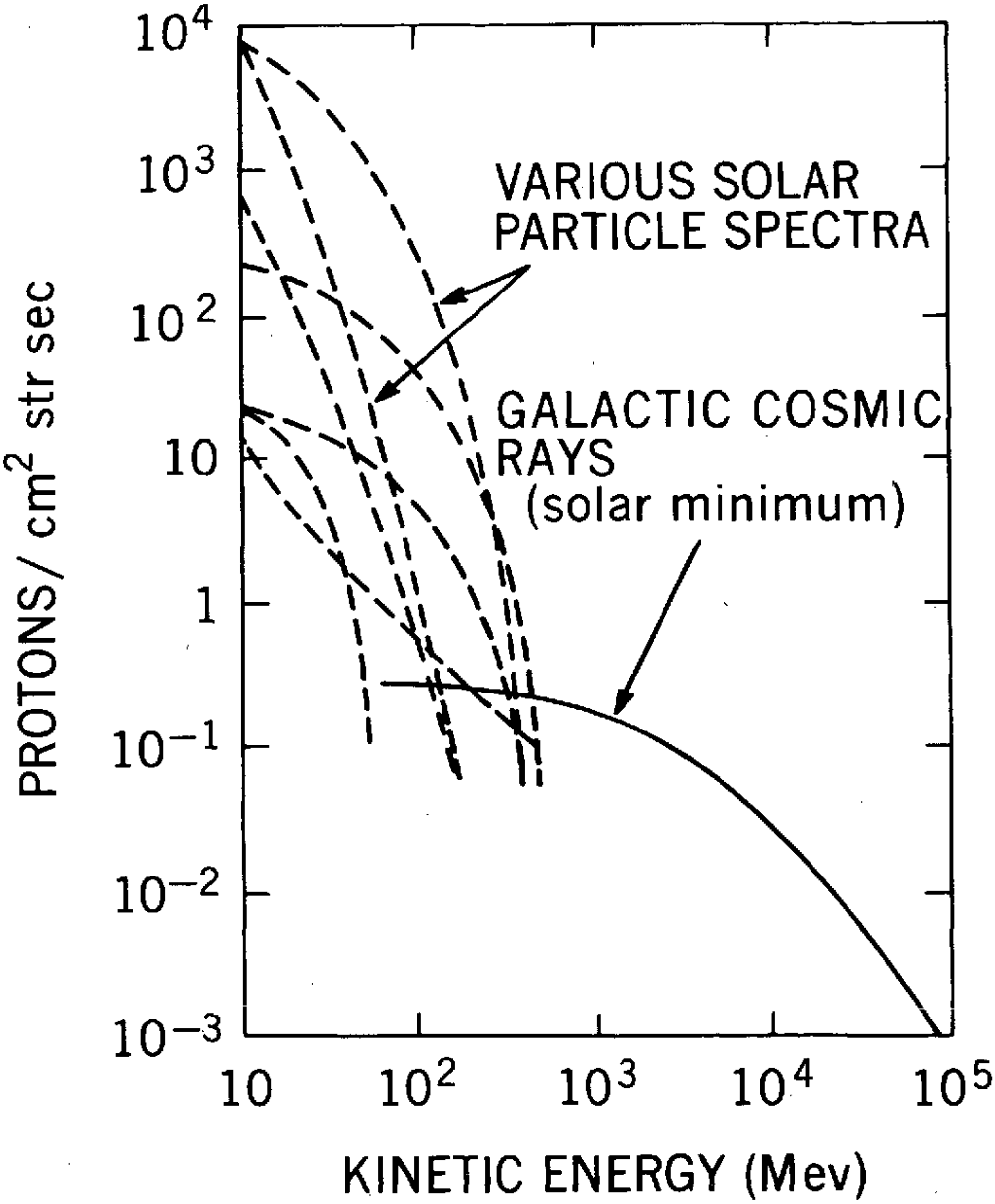
