Observations and Theory of Dynamical Triggers for Star Formation

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Abstract. Star formation triggering mechanisms are reviewed, including the direct compression of clouds and globules, the compression and collapse of molecular clouds at the edges of HII regions and supernovae, the expansion and collapse of giant rings and shells in galaxy disks, and the collision and collapse between two clouds. Collapse criteria are given. A comprehensive tabulation of regions where these four types of triggering have been found suggests that dynamical processes sustain and amplify a high fraction of all star formation that begins spontaneously in normal galaxy disks.

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1. Introduction

Star formation requires dense, self-gravitating gas, so compression is often considered to be a requisite, or helpful, precursor. The earliest models of triggered star formation were not concerned with gas density however, but with gas and star motions: Opik (1953) and Oort (1954) proposed that the expansion of OB associations observed by Blaauw & Morgan (1953) was the result of star formation in moving gas, expelled from a supernova or stellar cluster by high pressures. Today, we think that the expansion of OB associations, aside from runaway O stars, is the result of gravitational unbinding of the young embedded stellar cluster after the gas leaves (Zwicky 1953; see review in Lada 1991), so the original motivation for triggering has changed.

The idea that high interstellar pressures could directly squeeze a pre-existing cloud and cause it to collapse to a star followed from the theory of this process by Ebert (1955) and Bonner (1956), and was part of the triggering scenario applied to dark globules by Dibai (1958). Star formation in such globules had previously been considered by Bok & Reilly (1947) and others. This process is now thought to have widespread application to bright rims, cometary globules, globules in HII regions, and isolated clouds subject to high pressures from supernovae and other disturbances (see review in Klein, Whitaker & Sandford 1985).

Two other observations that led to early suggestions of triggering concerned spatial progressions of star formation. For OB associations, Blaauw (1964) observed a spatial separation of subgroups with different ages, and this led to
the idea that star formation is triggered sequentially in a molecular cloud by pressures from HII regions (Elmegreen & Lada 1977) and supernovae (van Till, Loren, & Davis, 1975; Ögelman & Maran 1976; Herbst & Assousa 1977). Larger scale models of triggering (Dopita, Mathewson & Ford 1985) similarly followed from an age progression in the Constellation III region of the Large Magellanic Clouds (Westerlund & Mathewson 1966). (This particular age progression is now in doubt; see Braun et al. 1997). Observations of age progressions are the basis for a second triggering scenario in which relatively low density gas is moved and compressed by a shock front until it collapses gravitationally. The collapsed gas forms dense molecular cores in which star formation occurs at a high rate, possibly forming a dense cluster.

A third suggestion for triggering came from models of gravitational instabilities during shock compression (Woodward 1976) and from the observation of star formation at the interface between an expanding HI shell and a dense cloud (Loren 1976). This led Loren to the cloud-collision model, even though he recognized that cloud collisions could be destructive (Stone 1970). A second example, LkHα, was soon found (Loren 1977), and then Dickel et al. (1978) applied the model to the W75/DR 21 region. Since then, numerous models (e.g., Kimura & Tosa 1996) have shown how collisions can lead to gravitational instabilities in the shocked region, and recent observations of this affect have been reported by Hasegawa et al. (1994) and others.

Evidently, there are three distinct triggering mechanisms that are commonly discussed: (1) direct compression of pre-existing globules or density enhancements in a cloud (“globule squeezing”), (2) accumulation of gas into a dense ridge or shell that collapses gravitationally into dense cores (“collect and collapse”), and (3) cloud collisions.

Another early line of evidence for triggered star formation was the observation of abundance anomalies of short-lived radio isotopes in meteorites (Lee, Papanastassiu, & Wasserburg 1977). Cameron & Truran (1977) suggested this implied the solar system was triggered by a supernova. Recent theoretical studies of such triggering in the globule-squeezing scenario were made by Ramadurai (1986), Boss (1995), Cameron, et al. (1995), and Valhala, Cameron, & Hoflich, (1996). Boss considered slow stellar winds for the triggering, and the others considered supernovae.

These triggering mechanisms may also have applications to star formation in galactic spiral arms. The need for such triggering was first discussed in the context of the density wave theory by Roberts (1969) when it was thought that underlying spiral waves were weak and most of the appearance of spiral structure was from triggered star formation. Today, infrared observations (Elmegreen & Elmegreen 1984; Rix & Rieke 1993; Block et al. 1994) reveal strong arms in even the old stellar disk, and Hα/CO arm-interarm contrasts suggest only moderate amounts of triggering, or, perhaps no triggering in some cases (Garcia-Burillo, Guelin, & Cernicharo 1993; Sempere & Garcia-Burillo 1997). Nevertheless, the globule-squeezing mechanism has been applied to spiral arms by Woodward (1976), the collect-and-collapse mechanism by Elmegreen (1979a, 1994a), Balbus & Cowie (1985), Tomisaka (1987), Balbus (1988), and others, and the cloud-collision mechanism by Kwan & Valdes (1983), Scoville, Sanders, & Clemens (1986) and others.
At the present time, there is no evidence in high resolution dust maps of in M51 (Block et al. 1997) for the comet-shaped, shocked clouds that were predicted by Woodward (1976) to trigger star formation in spiral arms. Indeed, the large interarm clouds tend to be spiral-shaped rather than spherical (Block et al. 1997). There is also no direct evidence for cloud collision-induced star formation in spiral arms, although this would be very difficult to obtain with present-day resolution. Instead, extensive observations of regularly-spaced, giant HI (Elmegreen & Elmegreen 1983, 1987; Boulanger & Viallefond 1992) and CO (Rand & Kulkarni 1990; Rand 1993) clouds in spiral arms, combined with observations of a dominant scale for star formation that has the same size, i.e., that of Gould’s Belt (Efremov 1995), plus observations of a gravity-sensitive criterion for the onset of galactic star formation (Kennicutt 1989), all suggest that the dominant star formation process in spiral galaxies is a spontaneous gravitational instability in the arms and disks. The same appears to be true in some starburst rings, because most star formation takes place in giant hot-spots having a size or separation equal to the local Jeans length (Elmegreen 1994b). Thus spontaneous processes probably dominate the onset of star formation on a galactic scale, but triggered star formation sustains, amplifies, and disperses what large-scale instabilities begin.

Spiral wave and other galactic-scale triggering will not be discussed further here; a comprehensive review is in Elmegreen (1995). A review of all aspects of triggered star formation is in Elmegreen (1992). Here we cover recent observations and some of the most basic aspects of the theory.

2. Observations of triggered star formation

There are numerous observations of star formation in OB associations and other high pressure regions that are thought to be the result of triggering. Tables 1-4 in the next sections list some these observations by region, going back about 10 years. This list was generated by a survey of the literature, with help from the ADS abstract service, so it may miss some references that are not in this service or that do not specifically discuss this topic in the abstract.

There is some attempt to organize the list into the proposed mechanisms of triggering, with a distinction made between small, intermediate, and large scales in the following sense: Small scale triggering (Table 1): direct squeezing of pre-existing clouds or globules by high pressure that nearly surrounds the whole cloud. This includes triggering in bright rims, proplyds, and small cometary globules. Intermediate scale triggering (Table 2): compression of a nearby cloud from one side, leading to a dense ridge of moving gas that presumably collapses or recollects into denser cores in which star clusters eventually form. Large scale triggering (Table 3): accumulation of gas into an expanding shell or ring partially surrounding the pressure source, with star formation in the shell or ring presumably triggered by gravitational collapse of swept-up gas. These three scales follow from a comparison between the size of the triggered region (small, medium and large) and the size of the triggering region.

The observations suggest that triggering occurs in most of the large star-forming regions near the Sun. Indeed, triggering of one type or another may occur inside and around most of the clouds in which OB stars form. For trig-
Figure 1. Schematic diagram of a young cluster (left) interacting with the clumpy structure at the edge of a molecular cloud (right). Some of the clumps are squeezed into gravitational collapse, forming stars.

