The role of cosmic rays on magnetic field diffusion and the formation of protostellar discs

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in collaboration with

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• Development of new telescopes with higher and higher resolution
• New observation techniques with consequent large data sets

• \( \text{H}_3^+ \): UKIRT, VLT-CRIRES — Indriolo+ (2012)
• \( \text{OH}^+, \text{H}_2\text{O}^+ \): Herschel — Neufeld+ (2010), Gerin+ (2010)
• \( \gamma \)-ray emission: Fermi-LAT (CTA) — Montmerle (2010)
• magnetic field morphology: SMA, Planck (ALMA) — Girart+ (2009)

All these observations require a solid theoretical support. A detailed effort in the modelling of the CR spectrum, and more precisely of its low-energy tail, was missing as well as the integration of the models in chemical and numerical codes for interpreting observations.
• Diffuse clouds ($A_v \sim 1$ mag) → the UV radiation field is the principal ionising agent (photodissociation regions);

• Dense clouds ($A_v \gtrsim 5$ mag) → the ionisation is due to low-energy CRs ($E < 100$ MeV) and, if close to young stars, to soft X–rays ($E < 10$ keV).

\[
\begin{align*}
\text{CR} & \quad \text{He} & \quad \text{He}^+ & \quad H^+ & \quad H_2 & \quad H_2^+ & \quad H_3^+ \\
& & & & & & \quad \text{OH}^+ & \quad \text{HCO}^+ & \quad \text{N}_2\text{H}^+
\end{align*}
\]

polyatomic ions

\[
\begin{align*}
\text{CH, NH, OH} & \quad \text{H}_2\text{O, NH}_3, \text{H}_2\text{CO} \\
\text{O}_2, \text{N}_2 & \quad \text{NO, CN}
\end{align*}
\]

electronic recombination

reactions among neutral molecules
Dense cores
(HCO\(^+\), DCO\(^+\))

Caselli+ (1998)

Diffuse clouds
(OH, HD, NH)

Black & Dalgarno (1977),
Hartquist+ (1978), Black+ (1978),
vvan Dishoeck & Black (1986),
Federman+ (1996)

(H\(_3^+\))

McCall+ (1993), Geballe+ (1999)

(OH\(^+\), H\(_2\)O\(^+\))

Neufeld+ (2010), Gerin+ (2010)

\[ \zeta_{H_2} \approx 10^{-18} - 10^{-17} \text{ s}^{-1} \]

\[ \zeta_{H_2} \approx 10^{-16} - 10^{-15} \text{ s}^{-1} \]
**Main questions:**

- origin of the CR flux that generates such a high ionisation rate ($\zeta_{CR}$) in diffuse regions;
- how to reconcile these values with those ones measured in denser regions;

**Different strategies approaching these problems:**

- effects of Alfvén waves on CR streaming
  - Skilling & Strong (1976); Hartquist+ (1978); Padoan & Scalo (2005);
  - Everett & Zweibel (2011); Rimmer+ (2012); Morlino & Gabici (2014);

- magnetic mirroring and focusing
  - Cesarsky & Völk (1978); Chandran (2000); **PM & Galli (2011,2013)**;

- possible low-energy CR flux able to ionise diffuse but not dense clouds
  - Takayanagi (1973); Umebayashi & Nakano (1981); McCall+ (2003); **PM, Galli & Glassgold (2009)**
The story so far...

Theoretical model (PM, Galli & Glassgold 2009)

computing the variation of the ionisation rate due to cosmic rays, $\zeta_{\text{CR}} \, [\text{s}^{-1}]$, inside a molecular cloud, with the increasing of the column density, $N \, [\text{cm}^{-2}]$, of the traversed interstellar matter.

$$\zeta_{\text{H}_2} (N) = \eta_h \zeta_{\text{p}} (N) + \zeta_{\text{e}} (N)$$

- correction due to the presence of heavy nuclei among CRs
- correction due to the ionisation of secondary electrons

$$4\pi \int_0^\infty j_p(E, N) \eta_p^\text{sec}(E) \sigma_{p+\text{H}_2}(E) dE$$

$$4\pi \int_0^\infty j_e(E, N) \eta_e^\text{sec}(E) \sigma_{e+\text{H}_2}(E) dE$$
**CR–proton and electron energy loss function**