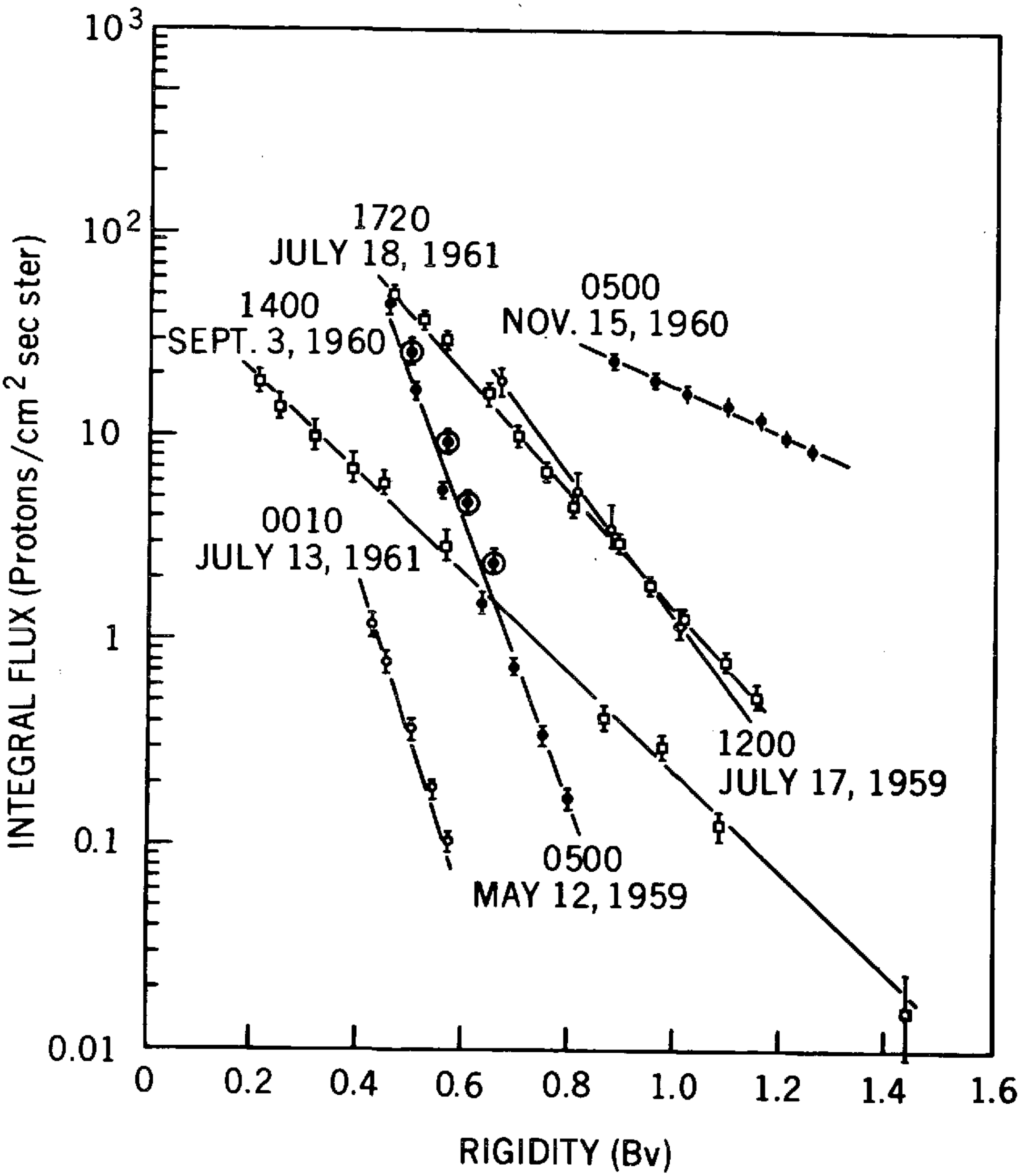