Triggering on small and intermediate scales, the extent of triggering is probably limited to the pre-existing dense gas, i.e., to the molecular clouds in which the first generation of stars formed. For large scales, triggering apparently occurs in ambient gas that was not previously part of the star-forming cloud. Whether all star formation, including the first generation inside molecular clouds (e.g., by stray supernovae) and the formation of the molecular cloud themselves, has been triggered by a previous generation of stars is unknown. The evidence for galactic-scale gravitational instabilities discussed in the Introduction suggests that only a small fraction of star formation is triggered by a previous generation of stars, perhaps less than 50%. Nevertheless, if a high fraction, e.g., 99%, of all cloud and star formation is triggered, then there will be important implications for spiral structure in galaxies (Seiden & Schulman 1990; Jungwiert & Palous 1994).

3. Theory of triggering

3.1. Small Scale Triggering: Globule-Squeezing

Overview: Numerous observations of clumpy structure inside and on the periphery of molecular clouds (Stutzki et al. 1988; Falgarone, Phillips & Walker 1991) suggest that young HII regions begin their expansion by interacting with cloud clumps. The high pressures of the HII regions can squeeze these clumps and trigger star formation. A schematic diagram is shown in Figure 1 (Figs. 1-4 are from Elmegreen 1992).

Dense neutral clouds at the edges of HII regions may also be formed there, independent of pre-existing molecular cloud clumps. Dynamical instabilities in the swept-up gas between the ionization front and the molecular cloud can fragment the material into small pieces even without self-gravity. One instability occurs if the HII region becomes brighter with time (Elmegreen & Elmegreen 1978a). Another is the Rayleigh-Taylor instability (Schneps, Ho & Barrett 1980,
Schwartz 1985; Capriotti 1996), and still others result from transverse flows in
the thin swept-up layer (Giuliani 1979; Vishniac 1983). Recent applications of
the latter instability to expanding HII regions were made by Garcia-Segura &
Franco (1996).

Numerical simulations of the formation of bright rims and other peripheral
structures are in Bedijn & Tenorio-Tagle (1984), Sandford, Whitaker, & Klein
(1982, 1984), Klein, Sandford, & Whitaker (1983), Lefloch & Lazareff (1994),
and Elmegreen, Kimura, & Tosa (1995). Analytical work on the structure of
embedded globules and cometary clouds is in many references, including Oort &
Spitzer (1955), Kahn (1969), Dyson (1973), Brand (1981), Bertoldi (1989), and
Bertoldi & McKee (1990).

The direct observation of neutral globules in HII regions had a difficult
start. Dyson (1968) proposed they existed in order to power the observed HII
turbulence, but it was not until 10 years later that Laques & Vidal (1978) found
them in Orion. Shaver et al. (1983) then observed clumpy ionized structure in
a large sample of HII regions, and Felli et al. (1984) found bright HII emission
peaks inside the M17 nebula. Further studies of Orion by Garay, Moran, &
Reid (1987) and Churchwell et al. (1987) showed convincingly that some of the
embedded bright rims are from neutral globules.

Table 1. Star Formation in Cometary Globules and Bright Rims

| Region          | Reference                                                                 |
|-----------------|---------------------------------------------------------------------------|
| Gum nebula      | Brand et al. 1983; Reipurth 1983; Sahu et al. 1988; Harju et al. 1990;    |
|                 | Bhatt 1993; Sridharan 1992; Reipurth & Pettersson 1993; Villas-Boas,     |
|                 | Myers; & Fuller 1994; González-Alfonso, Cernicharo, & Radford 1995;      |
|                 | Bourke et al. 1995; Sridharan, Bhatt, & Rajagopal 1996; Schoeller et al.  |
| Orion           | Sugitani et al. 1989; Ramesh 1995; O’Dell et al. 1993, 1994; McCullough et |
|                 | al. 1995; Cernicharo et al. 1992                                       |
| IC 1396         | Kun, Balazs, & Toth 1987; Balazs, & Kun 1989; Nakano et al 1989; Sugitani |
|                 | et al. 1989; Duvert et al. 1990; Serabyn, Güsten, & Mundy 1993; Patel et |
|                 | al. 1995; Weikard et al. 1996; Moriarty-Schieven, Xie, & Patel 1996;     |
|                 | Saraceno et al. 1996                                                    |
| Rosette         | Block 1990; González-Alfonso, & Cernicharo 1994; Patel, Xie, & Goldsmith |
|                 | 1993; Indrani & Sridharan 1994;                                          |
| Ophiuchus       | De Geus 1992                                                             |
| L810            | Neckel & Staude 1990                                                     |
| L1206           | Sugitani et al. 1989; Ressler, & Shure 1991                             |
| L1780           | Toth et al. 1995                                                         |
| L1582           | Zhou, Butner, & Evans 1988                                               |
| IC 1805         | Heyer et al. 1996                                                        |
| IC 1848         | Loren & Wootten 1978; Lefloch & Lazareff 1995                            |
| IC 4628         | King 1987                                                                |
| NGC 281         | Megeath & Wilson 1997                                                    |
| NGC 2264        | Tauber, Lis, & Goldsmith 1993                                            |
| NGC 5367        | White 1993                                                               |
| M8              | Caulet 1997                                                              |
| M16             | Hester et al. 1996                                                       |
| G10.6-0.4, W33  | Ho, Klein, & Haschick 1986                                               |
| G110-13         | Odenwald et al. 1992                                                    |
| Thumbprint glob.| Lehtinen et al. 1995                                                    |
| Carina          | Ogura & Walsh 1992; Megeath et al. 1996                                  |
| BD +40°4124     | Hillenbrand et al. 1995                                                  |
High resolution observations with *Hubble Space Telescope* now find solar-system size globules in Orion (Hester et al. 1991; O’Dell, et al. 1993, 1994), M16 (Hester et al. 1996), and M8 (Caulet 1997). Other similar regions have been found in the Sco-Cen association (Bertoldi & Jenkins 1992). Some of these neutral regions may be protostellar disks, as argued in the papers on Orion, but Hester et al. (1996) claims that many are evaporating globules rather than disks. This follows from the short exposure times of the neutral regions, as determined from their proximity to the rest of the molecular cloud in M16.

Recent evidence for star formation in embedded globules is summarized in Table 1. Young embedded or adjacent stars are connected with these globules, but it is uncertain whether this star formation was triggered by compression or was there before. For example, star formation could have occurred in the dense gas independently of the ionization and simply been exposed when the ionization cleared away the peripheral gas. This situation was discussed in some detail by Elmegreen (1992), and a schematic diagram of it is in Hester et al. (1996).

Dense neutral clouds can also become engulfed by high pressure from more distant sources, such as supernovae or stellar winds, or perhaps from older HII regions that have already expanded. These make comet shapes when the pressure is one-sided. Catalogs of such cometary globules are in Hawarden & Brand (1976), Sandqvist (1976), Reipurth (1983), and Zealey et al. (1983). Observations of star formation in cometary globules are also included in Table 1.

The observations suggest there are three distinct morphologies for dense neutral structures exposed to high pressures: (1) they may be isolated neutral globules that are either protostellar disks or pressurized cloud pieces, (2) they may be cometary or elephant trunk globules, which contain a dense cloud at the brightest end of an elongated structure and a connected neutral tail at the other end, and (3) and they may be bright rims, which are like cometary globules but with very short tails, or no tails, connecting the dense head to the rest of the cloud.

Catalogs of bright-rimmed clouds with embedded star formation are in Sugitani, Fukui, & Ogura (1991), Sugitani & Ogura (1994), and Indrani & Sridharan (1995). Other recent observations are in Table 1. Sugitani, Tamura & Ogura (1995) found small-scale age sequences inside bright rimmed clouds. Similar small scale sequences were found in the Orion ridge by Chandler & Carlstrom (1996).