![Diagram of CR–proton and electron energy loss functions](image)

**BUT... ENERGY LOSSES DO ONLY HALF OF THE JOB!**

- **PM & GALLI (2011)**
- **PM & GALLI (2013)**
- **PM, HENNEBELLE & GALLI (2013)**
magnetic mirroring
  bounces many CRs
  out of the core

magnetic focusing
  increases CR flux
  in the core

non-uniformity of the magnetic field

The Larmor radii of ionising CRs are smaller than typical sizes of Bok globules (~ 0.05 pc), dense cores (~ 1-5 pc), and GMC (~ 25 pc).

Césarsky & Völk, 1978
Theoretical challenge: formation of protostellar discs.
- Magnetic fields entrained by collapsing cloud brake any rotational motion and prevents cloud’s collapse (Galli et al. 2006; Mellon & Li 2008; Hennebelle & Fromang 2008).

Observational evidence of the presence of discs

Class II YSO

Class I YSO

Class 0 YSO
A number of possible solutions to solve the magnetic braking

(i) non-ideal MHD effects (Shu+ 2006; Dapp & Basu 2010; Krasnopolsky+ 2011; Braiding & Wardle 2012)

(ii) misalignment between $B$ and $J$ (Hennebelle & Ciardi 2009; Joos+ 2012)

(iii) turbulent diffusion of $B$ (Seifried+ 2012; Santos-Lima+ 2013; Joos+ 2013)

(iv) depletion of the infalling envelope anchoring $B$ (Mellon & Li 2009; Machida+ 2011)
Numerical models: rotating collapsing core

Table 1. Parameters of the simulations described in the text (from Joos et al. 2012): mass-to-flux ratio, initial angle between the magnetic field direction and the rotation axis, time after the formation of the first Larson’s core (core formed in the centre of the pseudo-disc with \( n \gtrsim 10^{10} \, \text{cm}^{-3} \) and \( r \sim 10 – 20 \, \text{AU} \)), maximum mass of the protostellar core and of the disc. Last column gives information about the disc formation.

<table>
<thead>
<tr>
<th>Case</th>
<th>( \lambda )</th>
<th>( \alpha_{B,J} )</th>
<th>( t )</th>
<th>( M_\star )</th>
<th>( M_{\text{disc}} )</th>
<th>Disc ?</th>
</tr>
</thead>
<tbody>
<tr>
<td>A_1</td>
<td>5</td>
<td>0</td>
<td>0.824</td>
<td>–</td>
<td>–</td>
<td>N</td>
</tr>
<tr>
<td>A_2</td>
<td>5</td>
<td>0</td>
<td>11.025</td>
<td>0.26</td>
<td>0.05</td>
<td>N</td>
</tr>
<tr>
<td>B</td>
<td>5</td>
<td>( \pi/4 )</td>
<td>7.949</td>
<td>0.23</td>
<td>0.15</td>
<td>Y</td>
</tr>
<tr>
<td>C</td>
<td>5</td>
<td>( \pi/2 )</td>
<td>10.756</td>
<td>0.46</td>
<td>0.28</td>
<td>K</td>
</tr>
<tr>
<td>D</td>
<td>2</td>
<td>0</td>
<td>5.702</td>
<td>0.24</td>
<td>–</td>
<td>N</td>
</tr>
<tr>
<td>E</td>
<td>17</td>
<td>0</td>
<td>6.620</td>
<td>0.43</td>
<td>0.15</td>
<td>K</td>
</tr>
</tbody>
</table>

\(^a\) A disc with flat rotation curve is formed (Fig. 15 in Joos et al. 2012).
\(^b\) No significant disc is formed \( (M_{\text{disc}} < 5 \times 10^{-2} \, M_\odot) \).
\(^c\) A keplerian disc is formed (Fig. 14 in Joos et al. 2012).
Numerical models: rotating collapsing core

Intermediate magnetisation $\lambda=5$

Aligned rotator $(J, B)=0$

It is not possible to unravel magnetic from column-density effects, but both intervene on the decrease of $\zeta_{\text{CR}}$. Deviations between iso-density contours and $\zeta_{\text{CR}}$ maps can be interpreted as due to magnetic imprints.