VARIOUS SOLAR PARTICLE SPECTRA
GALACTIC COSMIC RAYS (solar minimum)
PROTONS/ cm2 str sec
10-3
10-2
10-1
1
10
102
103
104
10
102
103
104
105
KINETIC ENERGY (Mev)

Fig. 4-1



EXPONENTIAL SPECTRA FOR SOLAR COSMIC RAYS OBSERVED AT SELECTED TIMES IN DIFFERENT EVENTS.

Fig. 4-2

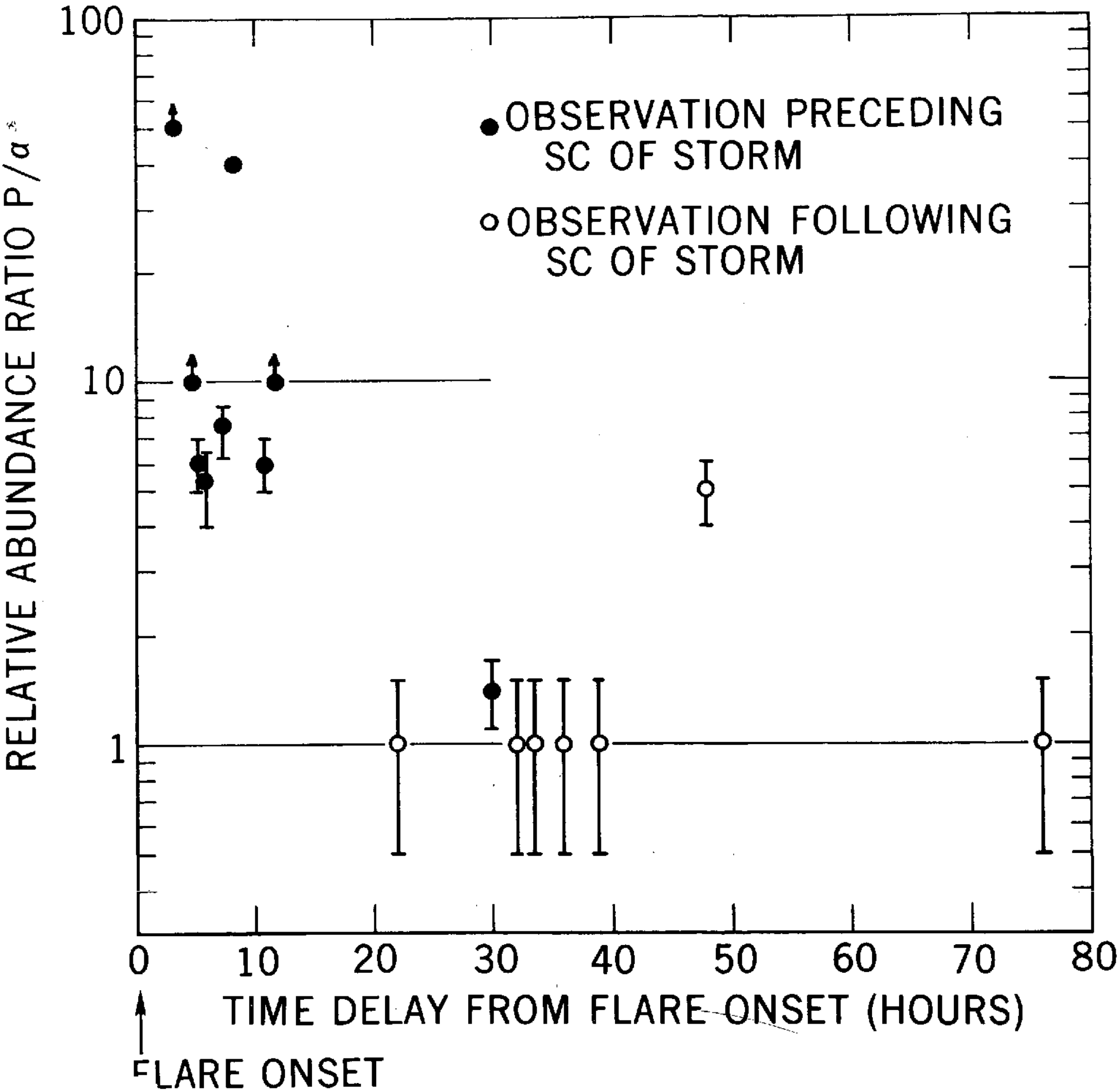

RELATIVE ABUNDANCE RATIO P/α
100
10
1
OBSERVATION PRECEDING SC OF STORM
OBSERVATION FOLLOWING SC OF STORM
0
10
20
30
40
50
60
70
80
TIME DELAY FROM FLARE ONSET (HOURS)
FLARE ONSET