The interaction between a supernova or other shock and a non-magnetic dense cloud was studied numerically by many authors, including Woodward (1976), Nittmann et al. (1982), Heathcote & Brand (1983), Tenorio-Tagle & Rozyczka (1986), Rozyczka & Tenorio-Tagle (1987), Falle & Giddings (1989), Bedogni & Woodward (1990), Stone & Norman (1992), Klein, McKee, & Colella (1994), and Xu & Stone (1995). Studies with a magnetic field were made by Nittmann (1981) and Mac Low et al. (1994). These tend to show compression of the cloud with the simultaneous formation of a tail of material shred off from the cloud’s surface. If the cloud is initially dense, the compression is strong, and the compressed gas cools so it becomes very dense after the compression, then the interior can collapse to make a star before the surface gets shred away by Kelvin-Helmholtz and other instabilities.
Observations of the interaction between the Cygnus supernova remnant and an ambient cloud are in Fesen et al. (1982, 1992), Graham et al. (1995), and Levinson et al. (1996). Observations of the interaction between the remnant W44 and small embedded clouds are in Rho et al. (1994). These interactions seem much too young for triggering to have happened yet. It is not even clear whether a single supernova exploding in the ambient medium or in the progenitor stars’ wind debris can trigger star formation by itself. Usually supernova triggering takes place in an environment that has already been highly perturbed by a long history of stellar winds and HII regions, such as the Gum nebula or an OB association.

**Theory of Triggering in Globules:** The basic process of triggering is the increased pressure at the edge of a clump. The maximum pressure for a stable, self-gravitating isothermal sphere was given by Ebert (1955) and Bonner (1956):

\[ P_{\text{max}} = \frac{1.4c^8}{G^3 M^2}. \]  

(1)

Château (1987) writes this as

\[ P_{\text{max}} = \frac{xGM^2}{R^4} \quad \text{where} \quad x = \frac{4 - 3\gamma}{8\pi\gamma} \]  

(2)

for equation of state \( P = \rho^\gamma \). When the boundary pressure of the sphere, \( P \), exceeds \( P_{\text{max}} \), the sphere collapses. In the Château condition, the sphere is always stable regardless of \( P \) if \( \gamma > 4/3 \), because then the pressure force that resists the collapse increases faster during compression than the gravitational force that drives it.

This stability at high \( \gamma \) also applies to magnetic clouds, for which there is an effective \( \gamma = 2 \) for compression perpendicular to the field lines without diffusion. The reason for this is that magnetic pressure equals \( B^2/(8\pi) \) and \( B \propto \text{density} \) with perpendicular compression and flux-freezing. Compression perpendicular to \( B \) cannot trigger collapse because the internal magnetic pressure goes up faster than the internal self-gravitational energy density. Thus we expect two conditions for the collapse of both polytropic clouds and isothermal magnetic clouds: one to limit how fast pressure increases with density, and the other to specify how much boundary pressure is needed for self-gravity to overcome internal pressure.

Mouschovias & Spitzer (1976) wrote the collapse condition for an isothermal sphere with an initially uniform magnetic field:

\[ P > P_{\text{max}} = \frac{1.89c^8}{G^3 M^2} \left( 1 - \left[ \frac{M_{\text{mag}}}{M} \right]^{2/3} \right)^{-3}, \]  

(3)

for \( M > M_{\text{mag}} \), where

\[ M_{\text{mag}} = \frac{B^3}{280G^{3/2} \rho^2}. \]  

(4)

This second condition is related to the relative contraction parallel and perpendicular to the field. If \( M << M_{\text{mag}} \), then a significant fraction of the
compression-induced motion is perpendicular to the field lines and the field pressure increases as fast as, or faster than, self-gravity; this gives absolute stability, as discussed above. If \( M > M_{mag} \), then there is significant motion parallel to the field lines during the contraction, driven by self-gravity. In the latter case, the central density goes up faster than in the \( M < M_{mag} \) case, and gravity can overcome internal pressure support when \( P > P_{max} \).

We can invert the \( P > P_{max} \) condition to write

\[
M > M_P = \frac{1.37c^4}{G^{3/2}P^{1/2}(1 - [M_{mag}/M]^{2/3})^{3/2}}.
\]

for specific angular momentum \( j \) parallel to the field.

Torsional magnetic fields decrease the critical mass because they pinch together the poloidal field lines, like an extra gravitational force; this implies that a torsional magnetic wave may trigger collapse in a marginally stable cloud (Tomisaka 1991). Tomisaka et al. (1989) point out that there are two stable solutions for given magnetic flux, mass and angular momentum, corresponding to low-density, low-spin, and pressure-bound, or high-density, high-spin, and gravity-bound. Tohline & Christodoulou (1988) discussed two similar solutions for the non-magnetic case.

Cooling of gas at higher density also decreases stability, in the sense that much milder pressure fluctuations are required to trigger collapse when the adiabatic index \( \gamma \) is small (Tohline, Bodenheimer, & Christodoulou 1987).

The parameters occurring in these equations are boundary pressure \( P \), magnetic field strength, \( B \), density, \( \rho \), mass, \( M \), velocity dispersion, \( c \), and specific angular momentum \( j \). They should all be evaluated before before any self-gravitational contraction occurs, i.e., they are the initial conditions for the globule. In the case of triggering, they are the conditions after the globule is squeezed, but before it collapses. The \( P \), for example, is the pressure in the HII region surrounding the neutral cloud, and \( B \) is the magnetic field strength in the compressed globule, which can be larger than in the pre-compressed molecular cloud.

These equations imply that if a spheroid is stable before compression because

\[
M < M_{mag} = \frac{B^3}{280G^{3/2}\rho^2},
\]

then it will be even more stable after compression perpendicular to the field lines because \( B \propto \rho \) in this case. Conversely, if a stable region of mass \( M < M_{mag} \) is compressed parallel to the field, where \( B \sim \) constant and \( \rho \) increases, then \( M_{mag} \) decreases giving \( M > M_{mag} \), and the condition for instability is essentially the field-free one: \( P > P_{max} \). Thus direct compression triggers instability only if a substantial part of it is parallel to the magnetic field or the field is initially weak.
Figure 2. Schematic diagram of compressions parallel (top) and perpendicular to the magnetic field with the collapse to star formation resulting from gravity overcoming internal magnetic pressure and from magnetic diffusion, respectively.

When there is magnetic diffusion, the situation changes. The diffusion rate is

\[ \tau_{\text{diff}} = \frac{\nu_{in}}{\omega_A} \propto \frac{n_i \sigma_{in} c}{k^2 v_A^2} \propto \frac{n_i \rho R^2}{B^2} \propto \frac{\rho^{3/2} R^2}{B^2} \]  

(8)

for ion-neutral collision rate \( \nu_{in} \), Alfvén wave frequency \( \omega_A \), ion density \( n_i \), ion-neutral collision cross section \( \sigma_{in} \), Alfvén wavenumber \( k \), Alfvén speed \( v_A \), and size of region \( R \); \( n_i \) is assumed to be proportional to \( \rho^{1/2} \) (Elmegreen 1979b). Compared to the free-fall rate, the diffusion rate is

\[ \frac{\tau_{\text{diff}}}{\tau_{\text{ff}}} \propto \left( \frac{\rho R}{B} \right)^2. \]  

(9)

For compression of a fixed mass parallel to \( B \), both \( \rho R \) and \( B \) are constant, so \( \tau_{\text{diff}}/\tau_{\text{ff}} \) is constant and the compression does not speed up the diffusion relative to the collapse. For compression perpendicular to \( B \), \( \rho/B \) is constant and \( R \) decreases, leading to more rapid diffusion.

Thus we have two cases:

1. compression of a stable region parallel to \( B \) may trigger collapse by decreasing \( M_{\text{mag}} \), but it can make magnetic diffusion relatively slow.
2. compression of a stable region perpendicular to \( B \) makes the region dynamically more stable at first, but collapse can follow quickly after the field diffuses away.
It is illustrative to write $\rho$ and $B$ in terms of $P$ using $\alpha = \rho c^2 / P$ and $\beta = B^2 / (8\pi P)$. Then the two conditions for the collapse are

$$M > M_{\text{mag}} = 0.011 M_\odot \frac{c_{0.1}^4 \beta^{3/2}}{P_6^{1/2} \alpha^2} \quad (10)$$

and

$$M > M_P = 0.034 M_\odot \frac{c_{0.1}^4}{P_6^{1/2}} \left( 1 - \left( \frac{M_{\text{mag}}}{M} \right)^{2/3} \right)^{-3/2}. \quad (11)$$

Here, $c_{0.1}$ is the velocity dispersion in units of 0.1 km s$^{-1}$, and $P_6$ is the pressure in units of $10^6 k_B$ for Boltzmann constant $k_B$. As before, the second condition on pressure has been written in terms of a critical mass. If the magnetic and thermal pressures in the globule are comparable to the external pressure, then these two conditions are about the same.