Field lines in the inner 600 AU

Padovani+ 2013
Numerical models: rotating collapsing core

Intermediate magnetisation $\lambda = 5$

Perpendicular rotator $(J,B) = \pi/2$

$\zeta_{CR} < 10^{-18} \text{ s}^{-1}$ down to $2 \times 10^{-21} \text{ s}^{-1}$ in the inner area with an extent of a few tenths of AU. We can assume that the gas is effectively decoupled with the magnetic field.

Field lines in the inner 600 AU

Padovani + 2013
Numerical models: rotating collapsing core

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Field lines in the inner 600 AU

Without magnetic effects

Padovani+ 2013
Numerical models: rotating collapsing core

Weak magnetisation $\lambda = 17$

Aligned rotator $(J, B) = 0$

The magnetic braking is very faint and the rotation acts in wrapping powerfully the field lines. The region with $\zeta_{CR} < 10^{-18} \text{s}^{-1}$ broadens out along the rotation axis where field line tangling up is very marked.

Field lines in the inner 600 AU

Padovani+ 2013
Numerical models: rotating collapsing core

Weak magnetisation \( \lambda = 17 \)

Aligned rotator \((J, B) = 0\)

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Field lines in the inner 600 AU

Padovani+ 2013
A useful fitting formula

When in presence of a magnetic field, the effective column density, $N_{\text{eff}}$, seen by a CR can be much larger than that obtained through a rectilinear propagation.

$$N(H_2) = \text{average column density seen by an isotropic flux of CRs}$$

$$N_{\text{eff}} = (1 + 2\pi F^s) N(H_2)$$

$$F = F(|B|, |B_\phi/B_p|)$$

$$s = s(n)$$

$$\zeta_{\text{eff}}^{H_2} \propto \zeta_{H_2}(N_{\text{eff}})$$
Cosmic rays and magnetic diffusion

Drifts of charged species with respect to neutrals determines different regimes of magnetic diffusivity.

\[ \frac{\partial \vec{B}}{\partial t} + \nabla \times (\vec{B} \times \vec{U}) = \nabla \times \left\{ \eta_O \nabla \times \vec{B} + \eta_H (\nabla \times \vec{B}) \times \frac{\vec{B}}{|\vec{B}|} + \eta_{AD} \left( (\nabla \times \vec{B}) \times \frac{\vec{B}}{|\vec{B}|} \right) \times \frac{\vec{B}}{|\vec{B}|} \right\} \]

Resistivities

\[ \eta_{AD} = \frac{c^2}{4\pi} \left( \frac{\sigma_P}{\sigma_P^2 + \sigma_H^2} - \frac{1}{\sigma_\parallel} \right) \]
\[ \eta_H = \frac{c^2}{4\pi} \left( \frac{\sigma_H}{\sigma_P^2 + \sigma_H^2} \right) \]
\[ \eta_O = \frac{c^2}{4\pi \sigma_\parallel} \]

Conductivities

\[ \sigma_\parallel = \frac{ecn(H_2)}{B} \sum_i Z_i x_i \beta_{i,H_2} \]
\[ \sigma_P = \frac{ecn(H_2)}{B} \sum_i \frac{Z_i x_i \beta_{i,H_2}}{1 + \beta_{i,H_2}^2} \]
\[ \sigma_H = \frac{ecn(H_2)}{B} \sum_i \frac{Z_i x_i}{1 + \beta_{i,H_2}^2} \]

The degree of diffusion is determined by the ionisation fractions that, in turn, are determined by \( \zeta^{H_2} \).
Abundance of charged species