Fig. 4-3

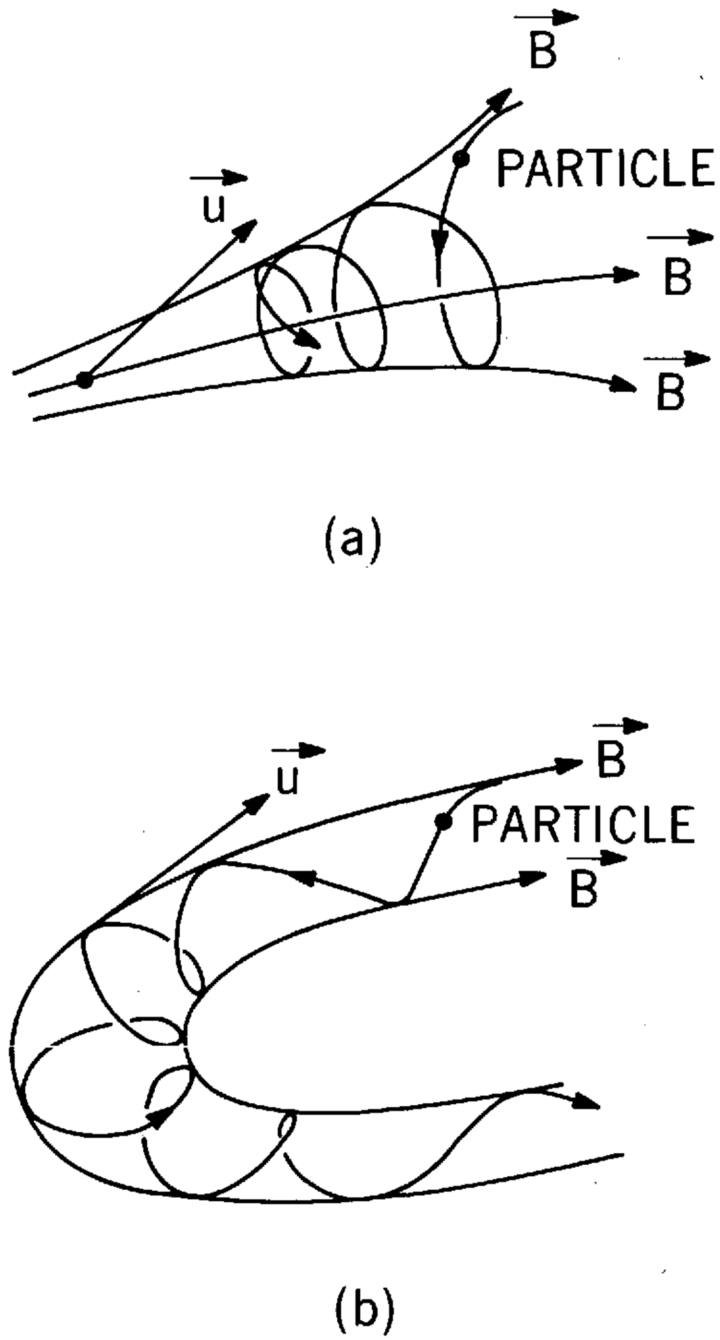


Fig. 5-1

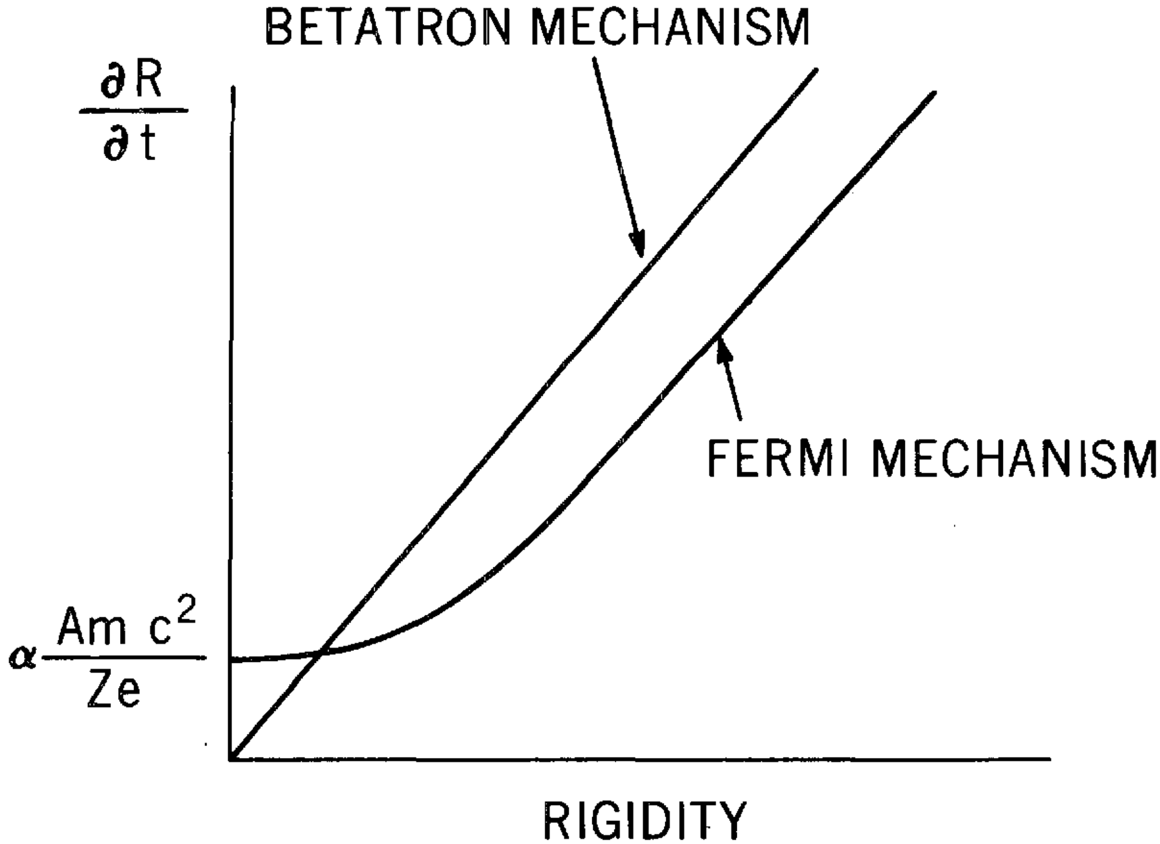
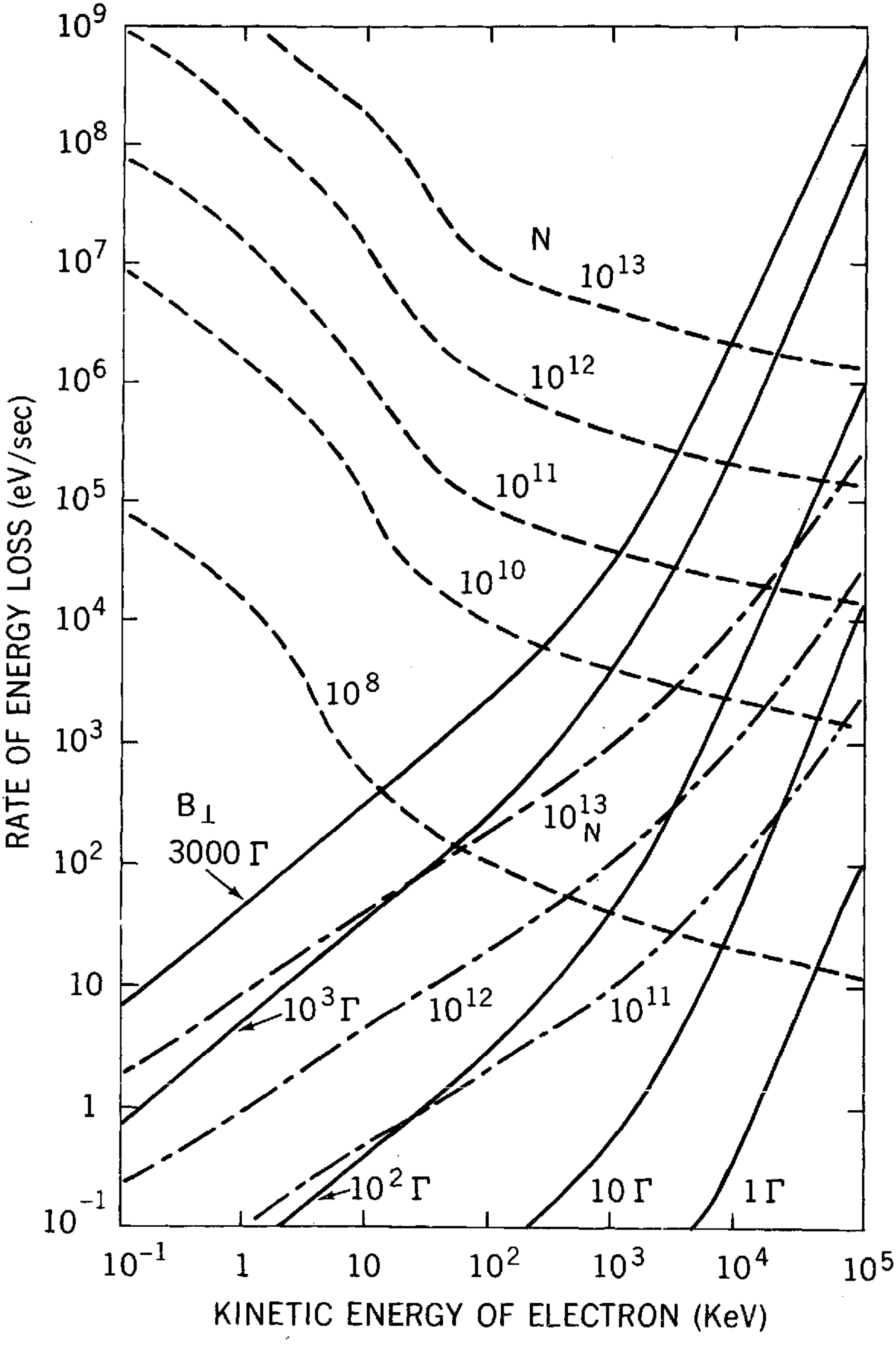

BETATRON MECHANISM
$\frac{\partial R}{\partial t}$
FERMI MECHANISM
$\alpha \frac{Am\,c^2}{Ze}$
RIGIDITY

Fig. 5-2



RATE OF ENERGY LOSS (eV/sec)
KINETIC ENERGY OF ELECTRON (KeV)
N
10^{13}
10^{12}
10^{11}
10^{10}
10^{8}
$B_\perp$
$3000\,\Gamma$
10^{13}_N
$10^3\,\Gamma$
10^{12}
10^{11}
$10^2\,\Gamma$
$10\,\Gamma$
$1\,\Gamma$
10^{9}
10^{8}
10^{7}
10^{6}
10^{5}
10^{4}
10^{3}
10^{2}
10
1
10^{-1}
10^{-1}
1
10
10^{2}
10^{3}
10^{4}
10^{5}

Fig. 5-3