Evidently, compression with a magnetic field is very different from compression without a field. Pressure alone does not trigger dynamical collapse when a magnetic field is present unless the field is weak enough to make $M \gg M_{\text{mag}}$ initially. For compression predominantly parallel to the field, collapse is triggered because $M_{\text{mag}}$ decreases during the compression and gravity overcomes magnetism. For compression predominantly perpendicular to the field, dynamical collapse is inhibited because of the increased field strength, but the field diffuses out of the gas more rapidly. Then enhanced diffusion is what triggers collapse, by lowering $M_{\text{mag}}$ to a value less than $M$. A schematic diagram of these two situations is shown in Figure 2. Of course, compression parallel to the field can be followed by collapse perpendicular to the field, and vice versa, and after both motions, the cloud can be unstable and have rapid diffusion.

The spontaneous formation and collapse of globules inside filaments is another way to form stars. Filaments are very common in the interstellar medium. They may form by magnetic processes (Elmegreen 1994d) or by the gravitational collapse of shocked layers (Miyama, Narita, & Hayashi 1987). A filament whose support in the transverse direction is strongly dominated by magnetic forces will collapse to stable oblate globules, while one with less initial magnetic energy will collapse to unstable globules and eventually stars. The threshold between these two cases occurs for an initial ratio of thermal to magnetic pressure equal to 0.02 (Tomisaka 1995). The fragmentation of a cylinder is strongly influenced by geometry, giving a characteristic separation between globules that scales with the filament width, rather than the Jeans length for sufficiently small Jeans length (Bastien et al. 1991). This is in agreement with the observed regular spacing between globules in dark filaments (Schneider & Elmegreen 1979) and spiral arms (Elmegreen & Elmegreen 1983). Numerical simulations of the collapse of a filament to the stage where dense protostellar disks form are in Nakamura et al. (1995) and Tomisaka (1996). The collapse and fragmentation of prolate clouds was considered by Nelson & Papaloizou (1993) and Bonnell & Bastien (1993).

The stability of interstellar clouds to outside fluctuations in pressure, radiation, and other variables is no doubt more complicated than these simple analyses suggest. Internal motions, turbulence, magnetic waves, fragmentation, heating and cooling, ionization, and other processes can alter the conditions and outcome of triggered star formation. Some of these processes are included in the
many published investigations on cloud collapse. Even without magnetic fields or rotation, there has been a considerable amount of analysis on instability and the collapse phase, as in Larson (1969), Penston (1969), Shu (1977), Hunter (1977), Stahler, Shu, & Taam (1980), Whitworth & Summers (1985), Blottiau, Chieze, & Bouquet (1988), Ori & Piran (1988), Suto & Silk (1988), Foster & Chevalier (1993), Pen (1994), Boily & Lynden-Bell (1995), Tsai & Hsu (1995), and Whitworth et al. (1996). The results of the most recent of these studies indicate that the initial and collapse phases of star-forming clouds cannot be well represented by solutions for a singular isothermal spheres, and that mild kinematic disturbances outside and inside the initial cloud can have important consequence for cloud evolution and collapse. This means that the trigger for some collapses can be very subtle.

Collapse studies including rotation usually investigate binary or multiple star formation; they are rarely concerned with triggering. Nevertheless, a few examples are useful for comparison, such as Boss & Black (1982), Terebey, Shu, & Cassen (1984), Myhill & Kaula (1992), Bonnell & Bastien (1992), Burkert & Bodenheimer (1993), Boss (1993), Sigalotti & Klapp (1994), Boss, & Myhill (1995), and Boss (1996). It is not clear whether the multiplicity of stars is affected by triggering. If triggering by one-sided compression tends to make flattened clouds, then multiplicity may be increased compared to the formation of stars in centrally-condensed, spherical clouds (Chapman et al. 1992).

The collapse of magnetic clouds, with or without rotation, also has unknown implications for triggering. Usually the collapse is taken to begin from a near-static initial condition, as in the Mouschovias & Spitzer problem reviewed above. If the collapse is initiated by an implosion, or a torsional magnetic wave, or some other fluctuation in the environment, then the conditions for star formation could change. Variations in ionizing radiation have been considered by McKee (1989), for example. The collapse of magnetic clouds, usually following ambipolar diffusion, has been considered recently by Galli & Shu (1993a,b), Fiedler & Mouschovias (1993), Basu & Mouschovias (1994, 1995a,b), Ciolek & Mouschovias (1994, 1995), and Boss (1997).

3.2. Intermediate scale triggering: Collect and Collapse

As an HII region ages, its expansion moves more and more of the neutral cloud, accumulating it into a dense neutral ridge at the nebula/cloud interface. This ridge may be unstable on a short, internal-crossing timescale because of kinematic, magnetic, and gravitational processes (Elmegreen & Elmegreen 1978b; Giuliani 1979; Doroshkevich 1980; Welter & Schmid-Burgk 1981; Welter 1982; Vishniac 1983, 1994; Bertshinger 1986; Vishniac & Ryu 1989; Wardle 1991; Nishi 1992; Yoshida & Habe 1992; Kimura & Tosa 1991, 1993; Lubow & Pringle 1993; Mac Low & Norman 1993; Strickland & Blondin 1995; Garcia-Segura & Franco 1996), and this fast instability may be oscillatory or monotonic, depending on the deceleration.

However, because the layer is thin in one dimension, it is also unstable on a longer timescale to gravitational collapse along its length. Gravity always acts to make the geometry of a cloud more spherical if it is pressure supported in one or two dimensions but not the third. The gravitational time scale is generally longer than the kinematic in this situation, so the large-scale collapse can occur
later, in a chaotic region that may have previously experienced several types of smaller scale instabilities. If the kinematic instabilities produce fine-scale structure, then the gravitational collapse pulls this structure together, forming large turbulent cores. A schematic diagram of this scenario is shown in Figure 3. Since the ridge is always unstable to collapse along its length, the condition for triggered cluster formation is the time when the large-scale collapse becomes important, considering its growth and motion. This is approximately the time when the gravitational collapse rate equals the inverse of the layer age.

Doroshkevich (1980) and Vishniac (1983) gave the instantaneous gravitational collapse time for a constant-velocity plane-parallel layer bounded by a shock on one side:

$$t \sim (2\pi G \rho_{\text{layer}})^{-1/2} \sim \left(2\pi G \rho_0 M^2 \right)^{-1/2}$$  \hspace{1cm} (12)

for preshock density $\rho_0$ and $M = v_{\text{shock}}/c_{\text{layer}}$. This result came from the collapse rate $\omega$ where

$$\omega^2 = 2\pi G \sigma k - k^2 c^2;$$  \hspace{1cm} (13)

$\sigma$ is the mass/area in the layer and $k = 2\pi/\lambda$ is the wavenumber for wavelength $\lambda$. For a steadily moving layer, we set the instantaneous collapse rate equal to the inverse of the time, and we set the net accumulated column density equal to $\sigma = \rho_0 vt$, to get the overall collapse time of the fastest-growing wavelength:

$$t \sim (\pi G \rho_0 M)^{-1/2}.$$  \hspace{1cm} (14)

For a decelerating layer, the growing dense core migrates to the shock front because it decelerates less rapidly than the lower density parts. At the shock front it can $\textit{erode}$ because of the transverse flow driven by the resulting shock curvature. Erosion competes with growth by self-gravity so the overall collapse can be $\textit{delayed}$ to the longer time (Elmegreen 1989; Nishi 1992):

$$t \sim 0.5 (G \rho_0)^{-1/2}.$$  \hspace{1cm} (15)
This solution again came from a collapse rate $\omega \sim 1/t$, but now

$$\omega^2 \sim 2\pi G\sigma k - 0.5k^2v_{\text{shock}}^2.$$  \hfill (16)

Note the substitution here of $v_{\text{shock}}$ for $c$ in equation (13) (Elmegreen 1989). Other processes leading to internal shear may delay the collapse to this timescale too (Doroshkevich 1980). Equation (16) for $\omega$ indicates that collapse occurs only when the layer is so strongly self-gravitating that forced transverse motions behind the shock front cannot disrupt the density perturbations.