We adopted a simplified chemical network that computes the steady-state abundance of $\text{H}^+$, $\text{H}_3^+$, a typical molecular ion $m\text{H}^+$ (e.g. $\text{HCO}^+$), a typical metal ion $M^+$ (e.g. $\text{Mg}^+$), $e^-$ and dust grains ($g^0$, $g^-$) as a function of:

— H$_2$ density ;
— Temperature ;
— Cosmic-ray ionisation rate.

\[
\begin{aligned}
\text{computed at each spatial position in our models.}
\end{aligned}
\]

Table 1. Parameters of the simulations described in the text: non-dimensional mass-to-flux ratio $\lambda$, initial angle between the magnetic field direction and the rotation axis $\alpha_{\text{B},\lambda}$, time after the formation of the first Larson’s core $t$ (core formed in the centre of the pseudo-disc with $n \geq 10^{10}$ cm$^{-3}$ and $r \sim 10 - 20$ AU), initial mass $M_{\text{in}}$, mass of the protostellar core $M_\star$ and of the disc $M_{\text{disc}}$.

<table>
<thead>
<tr>
<th>Case</th>
<th>$\lambda$</th>
<th>$\alpha_{\text{B},\lambda}$</th>
<th>$t$</th>
<th>$M_{\text{in}}$</th>
<th>$M_\star$</th>
<th>$M_{\text{disc}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$L_0$</td>
<td>5</td>
<td>0</td>
<td>0.824</td>
<td>1</td>
<td>--</td>
<td>--</td>
</tr>
<tr>
<td>$L_2$</td>
<td>5</td>
<td>$\pi/2$</td>
<td>10.756</td>
<td>1</td>
<td>0.46</td>
<td>0.28</td>
</tr>
<tr>
<td>$H$</td>
<td>$\sim 2$</td>
<td>no initial rotation</td>
<td>6.000</td>
<td>100</td>
<td>1.24$^a$</td>
<td>0.87$^a$</td>
</tr>
</tbody>
</table>

$^a$ This value refers to the densest fragment formed.

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Effects of the grain size distribution

Grains play a decisive role in determining the degree of coupling between gas and B.

**Three grain size distribution**

\[
a_{\text{min}} = 10^{-5} \text{ cm} \quad \text{(representative of large grains formed by compression and coagulation during collapse; Flower et al. 2005)}
\]

\[
a_{\text{min}} = 10^{-6} \text{ cm} \quad \text{(the minimum grain radius of a MRN size distribution Mathis et al. 1977 that gives the same grain opacity found by Flower et al. 2005)}
\]

\[
a_{\text{min}} = 10^{-7} \text{ cm} \quad \text{(typical size of very small grains)}
\]

\[
a_{\text{max}} = 3 \times 10^{-5} \text{ cm} \quad \text{(Nakano et al. 2002)}
\]

larger grains \(\rightarrow\) less number of grains
\(\rightarrow\) more free electrons (one electron per grain)
Dependence of chemical abundances on $\zeta^{\text{H}_2}$

Density of charged species usually parameterised as $n(i) \propto (\zeta^{\text{H}_2})^{k'}$ with $k' \approx 1/2$ (e.g. Ciolek & Mouschovias 1994, 1995), but... it depends on the grain size!

$n(\text{H}_2) = 10^6 \text{ cm}^{-3}$

$a_{\text{min}} = 10^{-5} \text{ cm} \text{ (solid lines)}$

$a_{\text{min}} = 10^{-7} \text{ cm} \text{ (dotted lines)}$

<table>
<thead>
<tr>
<th>$k'$</th>
<th>$10^{-5}$</th>
<th>$10^{-7}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$e^-$</td>
<td>1/2</td>
<td>1</td>
</tr>
<tr>
<td>$M^+$</td>
<td>1/2</td>
<td>0</td>
</tr>
<tr>
<td>$g^-$</td>
<td>0</td>
<td>1/2</td>
</tr>
<tr>
<td>$M_3^+, H^+, mH^+$</td>
<td>1/2</td>
<td>1/2</td>
</tr>
</tbody>
</table>
**ζ_H2** and the gas-magnetic field decoupling

Drift velocity of magnetic field **U_B** : it can be represented by the velocity of the charged species (frozen with the field lines) with respect to neutrals.