For a decelerating layer, collapse on the short timescale $(\pi G\rho_0 M)^{-1/2}$ for $M >> 1$ leads to small clumps that may not be self-gravitating (Lubow & Pringle 1993), but if cooling is sufficient and self-gravitating clumps do form, then any stars they produce will move out of the front of the layer and leave. This is because the deceleration of the layer, $\sim v_{\text{shock}}/t$, always exceeds the self-gravitational acceleration perpendicular to the layer, $G\sigma$, for timescales $t << (G\rho_0)^{1/2}$ (Elmegreen 1989; Nishi 1992). To see this, we set $\sigma \sim v_{\text{shock}}\rho_0 t$, then $v_{\text{shock}}/t >> G\sigma$ whenever $t << (G\rho_0)^{1/2}$. The observation of giant condensations inside and gravitationally bound to swept-up layers (Table 2), and of the ages and dimensions of these layers for molecular cloud densities exceeding $10^3$ cm$^{-3}$, are consistent with the long time scale of $\sim 0.5(G\rho_0)^{-1/2}$ for triggered clusters in swept-up gas. This implies that gravitational rather than kinematic instabilities govern the onset of cluster formation in a shocked layer. Any observation of moving young stars ahead of the swept-up layer, inside the unshocked molecular cloud, would be evidence for triggering on the short time scale.

The analyses of layer instabilities by Voit (1988) and Whitworth et al. (1994a,b) treated primarily the first significant collapse of the layer, on the short timescale, $(G\rho_0 M)^{-1/2}$, when the layer is still primarily pressure-bound. They did not consider the forward migration and erosion of condensations, shear behind the shock, and the possibility that stars migrate individually out of the front of the layer if the layer decelerates. Nevertheless, their assumptions are good for non-decelerating layers, as might occur between two colliding clouds, and this was one of the applications considered by Whitworth et al..

Lubow & Pringle (1993) also considered the collapse of a highly compressed layer, and explained analytically the result found in Elmegreen & Elmegreen
(1978b), that the layer has an early gravitational instability with a wavelength comparable to the layer thickness and a growth rate comparable to the internal gravitational timescale, $(G\rho)^{-1/2}$ for internal layer density $\rho$. This is true even though the layer is confined more by pressure than self-gravity in the perpendicular direction. The reason they gave is that the initial fast collapse is an incompressible deformation of the surface under these high-pressure conditions, and such deformations can happen quickly. The deformation is curved symmetrically about the center without deceleration, and so the boundary pressure is directed inward, aiding the self-gravity (see also discussion in Elmegreen 1989). Compressive gravitational instabilities on a larger scale, which may give rise to cluster formation, grow later. Lubow & Pringle (1993) considered applications to cloud collisions and concluded that substantial cooling is required to trigger the fast collapse into stars (as discussed also by Hunter et al. 1986).

Another discussion of the collapse of decelerating layers was by Nishi (1992), who explained the linearized results obtained numerically by Elmegreen (1989) in an analytical fashion, considering appropriate simplifying assumptions. Nishi (1992) found the short oscillating pressure-driven solutions and the long, monotonic, gravitationally-driven solutions, and pointed out the importance of a dimensionless parameter ($\alpha$), which is essentially the ratio of the Jeans length to the scale height inside the layer, or, similarly, the ratio of the deceleration of the layer to the self-gravitational acceleration perpendicular to the layer. A layer begins its evolution with a large value of $\alpha$ and ends with a small value as self-gravity becomes important. Nishi (1992) found that the monotonic gravitational collapse begins when $\alpha \sim 1$, and this corresponds to the long overall timescale, $\sim (G\rho_0)^{-1/2}$ for external (pre-shock) density $\rho_0$, as given by equation (15) above.

A 2D hydrodynamic simulation of the collapse of a decelerating isothermal layer was done by Yoshida & Habe (1992). They confirmed the oscillatory behavior at small wavelength and the monotonic collapse at large wavelength that was found by others, they got the predicted migration of the dense condensation to the front of the layer, and they got the expected timescale for gravitational collapse, i.e., comparable to about $0.5 \times$ the free fall time in the unshocked gas. They found several new effects however. At moderately long wavelengths where gravity begins to be important, the tangential flow behind the shock front that was found by Vishniac (1983) and others to cancel the pinch force from pressure there, was compensated in the 2D calculation by an opposing tangential force far behind the shock front in the dense gas, driven by this pinch force and by self-gravity. Because of this, a dense condensation could form at a leading perturbation and not be eroded away, as found by Elmegreen (1989) in the linear analysis, and so the condensation stayed at the leading perturbation and grew to significant densities, always with a slight protrusion out of the front. After a time equal to half the external free fall time, the condensation had an escape velocity larger than the transverse flow speed, so the erosion never got to be effective. At longer wavelengths, gravity is even more important and the collapse, still monotonic and unperturbed by erosion, occurs without much migration to the front of the layer. This is because the growth at longer wavelengths is slow enough that when it becomes important, the layer has slowed down so much that its deceleration is significantly less than the perpendicular gravitational acceleration toward the midplane. They also obtained a minimum wavelength
for gravitational collapse equal to about the layer Jeans length, $2\pi G\sigma/c^2$, for mass column density $\sigma$ and velocity dispersion $c$, evaluated at the time when the deceleration of the layer equals the self-gravitational acceleration perpendicular to the layer ($\alpha = 1$ in the notation of Nishi 1992). Under typical conditions, the mass contained in this minimum wavelength is several hundred solar masses or more, suggesting again the formation of embedded clusters rather than individual stars. These results were for an initial corrugation of fixed size; the results for other perturbations and for colder postshock gas are not known. The results with magnetic fields are also not known from any of these analyses.

It is interesting that the short and long timescale solutions for layer collapse were discussed in the Russian literature before these ideas surfaced in the West. Doroshkevich (1980) used the same shock boundary condition as Vishniac (1983) and got oscillatory solutions for unstable growth on the short time scale, as Vishniac and others did (see equations 7a and 10 in Doroshkevich), and he included shear flow behind the shock and got the long time scale for gravitational collapse as in Elmegreen (1989) for the analogous symmetric mode (i.e. collapse along the layer; see equations 19 or 26 in Doroshkevich and note that his $\beta$ is our $M$). Doroshkevich (1980) summarized his results with the prescient statement: "Shear flows therefore sharply diminish the maximum growth rate of the symmetric perturbation mode. The wavelength corresponding to the maximum growth rate will increase very rapidly. At the same time, however, a hydrodynamic instability will set in, developing within a comparatively short time interval on small scales and causing the layer to break up into small clouds. Later the clouds might clump together through the action of gravitational instability that has developed on scales of long wavelength. These results depend only weakly on the choice of velocity profile (provided the profile is reasonably smooth)." This is pretty close to what we think today. Doroshkevich (1980) applied his results to cosmological pancakes; here they are relevant to the formation of star clusters in compressed layers at the edges of H II regions and elsewhere.

What is the actual mechanism of star formation in this scenario? All that the instability analysis describes is the collapse of a layer into one or more dense cores. In fact, the collapse is probably first to filaments, and then to cores inside the filaments (Miyama, Narita, & Hayashi 1987). Presumably the mechanism of star formation is the same as in any other dense core once it forms, but faster in this case because of the higher density following the compression. This higher density can be quite large for initial compression parallel to the magnetic field, because it increases not only from the direct one-dimensional compression caused by the H II region, but also because of the subsequent collapse of the compressed gas perpendicular to the field. A schematic diagram of this two-step process is shown in Figure 4. The core density divided by the preshock density equals approximately the cube of the average density compression factor in the layer alone (Elmegreen 1985, 1992). Most of this density increase is from self-gravity; only a small amount is from the external pressure.

In Figure 4, the initial cloud width is $W_0$ perpendicular to the field, the length is $L_0$ along the field, the density is $\rho_0$, and the magnetic field strength $B_0$. After the compression (subscript 1), the width and field strength are the
Figure 4. Diagram of the collapse of a layer that moves parallel to the mean magnetic field orientation. The initial values are denoted by subscript 0; values after a purely lateral compression are denoted by 1, and after the gravitational collapse by 2. The result is written in terms of $L_1/L_2 \approx 1$. After the collapse, the density in the cloud has increased by the cube of the lateral compression factor $C$.

same, $W_1 = W_0$, $B_1 = B_0$, the density is higher by the shock-compression factor $C$, $\rho_1 = C\rho_0$, and the length is smaller by the inverse of this factor, $L_1 = L_0/C$.

The final configuration (subscript 2) considers that both the initial and final clouds are in equilibrium with the magnetic field resisting self-gravity. This requires magnetic field strengths proportional to the transverse column densities:

$$B_0 = K\rho_0 W_0 \quad ; \quad B_2 = K\rho_2 W_2.$$  \hspace{1cm} (17)

Then with mass and magnetic flux conservation, the final density becomes

$$\frac{\rho_2}{\rho_0} = \frac{1}{N} \left( \frac{CL_1}{L_2} \right)^3$$  \hspace{1cm} (18)

for $N$ cores. We expect $L_2 < L_1$ because of the extra gravity in the collapsed core, and typically $N$ for the formation of OB subgroups is in the range 1-5; then the compression factor after the collapse can be comparable to $C^3$, which may be a factor of 100 to 1000.

In the collect-and-collapse scenario, several dense fragments form in a ridge of swept-up gas because of gravitational instabilities. Stars form in these fragments at a much higher rate than they would have formed in the same gas without the compression because of the higher fragment density. The time scale for the whole triggering process is about the time scale for the layer to collapse, as given above by equation (15), because the star formation time inside each fragment should be relatively short compared to this, such as $(G\rho_2)^{-1/2}$.

The important aspect of this scenario is that there is a delay between the expansion of the HII region and the onset of triggered star formation in the swept-up layer. There is also bulk motion of the compressed gas and triggered clusters. The delay results from the layer’s initial stability against collapse that comes from the shock-forced transverse motions, erosion, and turbulence.
is evident from the term with $v_{\text{shock}}$ in equation [16]. A similar point was made by Vishniac (1983), but here we stress that the gravitational instability should not be prevented by the internal motions, but only delayed. As a result, all of the stars in a triggered cluster at the edge of an OB association should be significantly younger than the stars in the shock-driving cluster. That is, star formation should not be continuous in the layer, beginning with the onset of the expansion, but delayed. Then, when it begins, it happens rapidly, so all the triggered stars have about the same young age. Also because of this, there is not a smooth distribution of intermediate age stars between the old cluster and the young cluster, but essentially no intermediate age stars between the two clusters. These two aspects of the collect-and-collapse scenario are in agreement with observations of OB subgroups (Blaauw 1964, 1991).

In contrast, triggering in the globule-squeezing scenario is immediate because the pre-existing clumps are compressed quickly. Such triggering also takes place throughout the HII region because residual cloud pieces can be anywhere. Globule compression differs also because it probably forms individual stars or small stellar systems, while the collect-and-collapse scenario forms whole clusters if the compressed layer contains enough mass.

Other processes that can structure dense molecular gas and trigger star formation are likely to occur as well. For example, the forced motion of the ridge can cause pre-existing clumps in the moving part of the cloud to collide and coalesce with pre-existing clumps in the unshocked part of the cloud. Such collisions can lead to star formation on a clump-by-clump basis (Greaves & White 1991). In addition, the pre-existing clumps cause irregularities in the shock front, and these lead to transverse motions and the accumulation of large cloud fragments in the ridge, even without self-gravity (Kimura & Tosa 1993). The resulting fragments have the appearance of bright rims, and may contain star formation at enhanced rates when gravity becomes important (Elmegreen, Kimura, & Tosa 1995).

### 3.3. Large Scale Triggering: Shells and Rings

Expansion around a centralized pressure source leads to a shell or ring that can become gravitationally unstable to form clouds and new stars along the periphery. The actual physical size of the shell can be large or small, depending on the pressure and external density. If the external density is high, then the shell and propagation distance will be small. For expansion into the ambient medium, which generally has a low density, the shell can be several hundred parsecs in size. Examples of giant shells with young clusters along the periphery are found in our Galaxy and neighboring galaxies, such as M31 and the LMC, as summarized in Table 3. A review of giant shell formation and triggered star formation is in Tenorio-Tagle & Bodenheimer (1988).

After Heiles’ (1979) discovery of giant HI shells in our Galaxy, the first detailed model for their origin around high pressure OB associations was by Bruhweiler et al. (1980). Other models for shell formation soon followed, including one in which an extragalactic cloud collides with the galaxy disk and the resulting shell fragments into giant molecular clouds (Tenorio-Tagle 1981). Propagation of star formation over such distances had already been assumed by Gerola & Seiden (1978). Applications of the shell-collapse scenario to the Orion,
Perseus and Sco-Cen clouds in the Lindblad ring were then made independently by Olano (1982) and Elmegreen (1982), with the first detailed summary of observed regions in Elmegreen (1985).

Analytical work on gravitational instabilities in expanding shells began with Ostriker & Cowie (1981) and Vishniac (1983). An early computer simulation for the collapse of the Lindblad ring was in Elmegreen (1983, 1985). Consideration of both the theory for shell expansion and the theory for collapse was in McCray & Kafatos (1987), who used a model of shell formation in which the energy continuously increased, as if by a steady wind or continuous supernovae. Differential rotation was introduced by Tenorio-Tagle & Palous (1987) and Palous, Franco, & Tenorio-Tagle (1990).

Table 3. Triggered Star Formation in Swept-up Shells or Rings

| Region                  | Reference                                                                 |
|-------------------------|---------------------------------------------------------------------------|
| **Galactic Clouds**     |                                                                           |
| Lindblad’s Ring         | Elmegreen 1982; Olano 1982; Taylor, Dickman, & Scoville 1987; Franco et al. 1988; Comerón, & Torra 1994b |
| W3/W4                   | Lada et al. 1978; Dickel et al. 1980; Thronson, Lada, & Hewagama 1985; Routledge et al. 1991; Digel et al. 1996; Tieftrunk et al. 1997 |
| W5                      | Sato 1990                                                                 |
| IC 443, W28, W44, S147, HB21 | Odenwald & Shivanandan 1985                                               |
| Ara OB1 field           | Rizzo & Bajaja 1994                                                       |
| W75                     | Ward-Thompson & Robson 1991                                               |
| Cygnus superbubble      | Comerón & Torra 1994a                                                     |
| MonR2                   | Hughes & Baines 1985; Xie & Goldsmith 1994                                |
| NGC 1333                | Langer, Castets, & Lefloch 1996; Warin et al. 1996                        |
| GS235-02                | Jung, Koo, & Kang 1996                                                   |
| G214.6+0.0              | Handa et al. 1986                                                        |
| **Magellanic Clouds**   |                                                                           |
| Constellation III       | Dopita, Mathewson & Ford 1985                                            |
| LMC2                    | Wang & Helfand 1991                                                       |
| LMC4                    | Domgörgen, Bomans & De Boer 1995; Olsen et al. 1997                      |
| N11                     | Walborn & Parker 1992; Parker et al. 1992; Rosado et al. 1996            |
| DEM 152 in N44           | Oey & Massey 1995                                                        |
| NGC 2214                | Bhatt & Sagar 1992                                                        |
| **Galaxies**            |                                                                           |
| M33                     | Deul & den Hartog 1990; Palous 1991                                       |
| M31                     | Brinks & Bajaja 1986; Brinks, Braun, & Unger 1990                        |
| IC 2575                 | Martimbeau, Carignan, & Roy 1994                                         |
| NGC 1313                | Ryder et al. 1995                                                        |
| Ho II                   | Puche et al. 1992                                                        |
| BCD SBS 0335-052        | Thuan, Izotov & Lipovetsky 1997                                          |
| NGC 1620                | Vader, & Chaboyer 1995                                                    |
| Giant Extragalactic HII regions | Mayya & Prabhu 1996                                                      |

Applications of long-range propagating star formation to galactic structure and spiral arm formation were made by many others following Mueller & Arnett (1976) and Gerola & Seiden (1978). A recent review of their work is in Seiden & Schulman (1990), and other recent work is in Korchagin & Riabtsev (1992), Newkirch & Hesse (1993), Jungwiert & Palous (1994), and Palous, Tenorio-Tagle, & Franco (1994). These latter papers consider asymmetric propagation, as might result in the presence of galactic shear.
The criterion for star formation in an expanding shell or ring is analogous to that in the collect-and-collapse scenario discussed in the previous section. Shells and rings behind shock fronts contain a variety of instabilities, driven by kinematic, magnetic, and self-gravitational processes. Some of these stir up the gas and create small scale turbulence and structure not related to star formation. Gravitational instabilities lead to the formation of large condensations inside the swept-up material, and some of these may produce embedded clusters. Thus the criterion for the onset of cluster formation in a shell or ring is the time at which gravitational instabilities become important.