By comparing **U_B** with the fluid velocity, it is possible:

- to assess the degree of diffusion of the field;
- to estimate the size of the region where gas and B are decoupled.

Following Nakano et al. (2002), **U_B** can be written as a function of the resistivities.

\[
\vec{U}_B = \vec{U}_{AD} + \vec{U}_H + \vec{U}_O = \frac{4\pi}{cB^2} \left[ (\eta_{AD} + \eta_O) \vec{j} \times \vec{B} + \eta_H \left( \vec{j} \times \vec{B} \right) \times \frac{\vec{B}}{B} \right]
\]

\[
\frac{1}{t_B} = \frac{1}{t_{AD}} + \frac{1}{t_H} + \frac{1}{t_O}
\]

\[
t_k = \frac{R}{U_k}, \quad (k = AD, H, O)
\]

*R is the typical length scale of the region (assumed as the distance from the density peak)*
\[ \frac{1}{t_B} = \frac{1}{t_{AD}} + \frac{1}{t_H} + \frac{1}{t_O} \]

\[ t_k = \frac{R}{U_k}, \quad (k = AD, H, O) \]

\[ R \text{ is the typical length scale of the region} \]
\[ \text{assumed as the distance from the density peak} \]

Nakano et al. (2002): comparison between \( t_B \) and the free-fall time scale.
Here we compare \( t_B \) with the \( t_{\text{dyn}} = \frac{R}{U} \) (\( U \) = fluid velocity including infall and rotation).

In regions where:
- \( t_B < t_{\text{dyn}} \) : B is partially decoupled and it has less influence on gas dynamics;
- \( t_B > t_{\text{dyn}} \) : diffusion is not efficient and B remains well coupled to the gas.

Is there a relevant variation in \( t_B \) using a constant \( \zeta_{H_2} \approx 10^{-17} \text{ s}^{-1} \) or accounting correctly for the dependence of \( \zeta_{H_2} \) on \( N(H_2) \) and B?

YES!
Cosmic-ray ionisation rate and diffusion time scale
LOW-MASS CASE

The region of decoupling decreases with decreasing $a_{\text{min}}$ (from $\approx 50$ AU to $\approx 20$ AU of radius), but it is still present.

The correct evaluation of the CR ionisation rate as a function of density and magnetic field allows the diffusion time to decrease up to three order of magnitudes.
Cosmic-ray ionisation rate and diffusion time scale

**HIGH-MASS CASE**

- $t_B > t_{dyn}$
- $t_B < t_{dyn}$

For $a_{min} = 10^{-5}$ cm the decoupling region has a radius of $\approx 100$ AU, but for $a_{min} = 10^{-7}$ cm it vanishes.

Different behaviour wrt low-mass case: the field diffuses faster in the HM case (turbulent nature of the flow).


The turbulent diffusion present in the HM model reduces the field strength making the microscopic resistivities smaller (no decoupling).
Conclusions

- The study of low energy ($E < 1$ GeV) cosmic rays is fundamental for correctly dealing with chemical modelling and non-ideal MHD simulations;

- In order to study the cosmic-ray propagation we accounted for energy losses and magnetic field effects: an increment of the toroidal component, and in general a more tangled magnetic field, corresponds to a decrease of $\zeta_{H_2}$ because of the growing preponderance of the mirroring effect;

- The extent to which density and magnetic effects make $\zeta_{H_2}$ decrease can be ascribed to the degree of magnetisation; $\zeta_{H_2} < 10^{-18}$ s$^{-1}$ is attained in the central 300-400 AU, where $n > 10^9$ cm$^{-3}$, for toroidal fields larger than about 40% of the total field in the cases of intermediate and low magnetisation ($\lambda = 5$ and 17, respectively);

- We found an increase in $\eta$ in the innermost region of a cloud after the collapse onset: (1) field has to be considerably twisted; (2) dust grains had time to grow by coagulation.

A correct treatment of CR propagation can explain the occurrence of a decoupling region between gas and magnetic field that in turn affects the disc formation.