The divergence of expanding shells or rings gives them an initial stability against self-gravitational collapse. For continuous energy deposition into a shell, so that \( v_{\text{shock}} \propto t^{-0.4} \) (Castor et al. 1975), the shell is first unstable when (Elmegreen 1994c)

\[
t \sim (G\rho_0 M)^{-1/2},
\]

and the growth rate equals \( 1/t \) when

\[
t \sim 1.25(G\rho_0 M)^{-1/2}.
\]

In these equations, the preshock density is \( \rho_0 \) and the ratio of shock speed to velocity dispersion in the swept-up gas is \( M = v_{\text{shock}}/c \). Slightly different results were obtained by Whitworth et al. (1994a) without the effects of shell divergence. Theis et al. (1997) showed how the ratio of the fragmentation time to the time of first instability is constant, and they also showed that any density gradient in the surrounding medium has to be shallower than isothermal for the self-gravity of the accumulated matter to ever overcome the stability given by the expansion.

For a ring, the growth rate ~ \( 1/t \) when (Elmegreen 1994c)

\[
t \sim 1.5(G\rho_0 M^2)^{-1/2}.
\]

The extra power of \( M \) for the ring compared to the shell comes from the different geometry.

Equations (20) and (21) give the collapse conditions that should signal the onset of star formation. They include explicitly the self-gravity, internal pressure, and expansion of the region, while the kinematic instabilities that are also present are absorbed into the internal velocity dispersion through the term \( M \). These small scale instabilities increase \( c \) and decrease \( M \), which lengthens the timescale for collapse, as discussed in the previous section. Rings are more important than shells for large scale expansion in galaxies because most of the accumulated material is in the midplane. Indeed, Ehlerová et al. (1997) show with a numerical simulation how only the equator of a shell expanding into a thin galactic disk becomes unstable.

Another condition for star formation is that the swept-up layer has to be optically thick to starlight so it can cool and become dense. This condition was written by Franco & Cox (1986) as \( N > 10^{21} (Z_\odot/Z) \) cm\(^{-3}\) for total shell column density \( N \) and metallicity \( Z \). This second condition is implicitly included in the first through the term \( M \). If the swept-up region cannot cool, then \( M \) is small and the collapse takes a long time. Once cooling begins, \( M \) increases and the
collapse time drops. Thus cooling can trigger the instability, as well as the continued accumulation of mass. If $\omega \sim 1/t$ at the onset of cloud and star formation, as assumed above, and we write the variables in physically realistic units with $n_0$ the ambient density ahead of the front and $c$ the velocity dispersion inside the swept-up layer, then for a shell:

$$t_{SF} \sim 103 \left( \frac{n_0 M}{\text{cm}^{-3}} \right)^{-1/2} \text{ My} \quad (22)$$

$$R_{SF} \sim 176 M^{1/2} \left( \frac{c}{\text{km s}^{-1}} \right) \left( \frac{n_0}{\text{cm}^{-3}} \right)^{-1/2} \text{ pc.} \quad (23)$$

and for a ring:

$$t_{SF} \sim 124 \left( \frac{n_0 M^2}{\text{cm}^{-3}} \right)^{-1/2} \text{ My} \quad (24)$$

$$R_{SF} \sim 211 \left( \frac{c}{\text{km s}^{-1}} \right) \left( \frac{n_0}{\text{cm}^{-3}} \right)^{-1/2} \text{ pc.} \quad (25)$$

These results indicate that a shell or ring has to be relatively large before star formation appears along the periphery if the ambient density is about $1 \text{ cm}^{-3}$, the average for a galaxy disk at the Solar radius. Inside molecular clouds or in starburst regions, the density can be much larger, perhaps $10^3 \text{ cm}^{-3}$ or more. Then the triggering time and shell radius can be small.

Once a ring or shell stops expanding, the collapse proceeds quickly. Recall that for a moving ring, from equation (22),

$$t_{SF} \sim 1.5/(G\rho)^{1/2}$$

with $\rho = \rho_0 M^2$, but, using the same analysis for a static ring,

$$t_{SF} \sim e^{0.5(1+c^2/G\mu_0)} \left( \frac{4\pi G\rho}{(4\pi G\rho)^{1/2}} \right) \sim 0.8 \left( \frac{\rho}{G\rho} \right)^{1/2}. \quad (26)$$

$\mu_0$ is the mass/length of the ring and $\rho$ is the density inside. If the expansion history is unknown, then the SF time can be assumed to be between these two limits, which corresponds to

$$t_{SF} \sim 66 - 124 \ n_{\text{inside}}^{-1/2} \text{ My.} \quad (27)$$

### 3.4. Star Formation Triggered by Cloud Collisions

The dense shocked gas between two colliding clouds is another region where gravitational instabilities can lead to triggered star formation. The direct observation of this effect has been difficult, however.

The problem is that very few clouds are likely to be undergoing a collision at any one time. Considering a cloud cross section $\pi R^2$, speed $v$, and density $n$, the collision time is $\sim 1/(\pi R^2 vn)$ and the duration of each collision is $R/v$. This implies that the fraction of the time spent in the collision, or the fraction of all clouds in an ensemble that are currently colliding, is $\pi R^3 n$, which is about the volume filling factor of the clouds. This filling factor is usually very low for both
clouds and their clumps, i.e., less than 10% for diffuse clouds and less than 1% for molecular clouds, so very few clouds are likely to be involved in a collision at any one time in a Solar neighborhood environment. The fraction can be much higher in spiral density wave shocks where the clouds converge (Kwan & Valdes 1983; Roberts & Steward 1987; Kenney & Lord 1991). However, in other regions, cloud collisions are so rare that very few examples have been found. Some of the proposed regions are in Table 4.

The collapse of shocked layers between converging flows has been studied both analytically and numerically. Many of the results are the same as in the case of a shocked layer or shell so they are not repeated here. This section summarizes the recent calculations that deal specifically with gravitational collapse during cloud collisions.

Non-magnetic clouds typically destroy each other during an off-axis collision at relative velocities exceeding several times the internal sound speed (Kahn 1955; Chieze & Lazareff 1980; Hausman 1981; Gilden 1984) or several times the escape speed (Vázquez & Scalo 1989), whichever is greater. This destruction occurs for two reasons: (1) the shock between the clouds is confined only in the direction perpendicular to the contact plane, so high-pressure shocked material can squirt out parallel to this plane, and (2) the non-overlapping portions of the clouds do not feel the pressure from the shock and do not slow down to a common speed.

Table 4. Triggered Star Formation in Cloud Collisions

| Region            | Reference               |
|-------------------|-------------------------|
| NGC 1333          | Loren 1976              |
| LkHα              | Loren 1977              |
| MBM 55 DR 15, 20  | Vallee & Avery 1990     |
| G110-13           | Odenwald et al. 1990    |
| HVC+disk collisions | Lepine & Duvert 1994   |
| Sgr B2            | Hasegawa et al. 1994    |
| W49N              | Serabyn, Güsten, & Schulz 1993 |
| W75/DR 21         | Dickel et al. 1978      |
| Orion             | Greaves & White 1991; Womack, Ziurys, & Sage 1993 |
| IRAS 19550+3248   | Koo et al. 1994         |
| IRAS 2306+1451    | Vallee 1995             |

Early calculations suggested that because of this destruction, star formation requires head-on collisions (Stone 1970), or collisions between identical clouds (Gilden 1984), or collisions between clouds that are almost unstable initially (Lattanzio et al. 1985; Nagasawa & Miyama 1987). In such head-on collisions, or in the shocked layer between two unconfined gas streams (Hunter et al. 1986; Whitworth et al. 1994a), the compressed gas collapses on the time scale \( (G\rho)^{-1/2} \) for compressed density \( \rho \). This can be very short if there is cooling and the compressed density is large (Pongracic et al. 1992).

Now it is believed that even oblique collisions can be effective in triggering gravitational instabilities and star formation in the compressed layer (Usami, Hanawa, & Fujimoto 1995). Velocity shear at the interface reduces the growth
rate of the gravitational instability, as discussed above, but it may not eliminate the collapse altogether.

Recent models suggest that star formation can also be triggered by more general collisions if there is small scale structure. Habe & Ohta (1992) found that for collisions between two different clouds in hydrostatic equilibrium, the large cloud is disrupted by the processes discussed above, but the small cloud is compressed to trigger star formation. This is because the small cloud behaves like a cometary cloud, completely engulfed and confined by a large-scale gas flow. An even more realistic case was considered by Kimura & Tosa (1996), who found that collisions between clumpy clouds trigger collapse in the colliding clumps.

Stars may also form during the collisions between clumps in a turbulent cloud. This process is spontaneous in terms of the cloud evolution, that is, it does not need an external trigger for star formation, but it is triggered in terms of the stability of each clump, i.e., each clump needs the collision to form a star. A recent model using SPH of the gravitational instability of a compressed interface between two colliding clumps is in Pongracic et al (1992). They find that the instability operates much faster than the gravitational timescale in the uncompressed clouds. It is probably more like \((G\rho)^{1/2}\) for compressed density \(\rho\) (Whitworth et al. 1994a). This is consistent with the discussion in section 3.2 of the collapse time in a non-decelerating layer.

Even without pre-existing clumps, supersonic turbulence inside clouds will make dense sheet-like structures in the converging flows, and these structures can be unstable to form stars by gravitational collapse, as in the clump collision scenario. Early versions of this model were in Sasao (1973), with more recent work in Sabano, & Tosa (1985), Elmegreen (1993), Padoan (1995), Vazquez-Semadeni, Passot, & Pouquet (1996), and Padoan, Nordlund & Jones (1997). The detailed physical processes are similar to those discussed by Hunter et al. (1986) and others. A new instability for oblique and shearing shock fronts at the interface between colliding streams combines self-gravity with the Kelvin-Helmholtz instability; it may give small scale structure and even gravitational collapse into stars (Hunter, Whitaker & Lovelace 1997).

Models of clump or cloud collisions with magnetic fields are in their infancy. Some first studies are in Byleveld, Melrose, & Pongracic (1994) and Byleveld, & Pongracic (1996).

All of these models, whether they involve external shocks, converging flows, or clump collisions, make flattened gas sheets at some point during the star formation process. Larson (1985) has emphasized this point, and Hartmann et al. (1994, 1996) have considered the collapse of such layers and the spectra of the resulting stars.

4. Turbulence Effects

Turbulence changes our ideas about star formation and triggering in several ways. For example, it makes clouds clumpy even without self-gravity, presumably making fractal structure in the gas. This gives the clouds a hierarchical nature, and may ultimately cause stars to form in clusters, associations and complexes.
Pre-compression density and kinematic structures resulting from turbulence can also distort a shock front, causing bright rims to form and the acceleration to be intermittent. Positive acceleration makes comet shapes, while negative acceleration makes bullets.

In a turbulent medium, the ambient pressure affects the degree of gravitational self-binding because the velocity dispersion of a virialized cloud depends on the pressure:

$$\Delta v_{\text{virialized cloud}} \sim G^{3/8} P^{1/8} M^{1/4}.$$  

When the velocity dispersion is very high, significantly greater than $\sim 10 \text{ km s}^{-1}$, OB stars cannot easily destroy the clouds in which they form, and then massive clusters can form with high efficiency, allowing themselves to remain bound when the little remaining gas eventually leaves (Elmegreen & Efremov 1997). Even for lower velocity dispersions, greater cloud binding at higher pressure should lead to a greater probability of forming a bound cluster. This implies that triggered star formation in high pressure regions may preferentially form bound clusters. For example, cloud collisions have been proposed to provoke the formation of globular clusters in the Large Magellanic Clouds (Fujimoto & Kumai 1997). Globular clusters do indeed form in high pressure environments, such as early galaxy halos, interacting galaxies and starbursts, and occasional dwarf galaxy GMC cores, particularly in or near high pressure HII regions (Elmegreen & Efremov 1997).

Another effect of turbulence is the time-size correlation, $t \propto R^{1/2}$, which implies that larger regions form stars for a longer time (Elmegreen & Efremov 1996). This observation may be related to the observed expansion of OB associations in the sense that every large region of star formation, which in this case may be a whole OB association, contains smaller regions as part of the natural hierarchical structure, and the large regions have larger velocity dispersions at birth than the small regions. Thus there will be a systematic progression in size, age, and velocity dispersion for all regions of star formation, including the progression from OB subgroups to whole associations. This will make the large scale association appear to be an expanded version of the smaller scale subgroup, when in fact all may be parts of a continuous hierarchy of structure that extends all the way from small, multiple stellar systems to giant star complexes.

The problem of star formation is considerably harder if much of the small scale structure observed in clouds is from turbulence, as opposed to, say, gravitational instabilities. Turbulent structures need not be self-gravitating, and indeed many of the small clumps in giant molecular clouds appear to be non self-gravitating. Then one wonders what transition takes place before they form stars. Perhaps collisions between turbulent clumps are required to trigger star formation, as in the models discussed above. Vazquez-Semadeni, Passot, & Pue- quet (1996) discuss constraints on the polytropic index for turbulence to induce the collapse of clumps.

Figure 5 (right) shows the ratio of virial mass to luminous mass for clumps in several surveys, as indicated by the symbols. The virial mass is taken to be $5Re^2 / G$ for FWHM radius $R$, Gaussian velocity dispersion $c$, and gravitational constant $G$. Evidently, the smaller clumps are less self-gravitating than the larger clumps, with a systematic progression toward higher $5Re^2/(GM)$ with lower $R$ (Bertoldi & McKee 1992; Falgarone, Puget, & Pérault 1992;
Vazquez-Semadeni, Ballesteros-Paredes, & Rodriguez 1997). Considering that the luminous mass is \( M \propto c T R^2 \) for temperature \( T \), this ratio should be \( \propto c/(RT) \propto 1/(R^{0.5}T) \) for the a velocity-size trend \( v \propto R^{0.5} \). The solid line shows the predicted trend for constant \( T \).

5. Conclusions

Star formation can be triggered by dynamical processes when high pressures surround pre-existing clouds, when high pressures accumulate dense cloudy material into layers that collapse gravitationally into dense cores, when high pressures accumulate the ambient material into shells or rings, which then collapse into cores, and when clouds, clumps, or turbulent streams collide. Examples of triggered star formation in these cases were given, as were the thresholds for the onset of star formation. The evidence suggests that dynamical triggering is widespread and operates over a large range of scales.

The primary reason for the ubiquity of triggered star formation is that the gravitational timescale in a molecular cloud, \( (G\rho)^{-1/2} \), is shorter than the lifetime of an O-type star, so dense molecular gas is commonly induced to collapse in the midst of high pressures from HII regions, stellar winds and supernova explosions. Also, the gravitational timescale in the ambient disks of galaxies is shorter than or comparable to the lifetime of an OB association, so whole clusters can make giant shells and new molecular clouds. In different regions, the balance between these basic time scales may change, and then triggered star formation can have more or less importance compared to spontaneous processes.
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